Accretion Flow and Precession Phenomena in Cataclysmic Variables

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Accretion Flow and Precession Phenomena in Cataclysmic Variables

Daniel James Rolfe M.Phys.

Submitted for the degree of Doctor of Philosophy

June 2001
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ABSTRACT

Accretion Flow and Precession Phenomena in Cataclysmic Variables

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Cataclysmic variables are interacting binaries where the accretion flow is usually the dominant source of emission. They provide a valuable opportunity to study accretion, the most potent source of energy in the universe. This thesis presents optical studies of two such systems, V348 Pup and IY UMa, providing new insights into a variety of accretion phenomena.

V348 Pup has a persistently hot accretion disc, while in IY UMa the disc undergoes a series of outbursts in which it changes from a cool state to a much hotter, brighter state. We view both systems from close to the plane of the disc, making it possible to employ indirect imaging techniques to map both the spatial and velocity distributions of accretion flow.

V348 Pup harbours a precessing asymmetric disc, as does IY UMa during superoutburst, leading to a superhump modulation in luminosity as the companion tidally interacts with the precessing disc. The disc shape in IY UMa as it shows 'late' superhumps at the end of a superoutburst is found, and a simple model reveals varying energy dissipation at the stream-disc impact to be the source of the late superhumps. This study finds that this might also be the mechanism producing the superhumps in V348 Pup.

Spectroscopy reveals that V348 Pup is an SW Sex type star, where anomalous line emission is dominated by a source other than the accretion disc. This study supports the model in which some of the accretion stream overflows the disc and is expelled from the system by a magnetic propeller anchored in the disc.

Spectroscopy of IY UMa during outburst reveals rapidly changing accretion flow, with reprocessing of radiation in the outer disc and on the companion star, and possibly spiral shock waves in the disc. Quiescent observations reveal absorption features from the donor star, along with emission from the stream-disc impact and accretion disc.
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Thanks to Mum and Dad for their confidence in me, and for providing a great place for many much needed breaks in the country.
For Mum and Dad
Chapter 1

Introduction

1.1 Accretion

Accretion, where gravitational potential energy is released when material is captured by a compact object, is the most potent source of energy in the Universe. The other major source of energy is nuclear fusion in the centres of stars. Fusing a mass $m$ of hydrogen nuclei to form helium, the most common nuclear energy source in stars, releases an energy of

$$E_{\text{nuc}} = 0.007mc^2$$

(Frank, King & Raine 1992). In contrast, when a mass $m$ of material is captured by a compact object of mass $M_{\text{comp}}$ and radius $R_{\text{comp}}$ the gravitational energy released is

$$E_{\text{acc}} = \frac{GM_{\text{comp}}m}{R_{\text{comp}}}.$$ 

If the accretor is a typical neutron star (mass $1.4M_\odot$ and radius 10 km), $E_{\text{acc}} \approx 30E_{\text{nuc}}$, i.e. accretion onto a neutron star liberates 30 times more energy per unit mass than hydrogen burning. Accretion powers some of the most energetic astrophysical objects: active galactic nuclei and interacting binary stars.

Interacting binaries, where two stars orbit one another with matter from one star being accreted by the other, provide our best opportunity for studying accretion. Emission from the accretion flow in binaries provides a significant, often dominant, contribution to their luminosity. Accretion behaviour changes on accessible timescales, from fractions of seconds, through minutes, hours, days and months and longer. The orbital periods are short enough, less than two hours in the case of some cataclysmic variables, to view
systems from many orientations during one night of observations, and hence obtain detailed
information on the accretion flow. It is for these reasons that we are able to study in detail
accretion flow in interacting binaries, and in the work which follows a variety of techniques
for exploiting the information available from observations of binaries will be presented and
employed.

1.2 The Roche potential

When two masses, $M_1$ and $M_2$, execute circular orbits around their centre of mass, the
orbital period is given by Kepler’s third law:

$$P_{\text{orb}} = 2\pi \sqrt{\frac{a^3}{G(M_1 + M_2)}} \tag{1.1}$$

where $a$ is the separation of the masses. To understand the flow of matter in such a binary
system, it is informative to consider behaviour in a non-inertial reference frame corotating
with the binary about the centre of mass. The force, $\vec{F}$, experienced by a point mass
$m$ as measured in this frame will be the gravitational attraction from the two stars, the
centrifugal force, $\vec{F}_{\text{cent}} = -m\vec{\omega} \wedge (\vec{\omega} \wedge \vec{r})$, and the Coriolis force, $\vec{F}_{\text{cor}} = -2m\vec{\omega} \wedge \vec{v}$, where
$\vec{v}$ is the velocity of the point mass in the corotating frame. If the two stars are sufficiently
centrally condensed to be treated as point masses $M_1$ and $M_2$ at position vectors $\vec{r}_1$ and
$\vec{r}_2$ respectively (with the origin of the coordinate system at the centre of mass), then

$$\vec{F} = -\frac{GM_1m}{|\vec{r} - \vec{r}_1|^3}(\vec{r} - \vec{r}_1) - \frac{GM_2m}{|\vec{r} - \vec{r}_2|^3}(\vec{r} - \vec{r}_2) + \vec{F}_{\text{cent}} + \vec{F}_{\text{cor}}. \tag{1.2}$$

Using $\vec{F} = -m\nabla U$ we can express $\vec{F}$ in terms of an effective potential, $U$, in the corotating
frame. Note that this does not include the Coriolis force which depends on velocity and
so cannot be expressed in terms of a potential:

$$U = -\frac{GM_1}{|\vec{r} - \vec{r}_1|} - \frac{GM_2}{|\vec{r} - \vec{r}_2|} - \frac{4\pi^2\nu^2}{2P_{\text{orb}}^2}. \tag{1.3}$$

This is known as the Roche potential. The last term is the centrifugal term, rewritten
using Equation 1.1. Equation 1.3 can be reduced to

$$U \times \frac{a}{GM_1} = \frac{1}{d_1} - \frac{q}{d_2} - \frac{(1 + q)d^2}{2}, \tag{1.4}$$

where $q$ is the mass ratio of the system, $q = \frac{M_2}{M_1}$, and $d_1$, $d_2$ and $d$ are $|\vec{r} - \vec{r}_1|/a$, $|\vec{r} - \vec{r}_2|/a$
and $|\vec{r}|/a$ respectively. This shows that when distance is measured in units of the orbital
separation, the geometry of the Roche potential depends solely on the mass ratio, $q$. 
The Roche potential in the orbital plane\footnote{The orbital plane defines vertical coordinate $z = 0$, where $z$ increases out of the page, and the sense of rotation of the binary orbit in Figure 1.1 is anticlockwise. These axes, $(x, y, z)$, and direction of rotation are assumed throughout this work unless otherwise stated.} is plotted in Figure 1.1. There are five stationary points in the Roche potential, known as the Lagrangian points: three saddle points, $L_1$, $L_2$ and $L_3$, and two maxima, $L_4$ and $L_5$. The Roche lobes are the two teardrop-shaped volumes within which the potential is less than that at the $L_1$ point. Material within a Roche lobe is gravitationally bound, and moves most easily (i.e. requires the least energy to move) between the Roche lobes via the $L_1$ point.

When the two masses are stars, it is their sizes relative to their Roche lobes which determines if, or how, they interact. If both stars are within their Roche lobes they form...
1.3 Cataclysmic variables

A detached binary, with no exchange of matter possible between them except via a stellar wind. If both stars are larger than their Roche lobes, they form a double-cored system with a single dumbbell-shaped common envelope. These systems are known as contact binaries. If one star is within its Roche lobe, while the other star fills its Roche lobe, we have a semi-detached system, an interacting binary where material escapes the lobe-filling donor star through the L1 point and is accreted by the compact object (accretor). If the accretor is a normal star, then the system is an algol. If the accretor is a black hole or neutron star, the system is an X-ray binary, while systems with white dwarfs accreting from Roche lobe filling late type dwarfs are cataclysmic variables. Accretion in interacting binaries by this mechanism is commonly known as Roche lobe overflow.

1.3 Cataclysmic variables

As described above, cataclysmic variables (CVs) are interacting binaries where material is transferred via Roche lobe overflow from a late type dwarf star onto a white dwarf. They have orbital periods of typically just a few hours, with the shortest being just 78 minutes for CVs with hydrogen-rich donors. The accretion stream, which is the flow of material escaping the donor star through the L1 point, does not simply fall directly onto the white dwarf as the Roche potential suggests: the Roche potential does not include the Coriolis force. The angular momentum possessed by matter starting from the L1 point causes the stream to loop around the white dwarf, forming a ring as the stream intersects itself. This ring, which eventually develops into an accretion disc as described in the next section, will settle around the circularization radius, the radius at which material in a Keplerian orbit around the white dwarf has the same specific angular momentum as the L1 point. In CVs where the white dwarf has a strong magnetic field (typically ~ $10^7$–$10^8$ Gauss (Warner 1995)) the accretion stream will be channeled along magnetic field lines onto the white dwarf and an accretion disc is never formed. Such systems are called polars, or AM Her type CVs. Systems with a lower but still significant white dwarf magnetic fields are called intermediate polars or DQ Hers, and typically have an accretion disc which is truncated by magnetic pressure within a certain distance from the white dwarf. Note that white dwarf magnetic fields are expected in all systems, even those classed as non-magnetic, but in non-magnetic systems the field is too weak to disrupt the accretion disc.

The non-magnetic CVs are further classified into four classes. The classical novae and recurrent novae show large amplitude eruptions, thought to arise from thermonuclear
runaways as accreted material builds up on the surface of the white dwarf. Classical novae are systems where only one such eruption has been observed. Systems where more than one eruption has been seen are called recurrent novae. Dwarf novae show regular fainter outbursts in brightness caused by an increase in the mass accretion rate. Nova-likes are CVs which do not show eruptions or outbursts, and are thought to be like dwarf novae which spend all their time in the high accretion rate outburst state. This thesis presents studies of a dwarf nova (IY UMa) and a nova-like (V348 Pup).

An illustration of a disc-accreting CV is shown in Figure 1.2. The stream leaves the L1 point, falling down towards the disc. At the point where the stream hits the disc, there will be shocks and dissipation of kinetic energy as the stream merges with the disc flow. A bright spot, or hotspot, forms at this stream-disc impact. This region is likely to have vertical structure raising it above the disc, and can be much brighter than the disc (Warner (1995), see Chapter 5 for an example).

In non-magnetic systems, the region where the inner disc meets the white dwarf is called the boundary layer. In this region the velocity of the disc must decrease to match the rotation of the white dwarf. This is thought to occur over a radial range much less than the white dwarf radius (Pringle 1977).

1.4 Accretion discs

The formation of an accretion disc begins when material settles around the circularization radius. In order for accretion to occur, material must be able to move inwards. The velocity of a circular orbit at radius r from the white dwarf (mass \( M_{wd} \)) will be the Keplerian velocity,

\[ V_{Kep} = \sqrt{\frac{GM_{wd}}{r}}, \tag{1.5} \]

corresponding to a specific angular momentum \( l = \sqrt{GM_{wd}r} \), which increases with r. Therefore, for material to move inwards, it must lose angular momentum. The ring of material must have a finite radial thickness, and from Equation 1.5 it will rotate differentially. Since \( V_{Kep} \) decreases with r, the presence of a viscosity will lead to (i) angular momentum transport outwards by shearing forces, and (ii) loss of energy from the material as it radiates away heat produced by the viscous processes. Thus the disc will spread as the energy loss drives most mass deeper into the white dwarf potential well, with angular momentum being conserved by a smaller outward flow of mass. This mechanism is described in the review of accretion disc theory by Pringle (1981).
Figure 1.2: A disc accreting cataclysmic variable with mass ratio $q = 0.13$. The accretion streams leave the teardrop-shaped donor star via the L₁ point, impacting the accretion disc and forming a hot spot as it dissipates kinetic energy.
1.4 Accretion discs

It is usually assumed that most of the disc is close enough to the primary that it is unaffected by the gravitational pull of the donor. The outer disc will be affected by the tidal influence of the donor, with this interaction transferring the angular momentum transported out through the disc into the binary orbit. In some cases the interaction is sufficiently strong to distort or truncate the outer disc (Paczynski 1977). This is discussed further in Section 1.6.

The detailed structure of the accretion disc is governed by the form of the viscosity. Given a prescription for the viscosity, and suitable boundary conditions, we can solve for the structure of the disc in a steady state using the equations of mass and angular momentum conservation. In a non-magnetic system, the usual boundary conditions are that very close to the surface of the white dwarf $\frac{du}{dr} = 0$ and $u = u_{wd}$, where $u$ is the angular velocity of the disc and $u_{wd}$ is that of the white dwarf. Even without knowing the form for the viscosity, the radial temperature structure can be obtained, assuming all energy from viscous dissipation is immediately radiated away following a blackbody law.

The vertical structure of CV discs is governed by hydrostatic equilibrium, which for a thin disc with isothermal vertical structure gives

$$\frac{H(r)}{r} = \frac{c_s}{V_{Kep}(r)}, \quad (1.6)$$

where $H(r)$ is the disc height at radius $r$ and $c_s$ is the sound speed in the disc (Warner 1995). This shows that discs with highly supersonic flow (where $V_{Kep}(r) \gg c_s$) are thin, which is the case in CV discs.

Detailed solution of the disc structure equations requires knowledge of the form of the viscosity. This brings us to a major problem with accretion disc theory: we don’t yet understand the source of viscosity in accretion discs. The Reynolds number is the ratio of inertial forces to viscous forces in a fluid flow, and in the case of molecular viscosity in accretion discs is never much less than $10^{14}$ (Frank, King & Raine 1992). This shows that the molecular viscosity is far too low to explain the angular momentum transport in accretion discs. The large Reynolds number indicates that accretion discs are unstable to turbulent flow (Frank, King & Raine 1992). Turbulent gas flow could provide the viscosity mechanism in accretion discs, with our ignorance of turbulence being encapsulated in two parameters, the scale and turnover velocity of the turbulent eddies, $\lambda_{turb}$ and $v_{turb}$. An upper limit to $\lambda_{turb}$ is the disc height, $H$, while supersonic eddies would form shocks likely to thermalize their energy (Frank, King & Raine 1992), suggesting an upper limit for $v_{turb}$ of $c_s$. This leads to the $\alpha$-prescription for accretion disc viscosity (Shakura & Sunyaev...
1.5 The outburst cycle

\[ \nu = \alpha c_s H, \] with \( \alpha \) (which must be less than \( \sim 1 \)) parametrizing our ignorance of the viscosity mechanism. Observations of non-steady discs in CVs tell us that \( \alpha \sim 0.01-1 \) (Warner 1995; for a recent example see Baptista & Catalán 2001).

Much work has been put into modeling hydrodynamic instabilities in accretion discs, with results suggesting that that purely convective instabilities cannot transport sufficient angular momentum outwards to drive accretion discs. Magnetic stresses, which exist as magnetic fields in the disc are distorted by the differential flow, have also been suggested, with Balbus & Hawley (1991) discovering a dynamical instability present in accretion discs with (even) a very weak magnetic field. Tout & Pringle (1992) followed this work, demonstrating that a self-sustaining magnetic dynamo is possible, producing magnetic-viscosity with \( \alpha \) in the range 0.1–0.7.

Another possible method of angular momentum transport in accretion discs is through spiral shocks. Tidally induced shocks, which form a spiral structure in the accretion disc as the shock wave propagates faster azimuthally than radially, were found in numerical simulations by Sawada, Matsuda & Hachisu (1986), and have since been observed in many simulations, e.g. Armitage & Murray (1998), and detected in observations of several CVs, e.g. IP Peg (Steeghs, Harlaftis & Horne 1997). We present a possible detection of spirals in the accretion disc of IY UMa in Chapter 7. Sawada, Matsuda & Hachisu (1986) found that gas flowing through the shock loses angular momentum and thus could be accreted even without the presence of turbulent or magnetic viscosity. It is not yet clear whether spiral shocks can drive angular momentum transport in the inner disc (Steeghs 1999).

The importance of spiral waves in the transport of angular momentum in accretion discs is unclear. While the viscosity due to magnetic turbulence appears to be possible in most, if not all discs, spiral waves have so far only been detected in a few dwarf novae in outburst and one nova-like, where the disc radius is large enough for the tidal influence of the donor star to produce the spiral shocks. Simulations still have some way to go before they can demonstrate conclusively which process or combination of processes can fully account for angular momentum transport in discs.

1.5 The outburst cycle

The history of the various models for dwarf nova outbursts is discussed in Warner (1995). Here I outline the disc instability model, now generally accepted as the mechanism underlying outbursts.
1.5 The outburst cycle

Observations, such as changing eclipse profiles (see Section 1.7) through the outburst cycle (Warner 1974; Vogt 1983a; Patterson et al. 2000), reveal the accretion disc to be the source of the outburst emission. Once this was understood, it was soon realized that the basic scenario must consist of an unstable disc which can switch between two states, one with a low mass accretion rate corresponding to the faint non-outburst state, and the other a high accretion rate outburst state. In the quiescent state between outbursts the mass accretion rate through the disc is lower than the transfer rate from the donor, leading to a build up of mass in the disc, until at some point the disc instability causes the switch to the high-accretion rate outburst state, emptying some (though not necessarily all (see Section 1.6)) of the built-up mass onto the white dwarf. The source of the instability is the ionization of hydrogen at $\sim 10^4$K: the opacity of hydrogen as it is becoming ionized is significantly higher than that of neutral hydrogen. This leads to two stable thermal equilibrium solutions for the same region of a disc depending on its ionization state.

![Diagram of the thermal limit cycle](image)

Figure 1.3: The thermal limit cycle.

Meyer & Meyer-Hofmeister (1981) produced the first detailed model of this disc instability. Figure 1.3 shows schematically the thermal equilibrium solutions for a circular annulus in the accretion disc with surface density $\Sigma (= \int \rho(z) dz)$ and mass transfer rate $\dot{M}$. We see that for some values of $\Sigma$ there is more than one value of $\dot{M}$, with the branch below A corresponding to the annulus in a cool, neutral, low accretion rate state, while
the top branch above B represents the hot, ionized, high accretion state. If the rate of mass flow into the annulus is less than that at A, the annulus can find a steady equilibrium on this lower branch, while for an external mass transfer rate greater than that at B, the annulus will be in a steady equilibrium on the upper branch. However, an annulus for which the rate at which mass arrives at its outer radius is between the accretion rates at A and B can find no stable equilibrium. If such an annulus is initially on the lower branch, it cannot transport mass inwards as quickly as it receives it, and so $\Sigma$ increases moving it up the lower branch. At A, it cannot move up the S-curve since that would require $\Sigma$ to decrease. Instead the annulus heats up, becoming ionized and moving to the upper branch of the curve. Now mass leaves the annulus faster than it arrives, so $\Sigma$ decreases, the annulus moving down the S-curve until it reaches B. At B, the annulus once more leaves the equilibrium curve, and cools to the lower branch, where this limit-cycle begins again. As one annulus rises to the hot branch, it leads neighbouring regions to do the same, both by heating neighbouring annuli and by rapidly transporting mass to the inner annulus thus increasing its $\Sigma$ above the critical value. Thus heating waves propagate across the disc, bringing most of the disc into the hot outburst state. Similarly, once an annulus drops back into the low state, it will start a cooling wave bringing the disc back into quiescence. The viscosity in the two states is not necessarily the same, with models of outbursts requiring a higher viscosity during outburst than in quiescence (see e.g. Whitehurst 1988a).

In this model, nova-like systems are easily explained as systems where the mass transfer rate from the donor is high enough to keep the disc on the stable upper branch of the S-curve, making nova-likes effectively dwarf novae in permanent outburst.

Recent particle simulations of discs which include the disc instability (Truss, Murray & Wynn 2001) reveal complicated outburst behaviours in which the inner region of the disc can be continually cycling between the low and high states, and also occasions when mini-outbursts can occur affecting only the inner disc.

This model has been successful at modeling various properties of outbursts in CVs, but Smak (2000) points out a number of inconsistencies between the model and observations, e.g. the model predicts an increase in luminosity of the disc during quiescence as it builds up mass, but this is not observed. Smak suggests that including the effect of irradiation of the donor causing a change in the mass transfer onto the disc might help. Buat-Ménard, Hameury & Lasota (2001) show that including the effects of heating by the stream-disc
1.6 Superhumps

A phenomenon which has become ubiquitous among CVs and is now being observed in many black hole systems is the superhump, a quasi-periodic luminosity variation with period, $P_{sh}$, a few percent longer than the orbital period, $P_{orb}$.

Superhumps were first observed (Vogt 1974; Warner 1975) in SU UMa systems, a sub-class of dwarf novae which mostly show normal outbursts lasting $\sim$ 2–20 days, but which occasionally show superoutbursts which last $\sim$ 5 times longer than normal outbursts (Warner 1995). All SU UMa systems have orbital periods less than 3 hours. Superhumps appear shortly after the peak of the superoutburst, and have become the defining characteristic of superoutbursts, and therefore of SU UMa systems. SU UMa stars in superoutburst exhibit two distinct positive superhump phenomena (Vogt 1983b; Schoembs 1986). Common superhumps appear early in the superoutburst and fade away towards the end of the outburst plateau, to be replaced with late superhumps which persist into quiescence. These late superhumps are roughly anti-phased with the normal superhumps. Figure 1.4 shows the onset of the common superhumps in V1159 Ori from Patterson et al. (1995).

Persistent superhumps are seen in many nova-like systems (Patterson 1998), and are so called because they seem to be present at all times in these systems. Superhumps are also seen in several outbursting low-mass X-ray binaries (O'Donoghue & Charles 1996; Haswell et al. 2001). Low mass X-ray binaries are the X-ray binaries which are most similar to

Figure 1.4: The onset of superhumps in V1159 Ori (from Patterson et al. (1995)).
1.6 Superhumps

CVs, transferring mass via Roche lobe overflow from a low mass donor star onto a compact object, but with a black hole or neutron star rather than a white dwarf.

Some CVs also show *negative* superhumps - photometric signals with periods a few percent shorter than the orbital period. Negative superhumps are permanent, in that they are apparently independent of the eruptive state of the system (Patterson 1999). Negative superhumps are thought to result from the presence of a tilted or warped accretion disc with a regressing line of nodes. Negative superhumps are not discussed in this thesis, and so any reference to superhumps should be read as a reference to positive superhumps, where the superhump period is longer than the orbital period.

The standard model for positive superhumps is that they are caused by the interaction of the donor star with a precessing non-axisymmetric disc. If the accretion disc extends out far enough, the outermost orbits of disc matter can resonate with the tidal influence of the secondary star as it orbits the system. A 3:1 resonance can occur, where particles at the outer edge of the disc orbit the white dwarf three times for every one donor orbit, resulting in the disc becoming distorted to form an eccentric non-axisymmetric shape. The tidal forces acting on this eccentric disc will cause it to precess slowly in a prograde direction. This model was firmly established following simulations of tidally distorted discs (Whitehurst 1988b; Whitehurst & King 1991; Lubow 1991a; Lubow 1991b).

The superhump period, $P_{sh}$, is then given by

$$\frac{1}{P_{sh}} = \frac{1}{P_{orb}} - \frac{1}{P_{prec}},$$

(1.7)

where $P_{prec}$ is the disc precession period. $P_{sh}$ is the period on which the relative orientation of the line of centres of the two stars and the eccentric disc repeats. An interaction between the precessing disc and something moving with the donor then leads to the superhump emission. There are currently two physical mechanisms for producing the superhump modulations.

In the tidal model, the superhump is a result of tidal stresses acting on the precessing non-axisymmetric disc (Whitehurst 1988b; Murray 1996; Murray 1998). The shape and orientation of the disc change throughout the superhump cycle as it is tidally stressed by the gravitational influence of the donor star. Viscous dissipation will therefore lead to emission modulated at the superhump period. This is thought to be the source of common superhumps.

The bright spot model arises from noting that the energy gained by material in the accretion stream will depend on how far it falls before impacting on the disc (Vogt 1982).
The energy dissipated at impact will be modulated on the superhump period since the non-axisymmetric outer disc edge causes a stream-disc impact region at varying depths in the white dwarf potential well. In Chapter 5 a more detailed version of this model is developed. It is thought that late superhumps are a result of this mechanism, and strong evidence for this is presented in Chapter 5.

The disc is truncated at radius \( \sim 0.9R_L \) by the gravitational influence of the donor star (Paczynski 1977) (where \( R_L \) is the radius of the white dwarf Roche lobe). The 3:1 resonance only lies within this tidal disc radius for mass ratios \( q \) less than about 0.25 if the disc and accretion stream are in equilibrium, although it is possible for the disc to expand to the 3:1 resonance radius in systems with \( q \) as high as 0.33 if the mass transfer rate suffers a sustained drop below its equilibrium value (Murray, Warner & Wickramasinghe 2000). There is a strong correlation between the donor mass and the orbital period of CVs (see e.g. Smith & Dhillon 1998). In addition, most white dwarfs in CVs have mass \( 0.6-0.8M_\odot \). This upper limit on \( q \) then explains why the SU UMa systems have periods less than 3 hours. Most persistent superhumpers (nova-likes) also have periods below 3 hours, although there are a few with periods as high as 3.75 hours.

The presence of superhumps in all superoutbursts of SU UMa type CVs suggests a link between the underlying cause of these two phenomena. Osaki (1989) suggests a combination of the tidal instability which causes superhumps with the thermal instability responsible for outbursts. If less mass is emptied from the disc during each normal outburst than is added to the disc between outbursts, then from normal outburst to normal outburst the disc mass and radius will increase. The disc radius increases during outburst as some material moves outwards carrying the angular momentum of the inward flowing material. Eventually the disc radius will be sufficient that as the disc spreads out during outburst, the 3:1 resonance is excited leading to increased tidal-removal of angular momentum from the outer disc. Enhanced viscous dissipation in the outer disc and the increased removal of angular momentum prolong the outburst, producing a superoutburst. This is confirmed by the superoutburst simulations of Truss, Murray & Wynn (2001).

This thesis includes detailed studies of persistent superhumps in V348 Pup (Chapter 3) and late superhumps in IY UMa (Chapter 5). The various models for superhumps are discussed further in these chapters.
1.7 Optical observations of disc-accreting CVs

Cataclysmic variables exhibit a wide variety of observable behaviours on many different timescales. Outlined below are those exploited in this thesis.

1.7.1 Lightcurves

The inclination of the orbital plane of a binary star to the observer's line of sight drastically affects how the system appears, and can lead to a variety of phenomena valuable in unraveling the physics at work in these systems. The orbital inclination, $i$, of a binary system is defined as the angle between the axis of orbital rotation and the line of sight to the observer. We see a low inclination system, where $i$ is close to 0°, from almost directly above the orbital plane looking down at the disc. In a high inclination system, where $i$ is closer to 90°, we see the disc nearer to being edge on. This is illustrated in Figure 1.5.

![Inclination 0° and 65°](image)

Figure 1.5: A CV seen at orbital inclination 0° and 65°.

In systems with an inclination greater than about 70° the donor star will pass in front of the disc and white dwarf during the orbit, leading to eclipses in the lightcurve once every orbit\(^2\). Eclipses provide the most valuable opportunities for studying CVs. Not only does the eclipse enable us to determine easily and accurately the orbital period of a system, it also provides information on the spatial distribution of emission. As the donor star passes in front of the emitting regions, it occults different regions at different orbital phases. This

\(^2\)The inclination for which eclipse occurs depends on mass ratio. For $q=0.3$, the compact object will be eclipsed provided $i > 75°$.\n
1.7 Optical observations of disc-accreting CVs

1.7.1 Emission

Emission lines in the spectra of CVs are one of our most useful tools in understanding the physics of accretion flow. The relative strengths of emission lines from different elements and species provide information about the physical conditions, such as temperature and density, of the emitting material. The velocity of the line emitting region leads to Doppler shifts in the observed wavelength of a line. In high inclination systems these Doppler shifts are much larger than the intrinsic widths of the emission lines, so that the shape of the observed line profile depends only on the orbital phase and velocity distribution emitting material. This dependence is exploited by the method of Doppler tomography to map the velocity-space distribution of the emission-line regions (see Section 2.2.2). The most obvious manifestation of this behaviour is the double-peaked line profile from accretion discs, e.g. in IY UMa (Section 7). The material on one side of the accretion disc has a projected velocity towards the observer, and so is blue-shifted. Emission from the other side of the disc will have a velocity away from the observer, and so is red-shifted. This leads to emission line profiles showing two peaks - one from the red-shifted material and the other from the blue-shifted material. A concentrated region of emission in velocity space will be seen as an “S-wave” in the phase-dependent line profile - a single peak in the line profile whose radial velocity varies sinusoidally with time as it is projected onto the line of sight from the rotating binary frame. A common source of S-waves in CV spectra is the hotspot at the stream-disc impact, as seen in IY UMa (Chapter 7). Not all disc-accreting CVs show double-peaked emission. One particular group of systems with
several unusual emission line properties are the so-called SW Sex stars, first recognised by Thorstensen et al. (1991). They are thought to possess discs and yet show broad, single-peaked emission lines rather than the double-peaked disc emission usually seen in high inclination systems. Chapters 4 and 7 make extensive use of emission line profiles to study the accretion flow in two different CVs: V348 Pup, a new member for the SW Sex class, and IY UMa, a short period system with a very high inclination which presents a clear view of the stream-disc impact.

The spectra of CVs often show absorption lines from various different components of the system, notably infrared absorption features from the donor star and UV absorption from the white dwarf. These features are our only way of directly measuring the velocities of the donor and white dwarf, providing tight constraints on the system parameters. We exploit absorption features in the near infrared to identify the donor star in IY UMa (Chapter 6).
Chapter 2

Data reduction and analysis techniques

In this chapter I briefly outline the techniques for reducing CCD-based spectroscopic observations, as employed in the work presented in this thesis. Apart from the overscan and bias corrections described in Section 2.1.1, the techniques for reducing CCD-based photometry are different. I also introduce the indirect imaging techniques of Doppler tomography and eclipse mapping.

2.1 Reduction of CCD spectroscopy

Charge-coupled devices (CCDs) are two-dimensional array semiconducting devices, with each pixel recording (in electrons) a measure of the number of photons incident on that pixel during an exposure. The number of electrons in each pixel after an exposure is read out electronically, producing a two-dimensional image of the electrons per pixel. In the case of spectroscopic images, the aim of data reduction is to convert this two-dimensional image to a one-dimensional spectrum, providing a calibrated flux value as a function of wavelength. The data reduction described in this thesis was carried out using IRAF and its associated packages, unless otherwise stated.

2.1.1 Overscan correction and bias subtraction

The readout and analogue-digital conversion of CCD exposures introduces noise into the recorded counts per pixel. On readout, the CCD signal is automatically given a positive offset so that this readout noise never causes a negative signal. This bias is then subtracted in software during the reduction process. Bias frames are taken with the shutter closed.
2.1 Reduction of CCD spectroscopy

and zero exposure time. These images reveal the bias signal over the entire chip. An overscan region is also generated in which the CCD sequencer sends more readout pulses than are required to read the CCD. The resulting extra rows or columns in the CCD image will provide information on how the bias level varies from one frame to the next. The bias frames are averaged and subtracted from all CCD images, then the average value of, or a fit to, the overscan region of each image is subtracted.

2.1.2 Flat-field correction

The responses of different pixels on a CCD will not be identical. Some will be more or less sensitive than others. To account for this images are taken in which the chip is illuminated uniformly. The resulting flat-field images should reveal the non-uniform pixel response across the chip. In the dispersion direction it should also vary due to the spectrum of the flat-field lamp and the wavelength dependence of the CCD sensitivity.

The flat-field images are averaged and fitted with a smooth function in the dispersion direction. This function represents the variation across the flat-field due to the spectrum of the flat-field lamp. Thus dividing the flat-fields by this variation (and normalizing) should leave just the variation across the chip which results from non-uniform response of the CCD pixels.

These corrected flat-field images are divided into all object, sky and calibration images, thus correcting them for the sensitivity variation across the CCD.

2.1.3 Illumination correction

To see how effective the flat fielding is, the variation of the sky flat images in the spatial direction is inspected. Sky flat images are taken with the slit in the optical path, but with the telescope pointing at a blank region of sky. With no objects in the field of view, a sky flat should only vary in the dispersion direction. However, irregularities in the slit in the spatial direction can cause some spatial variations in the observed spectra. To correct for this, the average sky flat is broken into several equal-sized bins in the dispersion direction. Each of these is then collapsed in that direction and the spatial variation is fitted with a smooth function. The resulting wavelength-dependent illumination correction is used to flatten all object and calibration images in the spatial direction.
2.1 Reduction of CCD spectroscopy

2.1.4 Spectrum extraction

In order to extract the spectra from the CCD frames, the aperture is located and its path across the chip traced and fitted with a polynomial. The aperture is the region of the CCD which contains the spectrum of the object. It is a band extending across the object images in the dispersion direction, with an extent in the spatial direction determined primarily by seeing and focusing. Figure 2.1 shows an example CCD image from the Nordic Optical Telescope observations discussed in Chapter 7, which shows two spectra on the slit. The aperture for the brightest spectrum is centred around column 65, while the weaker spectrum is centred around column 15.

![Example object frame showing two spectra on the slit.](image)

The centre of each aperture is determined (traced) as a function of dispersion direction, with the software (APTRACE in IRAF) averaging a number of rows in the dispersion direction to increase signal-to-noise if necessary. A smooth function is then fitted to this aperture trace (the centre of the aperture), to avoid 'noise' in the measured trace. Regions defining the sky and aperture relative to the trace are chosen for some wavelength, and thus using the fitted trace, the aperture and sky are defined along the entire wavelength range. The spectra are then extracted by fitting and subtracting the sky before summing over the aperture, for each position in the dispersion direction. In this thesis, the IRAF implementation of the Horne (1986) optimal extraction algorithm is used. This algorithm assigns weights to the pixels when summing over the aperture. These weights are chosen to optimize the signal-to-noise ratio while maintaining the spectrophotometric accuracy of the extracted spectrum. The method also eliminates most cosmic rays from the spectrum. Those remaining are easily identified and removed manually by interpolating over affected
2.1 Reduction of CCD spectroscopy

2.1.5 Dispersion solution

The dispersion solution is a function which gives the wavelength as a function of position in the dispersion direction of the image. This solution is obtained by taking spectra in which the slit is illuminated using an arc lamp. The emission lines from the arc lamp have accurately known wavelengths. The arc lamp exposure is extracted using the aperture of the object image to be calibrated. The arc lines in the extracted spectrum are identified and assigned their known wavelengths, and a smooth function is fitted to the wavelength as a function of position. This smooth function can then be used to transform pixel number to wavelength for the object image. Arc lamp exposures are usually taken at each position of the telescope, with object spectra being calibrated using arc spectra from the same telescope position. This prevents the accuracy of the dispersion solution being degraded by slight changes in alignment of the instrument as the telescope moves.

2.1.6 Flux calibration

By taking exposures of flux standard stars (bright objects whose spectra have already been accurately calibrated), the wavelength dependent sensitivity of the instrument can be found. By summing the counts in several broad wavelength bins in the observed spectra of the flux standards, and comparing these to the fully calibrated standard star spectra, a wavelength dependent sensitivity function is calculated. The wavelength-dependent atmospheric extinction is also accounted for when calculating this sensitivity function using known tables of extinction for each observing site. The sensitivity function is then used to place all object spectra on a true flux scale. Note that unless the slit width is much wider than FWHM of the source, slit losses will occur. This is where inaccuracies in telescope pointing and focus, and variations in seeing, lead to a varying fraction of the object flux getting through the slit. I describe in each chapter how (or if) slit losses were corrected for.

2.1.7 Telluric correction

Molecules of O$_2$ and H$_2$O in the atmosphere lead to the existence of various absorption lines and bands in spectra obtained using ground-based telescopes. These telluric features are common in the infrared, and also affect the red end of optical spectra. Correction can
be made for this if spectra of a hot featureless star are taken. This is described in Chapter 6 where it is applied.

2.2 Indirect imaging techniques

Cataclysmic variables are one of our most valuable tools for understanding accretion processes, but they are far too small to resolve directly. With typical binary separations of a few solar radii, and at typical distances of 100 pc, they subtend $<10^{-4}$ arc seconds at the Earth, far smaller than can be resolved by even the largest telescopes. We therefore have to employ indirect imaging techniques to unravel the structure of CVs, and in this section I briefly introduce two of the most common such techniques, both of which are employed in this thesis.

2.2.1 Eclipse mapping

As described in Section 1.7, high inclination systems show eclipses as the donor star blocks our view of the emitting regions, and the profiles of these eclipses provide us with information about the spatial distribution of emission in the system. Unfortunately, eclipses provide us with only one dimensional information about a 2- (or more realistically 3-) dimensional emission distribution. This means that there is more than one possible emission distribution which will produce the same eclipse profile. It is therefore impossible to devise an algorithm which can take the eclipse shape and unambiguously obtain the emission distribution. However, Horne (1985) developed a technique for exploiting the information in eclipse profiles to obtain a map of the emission distribution in the orbital plane - the maximum entropy eclipse mapping technique. This works by placing an extra constraint on the emission distribution, in addition to requiring that it is consistent with the observed eclipse profile. This additional constraint usually requires that the emission distribution, or eclipse map, is as smooth as possible, or as axisymmetric as possible. By placing such a constraint on the eclipse map, we are effectively choosing which one of the many maps consistent with the observations is our preferred solution. Here I outline the basic principle of maximum entropy eclipse mapping, following closely the description in the recent review by Baptista (2001). We assume a map of emission which is a square grid in the orbital plane, rotating with the binary, and centred on the white dwarf. The grid has $N$ pixels of equal area, with pixel $i$ having constant brightness.
2.2 Indirect imaging techniques

The total flux seen by an observer at orbital phase $\phi$ will be

$$F(\phi) = \sum_{i=1}^{N} I_i V_i(\phi). \quad (2.1)$$

$V_i(\phi)$ is the visible fraction of pixel $i$ at orbital phase $\phi$, equal to 0 when the pixel is eclipsed and 1 when it is fully visible. This can easily be calculated if the mass ratio and orbital inclination of the system are known, and if we assume the donor star to fill its Roche lobe exactly. The consistency of the map with the observed lightcurve is tested using the reduced $\chi^2$ constraint function, defined as

$$\chi^2 = \frac{1}{M} \sum_{i=1}^{M} \left( \frac{F(\phi) - L(\phi)}{\sigma(\phi)} \right)^2, \quad (2.2)$$

where $M$ is the number of points in the lightcurve, $L(\phi)$ is the observed flux at phase $\phi$ and $\sigma(\phi)$ is its associated error. For a good fit, $\chi^2$ should be about 1. The additional constraint of smoothness/symmetry placed on the map is provided through the entropy, $S$, defined as

$$S = -\sum_{i=1}^{N} p_i \ln \left( \frac{p_i}{q_i} \right), \quad (2.3)$$

where

$$p_i = \frac{I_i}{\sum_{j=1}^{N} I_j} \quad \text{and} \quad q_i = \frac{D_i}{\sum_{j=1}^{N} D_j}.$$ 

$D_i$ are brightnesses of pixels in the default map, defined as

$$D_i = \frac{\sum_{j=1}^{N} \omega_{ij} I_j}{\sum_{j=1}^{N} \omega_{ij}}. \quad (2.4)$$

The default map is therefore a weighted average of the eclipse map, with the weights, $w_{ij}$. By maximizing the entropy $S$, subject to the consistency constraint that $\chi^2 = 1$, the final eclipse map will be that which is closest to the default map while remaining consistent with the observed lightcurves. The default map is recalculated from the eclipse map at each iteration of the fitting procedure. The definition of $w_{ij}$ prescribes our preferred final solution. A choice of $w_{ij} = 1$ produces a uniform default map, leading to a final eclipse map which is as uniform as possible within the constraints provided by the data. A variety of different choices of $w_{ij}$ are listed in Baptista (2001). The definition which is used in Chapter 3 is

$$w_{ij} = \exp \left[ -\frac{1}{2} \left( \frac{(R_i - R_j)^2}{\Delta R^2} + \frac{s_{ij}^2}{\Delta s^2} \right) \right]. \quad (2.5)$$

$R_i$ and $R_j$ are the distances between pixels $i$ and $j$ and the white dwarf, while $s_{ij}$ is the arc length between pixels $i$ and $j$ (centred on the white dwarf). This definition
effectively leads to an eclipse map which is as smooth as possible radially on a scale of $\Delta R$ and azimuthally on a scale $\Delta s$, while remaining consistent with the observations. By setting $\Delta R$ to 0.8 pixels and setting $\Delta s$ to correspond to an angle of $360^\circ$ at the edge of the map in Chapter 3, we looked for the most axisymmetric map consistent with observed lightcurves.

Baptista (2001) reviews the eclipse mapping method, along with various examples of its use.

2.2.2 Doppler tomography

Another indirect imaging technique used extensively to study accretion flow in binary systems is Doppler tomography, developed by Marsh & Horne (1988). This technique extracts the velocity information provided in the emission line profile variations to obtain a velocity-space image of the line emission distribution. A recent review of Doppler tomography, or Doppler mapping, is provided by Marsh (2001). The intrinsic width of the emission line is assumed to be small compared to the Doppler shifts resulting from the velocity of the emitting material, so that the line profile depends only on the velocity distribution of the emission regions. The emission distribution is assumed to be constant, and located in the orbital plane. Each point in velocity space is assumed to be equally visible at all orbital phases. This last assumption means that Doppler tomography cannot account for occultation effects such as eclipses.

I now outline the description of Doppler tomography in Marsh (2001). The coordinate axes, $x$, $y$ and $z$, follow the usual definition as described in Section 1.2 and shown in Figure 1.1 (where $M_1$ and $M_2$ are the white dwarf and donor respectively). The velocity-space of the Doppler map, $(V_x, V_y)$, is the velocity measured in an inertial frame with the same orientation as the binary. $V_x$ and $V_y$ are the $x$ and $y$ velocities we would measure in an inertial frame at orbital phase zero (when the donor is closest to the observer). $(0,0)$ is the velocity of the centre of mass. The radial velocity of the point $(V_x, V_y)$ at orbital phase $\phi$ will be

$$V_R = \gamma - V_z \cos(2\pi\phi) \sin i + V_y \sin(2\pi\phi) \sin i,$$  \hspace{1cm} (2.6)

where $\gamma$ is the systemic velocity and $i$ is the orbital inclination. Positive $V_R$ is directed away from the observer. If the velocity-space image of the emission line distribution is $I(V_x, V_y)$, then the line profile at velocity $V$ relative to the line centre at phase $\phi$ is an
2.2 Indirect imaging techniques

Integral of the Doppler map over velocity space,

\[ f(V, \phi) = \int_{-\infty}^{\infty} \int_{-\infty}^{\infty} I(V_x, V_y) g(V - V_R) dV_x dV_y, \]

(2.7)

where \( g(V) \) is the intrinsic line profile, usually assumed to be the same at every point. \( g(V) \) is assumed to be narrow, so \( f(V, \phi) \) is effectively an integral of the velocity-space image (the Doppler map) along a straight line defined by \( V_R(\phi) = V \) in Equation 2.6. Figure 2.2 illustrates this. The greyscale image in the centre is the Doppler map, with darkest regions corresponding to strongest emission. The line profile at orbital phase 0.5 is simply the map summed in the \( V_y \) direction, while the map at phase 0.25 is a sum in the \( V_x \) direction. Line profiles at other phases are just projections in other directions. If the emission distribution is constant in time, as assumed, then a line profile at phase \( \phi \) is a mirror image of that at phase \( \phi + 0.5 \). If \( f(V, \phi) \) is known for all \( V \) and \( \phi \), then it

\[ 0^\circ.0 = \phi \]
\[ 1 \text{ ms}/\text{arc A} \]

Figure 2.2: Line profiles projected from the Doppler map as a function of orbital phase.

is possible to unambiguously obtain \( I(V_x, V_y) \). In practice, since observations never have
full coverage of \( f(V, \phi) \), alternative techniques are used to obtain \( I(V_x, V_y) \).

The \textit{filtered back-projection} technique reconstructs the Doppler map by first filtering each line profile (see Marsh 2001), and then back-projecting each profile into velocity-space. The back-projection step is done by smearing each profile across the map in the same direction as the map is summed to produce that profile. So the line profile at phase 0.25 is smeared across the map in the \( V_x \) direction (see Figure 2.2).

The maximum-entropy method (Marsh & Horne 1988) uses the same principle as eclipse mapping described in the previous section. The entropy of the Doppler map is defined using Equation 2.3. The \( \chi^2 \) statistic is employed, this time comparing model line profiles predicted from the map to the observed line profiles. By maximizing entropy subject to the constraints of the observation imposed through \( \chi^2 \), a Doppler map is obtained. The weights, \( w_i \) (as used in eclipse mapping), are usually chosen to produce a blurred default image, so that the Doppler map is smooth on scales shorter than the blur size. Note that the definition of entropy (Equation 2.3) prevents the use of this method if the Doppler map contains negative values (absorption). This method is less prone to effects of incomplete phase or velocity coverage than back-projection.

In this thesis, maximum entropy Doppler maps are produced for V348 Pup (Chapter 4) using Tom Marsh’s \textsc{doppler} software. However, filtered back-projection is employed to produce the maps of IY UMa (Chapter 7) because there are strong absorption features which lead to negative values in the maps. The back-projections were done using the implementation in Tom Marsh’s \textsc{molly} software.

\textit{Interpreting Doppler maps}

The model Doppler map in Figure 2.2 has several features plotted on top to help us understand the structure of the map. The small teardrop shape is the velocity of the donor Roche lobe projected into the frame of the map. We would expect any emission from the donor star to appear within this teardrop. The larger broken teardrop is the velocity of the primary Roche lobe. It is not expected to correspond to any source of emission, but I occasionally find it useful as a reference when describing Doppler images. Where the teardrops meet is the velocity of the L1 point. The arc on the left starting at the L1 point shows the expected velocity of a ballistic accretion stream from the donor, with the open circles denoting steps of 0.1\( a \) along the stream. The other arc shows the Keplerian velocity along the stream trajectory, with the filled circles corresponding to
the open circles on the stream velocity. Emission from the stream-disc impact should lie somewhere between these two arcs, as the stream impacts the disc edge and gradually merges with it. This is seen as the dark blob of emission in the model Doppler map. The model map also shows a ring of emission centred on the white dwarf and passing through the hotspot velocity, with fainter emission outside. The dark ring is emission from the outer radii of a Keplerian disc, with the fainter emission at greater velocities coming from regions of the disc closer to the white dwarf. This shows how a Keplerian disc is turned ‘inside-out’ in velocity space, with the outer parts of the disc, which have low velocity, concentrated in a narrow range of velocities, leading to the dark ring. The black circle centred on the white dwarf has a radius equal to the Keplerian velocity at the tidal truncation radius of the disc. No emission is expected from a Keplerian disc within this circle.
Chapter 3

Superhumps in V348 Pup

3.1 Introduction

V348 Pup (1H 0709-360, Pup 1) is a nova-like cataclysmic variable: a system with a high mass transfer rate which maintains its accretion disc in the hot, ionized, high viscosity state reached by dwarf novae in outburst. It exhibits deep eclipses in its optical and infrared lightcurves (Tuohy et al. 1990): it is a high inclination system with orbital period $P_{\text{orb}} = 2.44$ hours (Baptista et al. 1996). The work described in this chapter is based on the work published in Rolfe, Haswell & Patterson (2000).

3.2 Observations

The observing campaign comprises 23 nights of rapid photometry from December 1991, February 1993 and January 1995 (see Table 3.1). The observations were carried out by Prof. Joe Patterson, Jonathan Kemp and Jessica Zimmerman. The 1991 and 1993 observations (8 and 11 nights respectively) were taken using the 40-inch telescope at CTIO with a blue copper sulphate filter. The January 1995 run consists of 4 nights of R band data. All the data have been corrected for atmospheric extinction and the 1995 data have also been calibrated to give an absolute flux. In 1995 the average out of eclipse R magnitude is 15.5 mag; at mid-eclipse $R = 16.8$. All three datasets are plotted in Figure 3.1. Before performing any analysis, each night of data was normalized by dividing by the average out-of-eclipse value.
Figure 3.1: The entire December 1991, February 1993 and January 1995 datasets. Vertical axis shows normalized flux.
### Table 3.1: Log of observations

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3.3 The orbital ephemeris

The mid-eclipse timings were determined as described in Section 3.8 from which an orbital ephemeris was determined for this dataset:

\[ T_{\text{mid}} = \text{HJD} \ 2448591.667969(85) + 0.101838931(14)E. \]

This is consistent with the Baptista et al. (1996) ephemeris within the quoted error limits. The above ephemeris is adopted for this analysis. The eclipse timings are given in Table 3.2.

3.4 Periodic signals in the lightcurves

3.4.1 Period searching and error estimation

To make detection of non-orbital modulations in the data easier, the average orbital lightcurves\(^1\) from each year’s observations were calculated and subtracted from the corresponding data. The resulting lightcurves contain no orbital variations, as can be seen from the power spectra: there is no power at the orbital frequency in Figures 3.2 and 3.3.

Lomb-Scargle periodograms were calculated for each year’s data, and are shown in Figures 3.2 and 3.3. The Lomb-Scargle method for calculating power spectra is designed to perform spectral analysis on unevenly sampled data, and is invariant to shifting the times of all the data points by the same constant amount. The Lomb-Scargle statistic for a particular frequency \( \omega \) is equivalent to estimating the strength of the harmonic at this frequency by carrying out a linear least-squares fit of the form \( f(t) = A \cos \omega t + B \sin \omega t \) (Lomb (1976)). The implementation of the Press & Rybicki (1989) algorithm in the PERIOD software package was used.

It is notoriously difficult to determine errors in periods measured from periodograms. To estimate the errors in the periods detected here, various fake datasets were generated. The lightcurves were smoothed (with phase width 0.02) and the residuals used to characterize the variance, \( \sigma_{\text{obs}}^2 \), of the random noise. Smoothing on this timescale will not smooth out the intrinsic flickering seen in the lightcurves. The flickering is therefore not present in the residuals, and so not accounted for in these error estimates. Five different sets of fake noise were generated as described below.

- Uniform distribution noise with variance \( \sigma_{\text{obs}}^2 \)

\(^1\)Superhump phase grouped average lightcurves are shown in Figure 3.10.
### Table 3.2: Eclipse timings

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<th>HJD mid-eclipse -2440000</th>
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</table>
3.4 Periodic signals in the lightcurves

- Normal distribution noise with variance $\sigma_{obs}^2$
- Poisson distribution noise with variance $\sigma_{obs}^2$
- The observed noise for each night reversed in time
- The observed noise for each dataset randomly re-ordered

Adding each of these sets of fake noise to the smoothed dataset produces five fake datasets for each observed one. By repeating a smoothed superhump modulation (obtained as described in Section 3.6) over the time series of the observations and adding each fake noise set to this, a further five fake datasets were produced. The uncertainty in each detected superhump period was then determined from the scatter in the 11 periods obtained from period searching the real data and the 10 fake datasets. Errors in the sideband periods (Section 3.4.3) were estimated using the real data and the first five fake datasets only. Since the smoothed data will still contain noise artefacts these are probably underestimates.

3.4.2 The superhump period

The 1991 periodogram reveals a periodicity with period 0.10763 days and simple aliasing structure. The 1993 periodogram has higher resolution, more complicated alias structure, and the strongest periodicity at 0.10857 days. The 1995 power spectrum (having the least time coverage) is lower resolution, however a clear signal at 0.10760 days and its aliases is present. By comparison with the clearer 1991 and 1993 spectra, it was surmised that the 0.10760-day peak is the true signal. These periods are all close to 6 per cent greater than $P_{orb}$ and correspond to the period excesses, $\epsilon$, shown in Table 3.4, defined using superhump period, $P_{sh} = (1 + \epsilon)P_{orb}$. The inferred disc precession periods are also shown in Table 3.4.

The 1993 dataset with its 20-day time base should provide the most precise period. Therefore, the possibility that the real 1991 and 1995 superhump periods are in fact closer to the 1993 value than these measurements suggest should not be ruled out, as Figure 3.3 shows. However, variability in the detected periodicity has also been seen before for superhumps (Patterson 1998).

Superhumping systems show a clear correlation between $\epsilon$ and $P_{orb}$, with $\epsilon$ increasing with $P_{orb}$; the superhump periods detected here in V348 Pup are consistent with this trend (Figure 3.4).
Figure 3.2: Power spectra. Arrows from left to right indicate the superhump period and the sidebands of the first and second harmonics of the orbital period (see Table 3.5).
Figure 3.3: Detail of the power spectra shown in Figure 3.2. Arrows indicate the superhump period. The orbital frequency \( \Omega_{\text{orb}} \) is marked, from which it is clear that the orbital variation has been successfully removed. The predicted period at \( \Omega_{\text{orb}} + \Omega_{\text{prec}} \) is also marked (see Section 3.4.3).
3.4 Periodic signals in the lightcurves

Figure 3.4: The superhump period excess, $\epsilon$, as a function of orbital period. The maximum and minimum values for V348 Pup (the 1993 and 1995 values from Table 3.4) are marked with the stars. Other values are from Patterson (1998), Patterson et al. (2000), Nogami et al. (2000) and Kiyota & Kato (1998).

3.4.3 Other periodic signals

The power spectra also reveal signals at frequencies corresponding to sidebands of harmonics of the orbital period. The strongest such detections are at periods corresponding to $2\Omega_{\text{orb}} - \Omega_{\text{prec}}$ and $3\Omega_{\text{orb}} - \Omega_{\text{prec}}$ (marked with arrows in Figure 3.2): the predicted and the directly measured values of these sidebands are shown in Table 3.5, and graphically in Figure 3.5. The errors in these periods were estimated using the method described earlier. The detected periods do not all agree to within the estimated errors, suggesting that the error estimates may be a little too low, as expected; the highest quality 1993 data agrees best.

The simplest way to produce these sidebands is by modulating the brightness or visibility of the superhump with orbital phase. If we consider the eclipse of the superhump light source as a product of a visibility function, $V(t)$, and the intrinsic variation in the superhump light source, $S(t)$, we can follow the approach of Warner (1986) and Norton, Beardmore & Taylor (1996) to studies of signals in magnetic CVs. We treat $V(t)$ as a sum of Fourier components, $V(t) = \sum_{n=1}^{\infty} v_n \cos n\Omega_{\text{orb}} t$, noting that the addition of a $\sin n\Omega_{\text{orb}} t$
3.4 Periodic signals in the lightcurves

term would complicate the maths without affecting the resulting frequencies. The intrinsic superhump variation is treated as \( S(t) = s \cos \Omega_{sh} t \), again ignoring the \( \sin \Omega_{sh} t \) term. The resulting lightcurve of the superhump including the effect of eclipse is then simply

\[
F(t) = V(t) S(t) = \sum_{n=1}^{\infty} sV_n \cos n\Omega_{orb} t \cos \Omega_{sh} t.
\]

Using basic trigonometric identities and the fact that \( \Omega_{sh} = \Omega_{orb} - \Omega_{prec} \) we get

\[
F(t) = \frac{1}{2} \sum_{n=1}^{\infty} sV_n \cos[(n+1)\Omega_{orb} - \Omega_{prec}]t + sV_n \cos[(n-1)\Omega_{orb} + \Omega_{prec}]t.
\]

This shows how the eclipse of the superhump light source will produce signals at frequencies \( (n + 1)\Omega_{orb} - \Omega_{prec} \). However, signals at frequencies \( (n - 1)\Omega_{orb} + \Omega_{prec} \) are also predicted, yet there is no evidence of these. Table 3.3 lists the predicted frequencies and coefficients for \( n=0-2 \).

<table>
<thead>
<tr>
<th>( n )</th>
<th>( V_n )</th>
<th>Frequency 1</th>
<th>Frequency 2</th>
</tr>
</thead>
<tbody>
<tr>
<td>0</td>
<td>( V_0 )</td>
<td>( \Omega_{sh} )</td>
<td>( -\Omega_{orb} + \Omega_{prec} )</td>
</tr>
<tr>
<td>1</td>
<td>( V_1 )</td>
<td>( 2\Omega_{orb} - \Omega_{prec} )</td>
<td>( \Omega_{prec} )</td>
</tr>
<tr>
<td>2</td>
<td>( V_2 )</td>
<td>( 3\Omega_{orb} - \Omega_{prec} )</td>
<td>( \Omega_{orb} + \Omega_{prec} )</td>
</tr>
</tbody>
</table>

For \( n=0 \), the predicted signals are the superhump (which we see) and a signal with negative frequency which we therefore do not expect to see. For \( n=1 \) the signals predicted have frequency \( 2\Omega_{orb} - \Omega_{prec} \) (which we also see) and the precession frequency \( \Omega_{prec} \). There is no significant signal in the observations around the precession period, but the observations will be insensitive to such low frequency signals since changes in calibration from night to night and changing atmospheric conditions can affect the observations on these timescales. The periodograms in Figure 3.2 show much spurious structure at low frequencies, particularly that for 1995. For \( n=2 \), the observed signal at \( 3\Omega_{orb} - \Omega_{prec} \) is predicted, as is a signal at \( \Omega_{orb} + \Omega_{prec} \). This frequency is indicated with arrows in Figure 3.3. The strength of this signal should be similar to that of the \( 3\Omega_{orb} - \Omega_{prec} \) since both have the coefficient \( V_2 \). This signal should therefore be much weaker than the superhump signal, and would not be identifiable in the 1991 and 1995 periodograms where it would
be lost in the 1-day alias peaks of the superhump period. We might expect to see it in the 1993 periodogram, which has narrower aliases, but it is not apparent, however it might be that the signal is simply too weak to be detected so close to peaks from a much stronger signal.

In Section 3.7 evidence is presented for a correlation between superhump amplitude and orbital phase; this modulation of superhump amplitude with orbital phase would also lead to the observed sideband signals, following exactly the same treatment as detailed above. This would also predict the signals \((n - 1)\Omega_{\text{orb}} + \Omega_{\text{prec}}\) frequencies which are not observed.

The SPH models of Simpson & Wood (1998) (and many others - see Chapter 1.4) predict the formation of double armed spiral density waves in the disc whose rotation rate, they suggest, might lead to observed signals at about three times the superhump frequency. They also suggest that viewing these structures from non-zero inclination could lead to the detection of further frequencies, although they do not make precise predictions. Observations of the dwarf nova IP Peg in outburst have revealed evidence of such spiral structure (Steeghs, Harlaftis & Horne 1997). Several other systems also show evidence for spiral waves. These are summarized in Section 7.7.2, where a possible detection of spirals waves in IY UMa is presented.

Table 3.4: Measured superhump characteristics, period \(P_{\text{sh}}\), the period excesses \(\epsilon\), the inferred disc precession periods \(P_{\text{prec}}\) and the consequent estimates of the mass ratio \(q\).

<table>
<thead>
<tr>
<th>Year</th>
<th>(P_{\text{sh}}) (d)</th>
<th>(\epsilon)</th>
<th>(P_{\text{prec}}) (d)</th>
<th>Fractional amplitude</th>
<th>(q)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1991</td>
<td>0.107628(8)</td>
<td>0.0568</td>
<td>1.89</td>
<td>0.20</td>
<td>0.26</td>
</tr>
<tr>
<td>1993</td>
<td>0.108567(2)</td>
<td>0.0661</td>
<td>1.64</td>
<td>0.15</td>
<td>0.31</td>
</tr>
<tr>
<td>1995</td>
<td>0.10760(7)</td>
<td>0.0566</td>
<td>1.90</td>
<td>0.13</td>
<td>0.26</td>
</tr>
</tbody>
</table>

3.5 The orbital parameters

The mass ratio, \(q\), of a superhumper can be estimated, given its period excess, \(\epsilon\), following Patterson (1998). The classical precession rate of an elliptical disc calculated by Osaki
3.5 The orbital parameters

Table 3.5: Sidebands of the orbital period harmonics

<table>
<thead>
<tr>
<th>Year</th>
<th>$\Omega = 2\Omega_{\text{orb}} - \Omega_{\text{prec}}$</th>
<th>$\Omega = 3\Omega_{\text{orb}} - \Omega_{\text{prec}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>1991</td>
<td>0.052327(1)</td>
<td>0.05241(1)</td>
</tr>
<tr>
<td>1993</td>
<td>0.0525477(4)</td>
<td>0.052540(1)</td>
</tr>
<tr>
<td>1995</td>
<td>0.05232(1)</td>
<td>0.052493(8)</td>
</tr>
</tbody>
</table>

Figure 3.5: Sidebands of the orbital period harmonics. Dotted lines mark predicted frequency range for each signal, solid lines show the measured range.

(1985) gives

$$\frac{P_{\text{orb}}}{P_{\text{prec}}} = 3 \frac{q}{4 \sqrt{1 + q}} \left( \frac{R}{a} \right)^{3/2}$$

where $R$ is the outer radius of the disc and $a$ the orbital separation. Using Equation 1.7 and the definition of superhump period excess, $\epsilon$, we get

$$\frac{P_{\text{orb}}}{P_{\text{prec}}} = \frac{\epsilon}{1 + \epsilon}.$$ 

Taking $R$ to be at the location of the 3:1 resonance, $R = 0.46a$ (Osaki 1996), this gives us

$$\epsilon = \frac{0.23q}{\sqrt{1 + q} - 0.23q}$$

which by assuming small $q$ (Patterson 1998) approximates to

$$\epsilon = \frac{0.23q}{1 + 0.27q}.$$  (3.1)
Equation 3.1 leads to the estimates of $q$ for V348 Pup shown in Table 3.4. We favour the mass ratio, $q = 0.31$, estimated from the most precise 1993 superhump period. SPH simulations of eccentric discs (Murray 1998; Murray 2000) suggest a more complicated relationship between $\epsilon$ and $q$ with the disc precession rate depending on the gas pressure and viscosity of the disc in addition to the mass ratio, but this does not affect the substantive results presented here.

Assuming that the secondary star is Roche lobe filling, the width, $w$, of eclipse of a point source at the centre of the compact object uniquely defines orbital inclination, $i$, as a function of $q$. Thus $i$ can be computed as a function of $q$ and $w$.

Figure 3.6: The eclipsed regions of the orbital plane at the start and end of white dwarf eclipse, showing how approximately half of the area of a $0.9R_L$ disc is eclipsed at this phase. This figure uses $q = 0.31$ and $i = 81^\circ 1$.

When the centre of the compact object (point $P$) is first eclipsed (orbital phase $\phi_1$), about half of the disc area will be eclipsed, and therefore for a disc whose intensity distribution is symmetric about the line of centres of the two stars, the fraction of disc flux eclipsed at this phase will be $\sim 0.5$ (Figure 3.6). Similarly at the end of the eclipse of $P$ (orbital phase $\phi_2$) the fraction of disc light visible is again $\sim 0.5$. Further assuming that the lightcurve consists purely of emission from a disc in the orbital plane, the full width of eclipse at half intensity will be equal to $w$.

Using the average eclipses from each dataset to give $w$, $i$ was obtained using both values of $q$ implied by the different values of $\epsilon$. The results are shown in Table 3.6. The error in eclipse width $w$ was estimated by measuring the width at half-flux, and also measuring it at half-flux+$\sigma(f_u)$ where $f_u$ is the out of eclipse lightcurve, and taking the uncertainty as the difference between the two measured widths.

The assumption that half of the disc area is eclipsed at $\phi_1$ and $\phi_2$ is a good one for
3.6 The superhump modulation

To study the form of the superhump modulation, each dataset had its average orbital lightcurve subtracted and was folded and phase-binned onto its detected superhump period. To define a zero point in superhump phase a sine-wave was fitted to each modulation (see Figure 3.7).

To assess the contribution of flickering to these curves the data were binned in three different ways. First, a simple average of the points in each bin was done. Secondly, an average of the lowest 25 per cent of the points in a bin was taken. Also, a weighted average of points in each bin was taken: if the minimum value in a bin is \(a\), the maximum value \(b\) and the points are \(x_i\), then each point is assigned weight

\[
w_i = \left(\frac{b - x_i}{b - a} + 1\right)^3.
\]

Since flickering in the lightcurve consists of brief increases in luminosity, by giving more

<table>
<thead>
<tr>
<th>Year</th>
<th>Full-width at half-flux (w) (minutes)</th>
<th>Mass ratio (q)</th>
<th>Inclination (i)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1991</td>
<td>10.6 ± 0.3</td>
<td>0.31</td>
<td>81.1 ± 0.6</td>
</tr>
<tr>
<td>1991</td>
<td></td>
<td>0.26</td>
<td>82.5 ± 0.7</td>
</tr>
<tr>
<td>1993</td>
<td>10.6 ± 0.5</td>
<td>0.31</td>
<td>81.1 ± 0.8</td>
</tr>
<tr>
<td>1993</td>
<td></td>
<td>0.26</td>
<td>82.4 ± 1.0</td>
</tr>
<tr>
<td>1995</td>
<td>10.7 ± 0.5</td>
<td>0.31</td>
<td>81.2 ± 1.0</td>
</tr>
<tr>
<td>1995</td>
<td></td>
<td>0.26</td>
<td>82.7 ± 1.2</td>
</tr>
</tbody>
</table>

these values of \(q\) and \(i\) and the disc size measured in Section 3.8. However, the average eclipses are slightly asymmetric, so that the mid-phase between the half-flux points of eclipse ingress and egress is not quite zero. This discrepancy is included in the uncertainty in \(w\), and thus propagated into the errors in \(i\). The 1993 data yielded the most precise superhump period, and so the orbital parameters \(q = 0.31, i = 81.1° \pm 1°\) were adopted, using the 1993 determination of \(q\) and all three measurements of \(w\). The conclusions drawn later are not sensitive to the choice of \(q\) provided the inclination is then chosen to be consistent with the observed eclipse width. 

3.6 The superhump modulation

To study the form of the superhump modulation, each dataset had its average orbital lightcurve subtracted and was folded and phase-binned onto its detected superhump period. To define a zero point in superhump phase a sine-wave was fitted to each modulation (see Figure 3.7).
3.6 The superhump modulation

Figure 3.7: The superhump modulations. For each dataset, the average orbital lightcurve was subtracted and the resulting data folded and binned on superhump phase. Dotted lines show the sine waves fitted to determine phasing.
weight to the lower values in each bin or using only the lower points in a bin, the impact of flickering on these superhump phase binned lightcurves should be reduced. This chosen weighting assigns 8 times more weight to the lowest point in the bin than to the highest.

Figure 3.8: The superhump modulation for the 1993 dataset, calculated using three different types of average when binning the data. The continuous curve is a simple average (the same as in Figure 3.7b), the dotted curve averages the lowest 25 per cent of values in each bin while the dashed curve does a weighted average of the points in each bin.

The curves produced by these different methods are the same except for a flux offset between them (the 1993 curves are shown in Figure 3.8). This suggests that flickering has little effect on the shape of the superhump curves. Since the datasets are extensive and the timescales of the flickering and the superhumps are very different, this is not unexpected.

For a high inclination system a modulation on the superhump period will arise due to the eclipses of the precessing accretion disc changing as the disc orientation changes. This effect will occur in addition to the intrinsic variations in luminosity which are observed in non-eclipsing systems. Figure 3.9 shows the 1993 superhump modulation as calculated for Figure 3.7b and also shows the 1993 superhump modulation calculated excluding points during eclipse. There is little difference between the two curves, which implies that in this system the form of the superhump lightcurve is not affected by the changing eclipse shape.

The broad form of the modulation is consistent for all three sets of observations. The peak-to-peak fractional amplitudes are shown in Table 3.4 and Figure 3.7 and decline
3.7 Average lightcurves

Figure 3.9: The superhump modulation for the 1993 dataset, calculated using three different types of average when binning the data. The continuous curve is a simple average (the same as in Figure 3.7b) while the dotted curve includes only those points out of eclipse.

steady from year to year.

Simpson & Wood (1998) calculate pseudo-lightcurves for superhumps. Assuming the light emitted from the disc is proportional to changes in the total internal energy of the gas in the disc, they present superhump shapes calculated for mass ratios of 0.050, 0.075 and 0.100 (their Figure 5). These curves have significant differences in morphology: the \( q = 0.050 \) curve has a sharp rise and slow decline, the \( q = 0.075 \) curve is reasonably symmetric, and the \( q = 0.100 \) curve has a slow rise and steeper decline. The cleanest and most reliable of the superhump curves, that from 1993, also shows an asymmetric shape, with a slow rise and sharper decline, agreeing best with their highest value of \( q = 0.100 \). We expect \( q \sim 0.31 \) so this is not inconsistent.

3.7 Average lightcurves

Because of the timing of the data (Table 3.1) and the \( \sim 2 \)-day disc precession period, the observed eclipses for each year appear in two rough groups, with superhump phases separated by about 0.5. An obvious approach to analysing the eclipses was therefore to compare average orbital lightcurves for each group. The average mid-eclipse superhump phase, \( \phi_{sh} \), of each group, and the range as indicated by the variance of \( \phi_{sh} \) for each group are shown in Table 3.7, the average orbital curves are shown in Figure 3.10.

The group A curve for 1991 (Figure 3.10a) displays a clear hump at orbital phase around -0.2: the mid-eclipse superhump phase of group A is 0.24, so eclipse should occur
Figure 3.10: Average orbital lightcurves for each year. The data for each year are sorted into two groups, so that all eclipses in each group have mid-eclipse times with a similar superhump phase. The values of $\phi_{\text{sh}}$ for each group are shown in Table 3.7. The third curve in each plot is the difference between the two lightcurves.
3.7 Average lightcurves

Table 3.7: Groups

<table>
<thead>
<tr>
<th></th>
<th></th>
<th></th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>$P_{\text{prec}} = 1.89d$</td>
<td>$P_{\text{prec}} = 1.64d$</td>
<td>$P_{\text{prec}} = 1.90d$</td>
</tr>
<tr>
<td>Group</td>
<td>$\bar{\phi}_{\text{sh}}$</td>
<td>$\sigma_{\phi}$</td>
<td>$\bar{\phi}_{\text{sh}}$</td>
</tr>
<tr>
<td>A</td>
<td>0.24</td>
<td>0.06</td>
<td>0.02</td>
</tr>
<tr>
<td>B</td>
<td>0.71</td>
<td>0.07</td>
<td>0.48</td>
</tr>
</tbody>
</table>

0.25 in orbital phase after a superhump maximum. Group B has mid-eclipse superhump phase 0.71, meaning that the eclipse should occur around 0.3 in orbital phase before the superhump. The post-eclipse flux in group B is higher than in A, although there is no distinct hump.

Group A in the 1993 data displays a hump peaking just before mid-eclipse, while the curve for group B is rather flat out of eclipse. The mid-eclipse superhump phase of group A is 0.02, while group B has superhump phase 0.48. This is again consistent with the position of the superhump. The difference curve (i.e. A-B) seems to show a broad superhump which is (possibly only partially) eclipsed by the secondary. We expect group B to display a hump around orbital phase 0.5. This is not obvious, meaning that the superhump is more prominent when it occurs at orbital phase 0 (group A) than when it occurs at phase 0.5. Schoembs (1986) also noted a similar effect in OY Car. This will be considered further in Section 3.10. This is evidence for an orbital modulation of the intrinsic superhump intensity or visibility, suggested in Section 3.4.3 as a possible cause of the higher frequency photometric signals detected.

The superhump phasing of the 1995 eclipses is almost the same as those observed in 1993, but the smaller extent of the 1995 data and the low fractional amplitude makes identifying the hump difficult without first subtracting the orbital lightcurve (compare Figs 3.7b and 3.7c). However, the flux during eclipse for group A is higher than for group B, consistent with the phasing which suggests that a superhump should occur at mid-eclipse.
Figure 3.11: Eclipse parameters. The O-C mid-eclipse times and the full eclipse widths at half depth. The superhump phase, $\phi_{sh}$, is calculated using the superhump period, $P_{sh}$ determined directly from the corresponding dataset. The continuous, dashed, dot-dashed and dot-dot-dot-dashed curves correspond to predictions from the lightcurve fits $f$, $e$, $a$ and $b$ described in Section 3.8.

3.8 Eclipse parameters

The O-C mid-eclipse times and the eclipse widths are shown in Figure 3.11. The mid-eclipse times were determined both by fitting a parabola to the deepest half of each eclipse and also by finding the centroid. The discrepancy between the two determinations provides an indication of the uncertainty.

As the eccentric disc precesses slowly in an inertial frame, we expect to see the eclipse width and midpoint phase modulated on the apsidal precession period. These quantities will be similarly modulated in superhump phase, since the superhump phase and precession phase of an eclipse are both measures of the relative orientation of disc and secondary star at mid-eclipse. Figure 3.11 shows that eclipse timings for all years exhibit a precession period modulation. The widths also show evidence of a modulation. The limited superhump phase coverage means that conclusions cannot easily be drawn from inspecting the datapoints alone. Such variations in eclipse asymmetry have been observed in other
superhumping systems e.g. OY Car (Schoembs 1986) and Z Cha (Warner & O'Donoghue 1988; Kuulkers et al. 1991). To investigate the disc shape further a simple model was produced which was then fitted to the observed lightcurves.

3.8.1 A simple eccentric disc model

Our simple eccentric disc prescription (shown in Figure 3.12) has a circular inner boundary with radius $r_{\min}$, centred on the white dwarf. The outer boundary is an ellipse of semi-major axis $a_{\max}$, eccentricity $e$, with one focus also centred on the white dwarf. The disc brightness at distance $r(\alpha)$ from the white dwarf at an angle $\alpha$ to the semi-major axis is

$$S(\alpha) \propto \left( \frac{r(\alpha) - r_{\min}}{r_{\max}(\alpha) - r_{\min}} + \frac{r_{\min}}{a_{\max}(1-e) - r_{\min}} \right)^{-n},$$

where $r_{\max}(\alpha)$ is the distance from the white dwarf to the outer disc boundary at orientation $\alpha$. Brightness contours are therefore circular at the inner boundary, smoothly changing to elliptical at the outer boundary. This form for $S(\alpha)$ reduces to $S \propto r^{-n}$ if the disc is circular. Our model is sensible for a tidally distorted disc, since the tidal influence of the secondary star is unimportant at the inner disc, so we expect a more or less circular inner disc.

In an inertial frame, the disc slowly precesses progradely with period $P_{\text{prec}}$. With respect to the corotating frame of the system, this disc then rotates retrogradely with
3.8 Eclipse parameters

period $P_{sh}$. Let the relative orientation of the line of apsides of the disc with the line of centres of the binary when superhump maximum occurs be $\phi_{disc}$ (in phase units). The structure of the disc in our model is therefore described by five parameters: $r_{\text{min}}$, $a_{\text{max}}$, $e$, $n$ and $\phi_{\text{disc}}$. This model was chosen as the simplest way to model an eccentric precessing disc, with as few parameters as possible. A similar model was used by Patterson, Halpern & Shambrook (1993) to model the disc in AM CVn.

3.8.2 Fitting the model

Synthetic lightcurves for eclipses of the model disc were generated using the orbital parameters from Section 3.5. By varying the parameters of the model to minimize the reduced $\chi^2_r$ of the fit, a best fit of the model to the lightcurves was obtained. $\chi^2_r$ is defined as

$$\chi^2_r = \frac{1}{N-M} \sum_{i=1}^{N} \left( \frac{f_i - m_i}{\sigma_i} \right)^2$$

(3.2)

where $N$ and $M$ are the number of points in the lightcurve and the number of free parameters in the model, and $f_i$, $m_i$ and $\sigma_i$ are the observed flux, model flux and the flux error for the $i$th point in the lightcurve. The smoothed superhump was subtracted from the lightcurves before fitting the model in order to remove the intrinsic variation in the disc flux, enabling study of the disc shape. In Section 3.7 we noted that the superhump is more visible for $\phi_{sh} = 0$ at $\phi_{\text{orb}} = 0$, so ideally we should subtract a superhump modulation which takes account of this variation in superhump prominence, but insufficient sampling of the disc precession phase by our data prevented us from doing this.

The downhill simplex method for minimizing multidimensional functions was used (the amoeba routine from Press et al. (1992)). Each orbit (centred on an eclipse) was allowed to have a different total disc flux and a different uneclipsed flux. This prevents the variation of $\sim 10$ per cent in flux from one orbit to the next from interfering with the results, and allows for the possibility of a contribution to the lightcurve from regions never eclipsed by the secondary star. The errors in the normalized (and therefore dimensionless) fluxes, $F$, were estimated as being $\sigma \sqrt{F}$ where $\sigma$ is the square root of the variance of the flux between orbital phase 0.2 and 0.8. This estimate therefore includes the effect of flickering.

3.8.3 Error estimation

Two methods were used to assess the robustness and accuracy of these fits: a Monte-Carlo method to estimate the size of the region in parameter space which has a $\sim 75$ per cent
3.8 Eclipse parameters

The five parameters, \( r_{\text{min}} \), \( a_{\text{max}} \), \( e \), \( n \) and \( \phi_{\text{disc}} \), form a five-dimensional space in which we are trying to locate a minimum value, \( \chi^2_{r,\text{min}} \), of \( \chi^2_r \), corresponding to the best fit solution. Points with higher values of \( \chi^2_r \) are worse fits, and are less likely to represent the 'true' set of parameters. A 5-dimensional volume in parameter space whose surface is defined by \( \chi^2_r \leq \chi^2_{r,\text{min}} + \Delta\chi^2_r \) therefore has a definite probability of containing the 'correct' solution. The range of parameter values in this confidence region tells us how precisely we can measure the parameters with a particular level of confidence. The level of confidence depends on \( \Delta\chi^2_r \) and the properties of the noise. Our job is to determine the value of \( \Delta\chi^2_r \) which gives us a particular confidence in the results for the noise in the observations. This was done as described below.

First, the best-fit parameters for the 1993 lightcurve were obtained. Then \( \chi^2_r \) was calculated for the \( 10^5 \) synthetic curves corresponding to parameters on \( 10 \times 10 \times 10 \times 10 \times 10 \)-element Cartesian grid in parameter space, centred on the best-fit point. Computing speed limited the test to this resolution. This defines the 5-dimensional region in parameter space around the best-fit point defined by \( \chi^2_r \leq \chi^2_{r,\text{min}} + \Delta\chi^2_r \).

The smooth best-fit synthetic lightcurve for the 1993 dataset was then used as a basis for generating 500 fake lightcurves. Since the noise in the observations is complicated by the addition of flickering, a normal distribution for fake noise is not realistic: there will be a disproportionate number of values with large deviations due to flickering. Therefore, to give a more reasonable estimate of errors, 250 lightcurves had normal distribution noise added while 250 had uniform distribution noise added. Each of these 500 lightcurves was then fitted using the fitting routine, and a set of best-fit parameters obtained for each. This technique tells us what spread of best-fit parameters is produced due to the noise. By binning the number density of solutions in parameter space onto the same grid as was used to calculate \( \chi^2_r \) for the 1993 lightcurve, the fraction of solutions within a particular surface constrained by \( \Delta\chi^2_r \) was determined. The result obtained is that 75 per cent of solutions lie within the surface for \( \Delta\chi^2_{r,\text{max}} = 13.3/(N - M) \).

This now enables us to estimate the 75 per cent confidence range for each parameter for each fit. As the fitting algorithm searched for the best fit, the values of each parameter...
and $\chi^2$ were recorded for each step. Once the best fit was found, the maximum deviation of each parameter during convergence from the best fit value was found. Assuming that close to the minimum of $\chi^2$ we have $\chi^2 - \chi^2_{min} \propto (x - x_{min})^2$, where $\chi^2_{min}$ is the minimum value of $\chi^2$, $x$ is the parameter and $x_{min}$ its value at the minimum, we can extrapolate to find the value of $x$ corresponding to $\Delta \chi^2_{x_{max}} = 13.3/(N - M)$. We call the 75 per cent confidence error $\sigma_{75}$.

**Reliability of convergence**

To assess the uniqueness and robustness of each solution the fitting process was carried out 20 times for each model/data combination starting each fit with a different random initial simplex. The solution chosen was that with the lowest value of reduced $\chi^2$. Extreme outlier solutions are rejected and the variance in the parameters for the remaining solutions was used as a measure of the accuracy with which the AMOEBA routine converges to a unique solution, which we call $\sigma_{converge}$. The parameters resulting from all the fits are shown in Table 3.8. $R_L$ is the Eggleton radius of the primary Roche lobe (Eggleton 1983). The convergence of the fit is not precisely unique because the $\chi^2$ surface in parameter space is not perfectly smooth. There is a broad global minimum superimposed with smaller amplitude bumps. Close to the global minimum the gradient of $\chi^2$ is low and so small bumps can lead AMOEBA to settle into a local minimum near the real minimum.

The errors quoted in Table 3.8 for each parameter are whichever is the greater of the confidence region, $\sigma_{75}$, and the convergence estimate, $\sigma_{converge}$.

### 3.8.4 Results from model fitting

Fits using the model as described above will be referred to as fit $f$. The emission extends out to 80–90 per cent of Roche lobe size, while the largest radius of the disc is $a_{max} (1+e) = 0.97 R_L$ for 1995 data. These results suggest that the disc does indeed extend out to the tidal cut-off radius, $r_{tide} \sim 0.9 R_L$ (Paczynski 1977).

The 1995 radius is 10 per cent larger than the 1991 and 1993 radii, but we note that the uncertainty in each value of $a_{max}$ is also about 10 per cent. The 1991 and 1993 observations are in blue CuSO$_4$ filter light, which should come from hotter inner disc regions; the 1995 observations are in R which is expected to weight the outer disc emission more heavily, perhaps causing the inferred disc radius to be largest for the 1995 observations. A simple model black-body disc with a $T \propto R^{-\frac{3}{4}}$ temperature distribution, with $T = 10,000K$
3.8 Eclipse parameters

Table 3.8: Parameters resulting from fitting various models to the lightcurves (see Section 3.8). For the behaviour of $a_{\text{max}}$ and $e$ in each model see Table 3.9. The numbers quoted in brackets correspond to the estimated uncertainty in the last two figures of each parameter value.

<table>
<thead>
<tr>
<th>Fit</th>
<th>Year</th>
<th>$r_{\text{min}}/R_L$</th>
<th>$a_{\text{max}}/R_L$</th>
<th>$e$</th>
<th>$n$</th>
<th>$\phi_{\text{disc}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td>$a_1/R_L$</td>
<td>$a_2/R_L$</td>
<td>$e_0$</td>
<td></td>
<td></td>
</tr>
<tr>
<td>$f$</td>
<td>1991</td>
<td>0.13(13)</td>
<td>0.81(10)</td>
<td>0.035(03)</td>
<td>0.61(81)</td>
<td>0.46(17)</td>
</tr>
<tr>
<td></td>
<td>1993</td>
<td>0.12(09)</td>
<td>0.83(07)</td>
<td>0.106(02)</td>
<td>0.94(65)</td>
<td>0.41(11)</td>
</tr>
<tr>
<td></td>
<td>1995</td>
<td>0.11(13)</td>
<td>0.90(09)</td>
<td>0.082(03)</td>
<td>0.64(71)</td>
<td>0.53(13)</td>
</tr>
<tr>
<td>$e$</td>
<td>1991</td>
<td>0.13(10)</td>
<td>0.82(08)</td>
<td>0.121(06)</td>
<td>0.62(76)</td>
<td>0.35(15)</td>
</tr>
<tr>
<td></td>
<td>1993</td>
<td>0.09(10)</td>
<td>0.81(07)</td>
<td>0.144(03)</td>
<td>0.67(60)</td>
<td>0.41(03)</td>
</tr>
<tr>
<td></td>
<td>1995</td>
<td>0.08(10)</td>
<td>0.88(06)</td>
<td>0.102(04)</td>
<td>0.48(52)</td>
<td>0.55(15)</td>
</tr>
<tr>
<td>$a$</td>
<td>1991</td>
<td>0.12(14)</td>
<td>0.72(12)</td>
<td>0.88(17)</td>
<td>0.029(02)</td>
<td>0.59(84)</td>
</tr>
<tr>
<td></td>
<td>1993</td>
<td>0.11(11)</td>
<td>0.76(08)</td>
<td>0.99(16)</td>
<td>0.099(04)</td>
<td>1.11(73)</td>
</tr>
<tr>
<td></td>
<td>1995</td>
<td>0.10(16)</td>
<td>0.78(11)</td>
<td>0.91(09)</td>
<td>0.068(05)</td>
<td>0.34(96)</td>
</tr>
<tr>
<td>$b$</td>
<td>1991</td>
<td>0.12(15)</td>
<td>0.74(09)</td>
<td>0.89(20)</td>
<td>0.104(08)</td>
<td>0.55(79)</td>
</tr>
<tr>
<td></td>
<td>1993</td>
<td>0.10(09)</td>
<td>0.76(08)</td>
<td>0.98(13)</td>
<td>0.140(05)</td>
<td>0.99(61)</td>
</tr>
<tr>
<td></td>
<td>1995</td>
<td>0.09(17)</td>
<td>0.80(09)</td>
<td>0.93(10)</td>
<td>0.084(06)</td>
<td>0.42(61)</td>
</tr>
</tbody>
</table>

at the inner disc radius $r_1 = 0.1R_L$ and with an outer disc radius $r_2 = R_L$ was used to calculate eclipse profiles in the R and B band. Fitting model $f$ to these eclipses showed no significant difference in outer disc radius between the R and B band fits.

The inner disc boundary and the index $n$ in the flux distribution are poorly constrained. The eccentricity is robustly non-zero; the changing eclipse shape demands a non-axisymmetric disc.

The most interesting result is that all three datasets have $\phi_{\text{disc}}$ around 0.4 to 0.5. This means that in this elliptical disc model the secondary star sweeps past the smallest radius part of the disc at superhump maximum - a result unexpected if tidal stressing of the disc by the gravitational influence of the secondary star is responsible for the superhump light. The implications of this result are discussed later.

The model was adjusted so that the eccentricity varied during the superhump cycle as $e(\phi_{sh}) = e_0 \cos^2 \pi \phi_{sh}$. This will be referred to as fit $e$. This variation in eccentricity follows
that in the simulations of Simpson & Wood (1998) in which the disc varies between being highly eccentric at the superhump maximum to almost circular away from the superhump. The results for $r_{\text{min}}$ and $a_{\text{max}}$ changed very little, with $a_{\text{max}}$ again larger in the red (1995) than the blue (1991 and 1993). $\phi_{\text{disc}}$ was unchanged from fit $f$ within the errors for all three years. The maximum eccentricity, $e_0$, was larger than when $e$ was constant. This is expected since the eccentricity is demanded by the variation in O-C mid-eclipse times, and these O-C times are non-zero at times when $e$ is less than $e_0$.

Next, fits were obtained in which the eccentricity was again constant, but where $a_{\text{max}}$ was allowed to vary between $a_1$ at superhump maximum and $a_2$ half a superhump period later; $a_{\text{max}}(\phi_{\text{sh}}) = a_1 + (a_2 - a_1) \sin^2 \pi \phi_{\text{sh}}$. This will be referred to as fit $a$. This was an attempt to reproduce the observed variations in eclipse width (Figure 3.11). $r_{\text{min}}$, $e$ and $\phi_{\text{disc}}$ were essentially the same as in the first fit, while the values of $a_1$ and $a_2$ implied a variation in $a_{\text{max}}$ of amplitude $20 - 25$ per cent in the blue and $14$ per cent in the red, with the disc being smallest at superhump maximum. At its largest, the disc extends to the edge of the Roche lobe. While these implied variations in disc size are large, they are comparable to the uncertainties in $a_1$ and $a_2$, and so must be treated with caution. They are, however, consistent with the area variations found in SPH simulations of discs during the superhump cycle and provide a simple mechanism through which superhumps in black hole binary systems can be explained (Haswell et al. 2001).

The final variation of the model was to allow both $e$ and $a_{\text{max}}$ to vary as described above. This will be referred to as fit $b$. The eccentricity and $\phi_{\text{disc}}$ were little different from the fits with $e$ varying periodically and $a_{\text{max}}$ constant, while the values of $a_1$ and $a_2$ follow those from the previous fit ($e$ constant and $a_{\text{max}}$ varying).

Table 3.9: Treatment of $a_{\text{max}}$ and $e$ in our four models. $r_{\text{min}}$, $n$ and $\phi_{\text{disc}}$ are constant in all four models.

<table>
<thead>
<tr>
<th>Fit</th>
<th>$a_{\text{max}}$</th>
<th>$e$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$f$</td>
<td>constant</td>
<td>constant</td>
</tr>
<tr>
<td>$e$</td>
<td>constant</td>
<td>$e_0 \cos^2 \pi \phi_{\text{sh}}$</td>
</tr>
<tr>
<td>$a$</td>
<td>$a_1 + (a_2 - a_1) \sin^2 \pi \phi_{\text{sh}}$</td>
<td>constant</td>
</tr>
<tr>
<td>$b$</td>
<td>$a_1 + (a_2 - a_1) \sin^2 \pi \phi_{\text{sh}}$</td>
<td>$e_0 \cos^2 \pi \phi_{\text{sh}}$</td>
</tr>
</tbody>
</table>
3.8 Eclipse parameters

The treatment of $a_{\text{max}}$ and $e$ for each fit is summarized in Table 3.9. Figure 3.13 shows part of the fit to the 1995 dataset using model $a$. The fit (shown as the continuous line) shows the different disc flux and uneclipsed flux allowed by our model in the form of the discontinuities at phase 30.5 and 31.5. The level of flickering out of eclipse can also clearly be seen, and was taken into account in the estimate of the errors as described above.

Figure 3.13: One night of the January 1995 data with the best fitting lightcurve using model $a$ plotted as a continuous line. The discontinuities in the fit illustrate the different disc fluxes and uneclipsed fluxes allowed for each orbit. See Section 3.8.

Formally, the best model is that which achieves the lowest value of reduced $\chi^2$ ($\chi^2_r$). In Figure 3.14 we show the values of $\chi^2_r$ achieved for each model and dataset relative to the lowest. The minimum $\chi^2_r$ achieved was around 0.8 for all datasets and models. This figure shows that the fits $a$ and $b$, i.e. those in which $a_{\text{max}}$ varies on the superhump cycle, produce significantly better fits to the 1993 and 1995 observations. The variation of $e$ during the superhump cycle has little effect on the quality of these fits. There is less significant difference between the $\chi^2$ achieved by the different fits to the 1991 observations, although allowing $a_{\text{max}}$ or $e$ to vary during the superhump cycle produces a better fit than when they are both constant. The significant reduction in $\chi^2_r$ achieved by allowing $a_{\text{max}}$ to vary implies that this model best represents the behaviour of the system.

It is also interesting to compare how well each model predicts the variation in the eclipse width and O-C mid-eclipse times. The predictions of each model are plotted in Figure 3.11. It is periodic variation in these two eclipse characteristics which requires the disc to be eccentric. The models poorly reproduce the O-C variations in the 1991 data. While the phasing of the predicted variation agrees with the observations, the amplitude is
too low. The fits in which $e$ varies during the superhump cycle predict a larger modulation in O-C times, a result of the larger eccentricity in these fits, but the agreement for these fits is still poor. The 1991 lightcurves suffer more from flickering than the 1993 and 1995 data, with many eclipses distorted as a result. This is the most likely explanation for the poor agreement between our model and the 1991 lightcurves. The agreement between the predicted and observed O-Cs is very good for all models for the 1993 observations. The variation in eclipse width is only reasonably modeled by those fits in which $a_{\text{max}}$ varies. The same is true of the 1995 fits.

The result of these comparisons between the different models, both the formal comparison of reduced $\chi^2$ and the more subjective 'chi-by-eye' considerations of the O-C times and eclipse widths is that the models in which $a_{\text{max}}$ varies during the superhump cycle predict the observations better than those in which $a_{\text{max}}$ is constant.

All four models agree on three important points. The values of $a_{\text{max}}$, $a_1$ and $a_2$ show that the disc is large, filling at least about 50 per cent of the Roche lobe area in the orbital plane. The disc is not axisymmetric. From the consistent values of $\phi_{\text{disc}}$ we see that when the superhump reaches maximum light, the light centre of the disc is on the far side of the white dwarf from the donor star.

Figure 3.14: The minimum values of $\chi^2_f$ achieved for the various models fitted (Section 3.8).
3.9 Eclipse mapping

In Section 3.8 the changing eclipse profiles were used to constrain the parameters of a model intensity distribution. An alternative method for investigating the distribution of emission in the orbital plane is the commonly used eclipse mapping technique developed by Horne (1985), described in Section 2.2.1. This technique has been widely used, and O'Donoghue (1990) employed it to locate the source of the strong common superhumps in Z Cha.

In order to study the shape of the precessing disc in V348 Pup, the PRIDA eclipse mapping code of Baptista & Steiner (1991) was modified so that the intensity distribution was fixed in the precessing disc frame rather than the corotating frame of the binary. This is a new approach. Past eclipse mapping and modeling of superhumps in CVs (Warner & O'Donoghue 1988; O'Donoghue 1990) has assumed a brightness distribution corotating with the binary. Each year's data was split into two groups (Section 3.7) but the lightcurves were not folded on orbital phase. This enabled us to obtain two maps for each year, corresponding to the groups given in Table 3.7. Since we expect the intensity distribution to change throughout the superhump cycle, grouping the eclipses as described means that the intensity distribution should be roughly the same for all eclipses in a group, an assumption of the eclipse mapping method. The superhump modulation was subtracted from each lightcurve as in Section 3.8. Normalization of the lightcurves was achieved by using the values for total disc flux and uneclipsed flux for each orbit obtained during the fitting procedure in Section 3.8. The uneclipsed flux was subtracted from each orbit and fluxes were then rescaled to produce an effective disc-only lightcurve. Various other normalization techniques were tested, and the detail of the reconstructed maps was sensitive to these changes. Only the gross structure discussed below was robust. Orbital parameters $q = 0.31$ and $i = 81^\circ$ were used, and the most axisymmetric solution consistent with the data was sought. The minimum values of $\chi^2$ achieved for the maps were in the range 0.6–0.8, suggesting that the errors were slightly over-estimated.

Figure 3.15 shows the two eclipse maps for each year, with the average superhump phase corresponding to each one. The maps are fixed in the precessing disc frame, so the average orientation of the primary Roche lobe is shown on each map.

While we limit the conclusions drawn from these eclipse maps, there are a number of points deserving consideration.

The eclipse maps do not show evidence of a bright spot, but this does not rule out
Figure 3.15: Maximum entropy eclipse maps of the intensity distribution in the precessing frame of the disc. The average orientation of the primary Roche lobe for each group is also shown (in white). See Section 3.9. Dark blue corresponds to no emission, and bright red to maximum emission. The colour scale is the same for both maps from each year.
the possibility that a bright spot is the source of the superhump light, for the following reasons. First, the superhump modulation was subtracted from the lightcurves before performing the eclipse mapping, which should reduce the contribution of the bright spot in the maps if it is the primary source of the superhump light. Also, the maps are fixed in the precessing disc frame, rather than the orbital frame of the system, so the hotspot should be blurred azimuthally in our maps by $\sim 70^\circ$ corresponding to the eclipse width of the system of $\sim 0.2$ in orbital phase. There will be additional azimuthal blurring since the eclipses contributing to each map have a spread of disc orientations at mid-eclipse corresponding to the values of $\sigma_\phi$ in Table 3.7. Azimuthal structure in the maps is also suppressed by looking for the maximally axisymmetric solution.

![Figure 3.16: Azimuthally averaged flux in maximum entropy eclipse maps. Quantity plotted, $rF_r$, is proportional to flux in annulus at radius $r$. See Section 3.9.](image)

The most consistent result revealed by these eclipse maps is that the emission at superhump phase 0.5 is less centrally concentrated than at superhump phase 0. This is illustrated in the second and third rows of Figure 3.15 which show the maps for the two 1993 and 1995 groups of eclipses. Figures 3.16a to 3.16c show the azimuthally averaged brightness distribution of each map. They show the flux at radius $r$, $F_r$, multiplied by $r$; this quantity is proportional to the total flux in an annulus at radius $r$. Figures 3.16b and 3.16c show how the disc extends further out at superhump phase 0.5 than at phase 0, while the curves in Figure 3.16a are both nearly the same, as expected since both of these curves show the situation roughly half way between superhump phase 0 and 0.5. This result agrees with the results of our fits of model $a$ in which the disc size was allowed to change. These fits showed that the size of the emission region is larger at superhump
phase 0.5 than at phase 0 (superhump maximum). The maps are asymmetric, but due to the sensitivity described above, no conclusions are drawn from the detailed structure.

3.10 Discussion

The phase of the superhump relative to the conjunction of the line of centres of the system and the semi-major axis of the disc should make it possible to determine whether the bright spot model or the tidal heating model better explains the source of the superhump. The simplest tidal model predicts that the superhump light should peak when (or slightly after) the largest radius part of the disc coincides with the line of centres. This is because the tidal interaction is strongly dependent on distance from the secondary, and so will be most significant in regions where the disc extends out close to the L\textsubscript{1} point. However, if the bright spot model is to be believed, then the superhump light source will be modulated as the stream-disc impact geometry changes during the precession cycle. In the model of Vogt (1982) the hotspot is brightest when the accreting material has the furthest to fall, i.e. the superhump maximum should occur when the stream impacts on the disc at its smallest radius. Simulations of superhumps present a more complicated picture, e.g. Murray (1998), which shows an extended superhump light source due to viscous dissipation as the disc is tidally stressed. In that model, the superhump emission came from the outer edge of the side of disc flowing away from the donor.

The mid-eclipse times shown in the top row of panels in Figure 3.11 show the eclipses to be earliest around superhump phase 0.75 in all cases. Assuming that the centre of light of the eccentric disc is offset from the white dwarf in the direction of the largest radius, we can deduce the disc orientation during these eclipses to be as shown in Figure 3.17a. A quarter of a superhump period later, the orientation of the disc has barely changed, the secondary will be lined up with the smallest radius part of the disc and the superhump phase will be 0.0 (Figure 3.17b). Therefore superhump maximum occurs when the secondary star is lined up with the smallest radius part of the disc. The values of $\phi_{\text{disc}}$ in Table 3.8 agree with this deduction. This phasing is consistent with the bright spot model for the superhump emission but is inconsistent with the simple tidal heating model.

In Rolfe, Haswell & Patterson (2001) a more realistic version of the bright spot model was developed, in which the bright spot is modulated by the relative kinetic energy of the stream and disc flows at the impact point. This takes into account both the depth of the hotspot in the white dwarf potential, and the angle at which the stream impacts the disc.
3.10 Discussion

Figure 3.17: The first row of Figure 3.11 shows the eclipses to be earliest at superhump phase $\phi_{sh} \sim 0.75$, which implies orientation (a), from which we deduce the relative phasing of disc and secondary star at superhump maximum ($\phi_{sh} = 0$) shown in (b).

A detailed discussion of this model is given in Section 5.5.1. Applying this model (model (ii) from Section 5.5.1) to the various fits from Section 3.8.4 enables a prediction of the superhump lightcurves to be made, assuming the hotspot is the superhump light source. A donor mass of $M_{\text{donor}} = 0.20M_\odot$ was assumed, using the empirical relation,

$$\frac{M_{\text{donor}}}{M_\odot} = 0.126 \left( \frac{P}{1 \text{ hour}} \right) - 0.11$$  \hspace{1cm} (3.3)

(Equation 9 from Smith & Dhillon (1998)). Using $q = 0.31$ this gives white dwarf mass $M_{\text{wd}} = 0.64M_\odot$, needed to calculate the stream velocity for the bright spot model. The predicted bright spot curves for all four model fits are shown in Figure 3.18 along with the observed curves. The predicted curves were scaled and shifted in flux to match the observations. For those fits with the disc size changing which had the disc extending outside the primary Roche lobe when the disc was at its maximum size, the maximum semi-major axis was reduced to keep the disc just within the Eggleton radius. These changes in $a_2$ are within the estimated uncertainties in $a_2$ and so are acceptable.

For 1991, the predicted bright spot variations for fits $e$ and $b$ are too narrow and peak too early to match the observed superhump. The curves for fits $f$ and $a$ peak closer to superhump phase zero and have wider peaks, better matching the observed curves. For
1993 the fit $f$ and $a$ curves are broadest again, in better agreement with the observations than fits $e$ and $b$. Fit $a$ agrees better in phase than fit $f$. For 1995 none of the curves provides a good match to the shape of the observed curve, probably a result of the limited extent of the 1995 dataset compromising the observed modulation. The conclusion is that the model superhump fits the observations best for fits $f$ and $a$.

The fact that no clear orbital hump is seen in the lightcurves tells us that the bright spot contributes a small fraction of the total flux. To obtain the largest possible contribution to the lightcurve from hotspot brightness variations while keeping the average hotspot brightness as low as possible, the fractional amplitude$^2$ of the hotspot brightness should be maximized. The ratio of the fractional amplitude of the model $a$ hotspot variation to that of model $f$ is 2.8, 1.4 and 0.9 for the 1991, 1993 and 1995 disc fits respectively. Therefore, for the two most extensive datasets analysed here (1991 and 1993), model $a$ produces a higher amplitude hotspot brightness variation than model $f$ for the same phased-averaged hotspot flux. Hence we should conclude that if the bright spot is the source of the superhumps in V348 Pup, model $a$, in which the disc area changes on the superhump cycle, best reproduces the superhump modulation.

In Section 3.7, it was noted that the superhump is strongest when it occurs around orbital phase 0. This is easily explained if the major contribution to superhump light is the bright spot: the bright spot is most visible when it is on the nearside of the accretion disc.

Schoembs (1986) observed late superhumps in the eclipsing SU UMa dwarf nova OY Car. When a superhump was coincident with a pre-eclipse orbital hump, the combined amplitude was greater than that predicted for a linear superposition of the individual amplitudes i.e. OY Car's late superhumps were strongest around orbital phase 0. van der Woerd et al. (1988) studied the dwarf nova VW Hyi, concluding that there was no correlation between the orbital phase and amplitude of late superhumps, but the precession phase coverage of their observations was insufficient to detect any such correlation.

Schoembs (1986) followed OY Car from early in a superoutburst almost until the return to quiescence, observing the $\sim 180^\circ$ phase change from common superhumps around the height of the outburst to late superhumps during the decline of the superoutburst. Patterson et al. (1995) observed the same change in superhump phase late in a superoutburst of V1159 Ori. Hessman et al. (1992) studied OY Car at the end of a superoutburst. By looking at the varying hotspot eclipse ingress times, and considering the trajectory of the

$^2$Here fractional amplitude is defined as the ratio (maximum flux - minimum flux)/average flux.
Figure 3.18: The predicted superhump lightcurves for each fitted disc model assuming the bright spot is the superhump light source. See Section 3.10.
accretion stream, they concluded that the disc was eccentric. Rolfe, Haswell & Patterson (2001) employed a similar analysis to trace the disc during the late superhump era in the SU UMa system IY UMa (described in Chapter 5). The orientation of the disc at late superhump maximum in both IY UMa and OY Car was very similar to that found in V348 Pup.

In the SPH simulations of Murray (1996) and Murray (1998) pseudo-lightcurves are produced by assuming the heat produced by viscous dissipation to be radiated away where it is generated. Murray (1996) and Murray (1998) reveal an extended superhump light source in the outer disc, while Murray (1996) also reveals an additional superhump modulation which arises from the impact of the accretion stream with the edge of the disc occurring at a varying depth in the primary Roche potential. This additional weaker superhump modulation is approximately 180° out of phase with the modulation due to tidal stressing, another similarity between late superhumps in dwarf novae, the persistent superhumps in V348 Pup and the bright spot model. A possible link between late superhumps and the bright spot model was originally suggested by Osaki (1985) and Whitehurst (1988b).

There are many other studies of the disc structure in SU UMa stars during superoutburst. Krzeminski & Vogt (1985) studied OY Car during a superoutburst and through variations in the O-C eclipse timings deduced the presence of an eccentric disc with phasing similar to that in V348 Pup. Vogt (1982) and Honey et al. (1988) found evidence for an eccentric precessing disc in Z Cha from the radial velocity variations of various absorption and emission lines respectively. The very prominent common superhump in Z Cha made it possible for Warner & O'Donoghue (1988) to study the location of the superhump light source by modeling eclipse profiles for emission distributions and comparing these with observed profiles. They found strong departures from axisymmetry in the superhump surface brightness. O'Donoghue (1990) employed a modified eclipse mapping technique to Z Cha lightcurves and found the common superhump light coming from three bright regions of the disc rim, located near the L1 point and the leading and trailing edges of the disc, concluding that the superhumps are tidal in origin, and that a highly eccentric disc with a smooth brightness distribution is not necessary to explain superhump behaviour. However, this analysis did not account for possible thickening of the accretion disc during outburst, and it is possible that re-analysis taking vertical disc structure into account may change the result. One anomalous eclipse did confine the superhump light source in Z Cha
3.11 Are the persistent superhumps in V348 Pup late?

to the region of the quiescent bright spot.

The maximum entropy eclipse maps tell us that the azimuthally averaged radial extent of the emission is lowest at superhump maximum, shown in Figure 3.16. If this change in extent of the emission region is interpreted as a result of a changing disc size, then the smaller disc radius at superhump maximum is consistent with the bright spot model for the superhump light source. This changing disc area is also supported by the results of the model fitting, and the superhump modulation prediction, in which the best fits were achieved using models a and e, where the size of the disc was modulated on the superhump period. SPH simulations of eccentric discs by Murray also reveal a superhump-phase modulated disc area (Haswell et al. 2001), with the disc area and viscous dissipation maxima coinciding. As discussed, and as found in Murray (1996), the viscous dissipation maximum is antiphase with superhumps originating from the bright spot. This implies that the disc area is smallest in the simulations around bright spot maximum, reinforcing the bright spot interpretation of the persistent superhumps in V348 Pup.

3.11 Are the persistent superhumps in V348 Pup late?

This study of persistent superhumps in V348 Pup has shown that there is an eccentric precessing accretion disc, and that the area of the disc is probably changing through the superhump cycle. The orientation of the disc at superhump maximum, the model superhump variation, the dependence of superhump visibility on orbital phase and the phasing of the disc area changes all point to a link between between the persistent superhump light source in V348 Pup and late superhumps in SU UMa systems. It seems almost certain that late superhumps arise from varying dissipation of kinetic energy at the stream-disc impact (Rolfe, Haswell & Patterson 2001; Section 5.5.1). The result of this work is that the same mechanism may be responsible for persistent superhumps in V348 Pup. The calculation of synthetic eclipse lightcurves from simulations such as those of Murray (1998) is important to determine whether the complicated viscous dissipation patterns in tidally distorted discs could conspire to imitate a superhump from the hotspot in observations like those analysed here.
Chapter 4

The SW Sex phenomenon in V348 Pup

4.1 Introduction

The previous chapter presented a photometric study of the accretion disc in V348 Pup. In this chapter, a high resolution spectroscopic study is presented, revealing detailed information about the accretion flow. The only previous spectroscopic study of V348 Pup was discussed in Tuohy et al. (1990). This study was at a lower resolution than that presented here, but revealed evidence suggesting that the emission lines in V348 Pup may not be dominated by disc emission. We will find that V348 Pup shares many of the features of the so-called SW Sex group of CVs, and consider in some detail the most promising model for this behaviour.

The SW Sex type stars are a group of disc-accreting nova-like CVs whose behaviour does not appear to fit in with the standard model of an accretion disc fed by a stream from the donor star via the stream-disc impact (hotspot). Thorstensen et al. (1991) first noticed the phenomenon in SW Sex, DW UMa, V1315 Aql and PX And. Since then, many more systems have been suggested as SW Sex stars. Smith, Dhillon & Marsh (2001) provides a detailed list of more recent additions: BH Lyn (Dhillon et al. 1992), BP Lyn (Still 1996), WX Ari (Beuermann et al. 1992), V795 Her (Casares et al. 1996), LX Ser (Hoard 1998; Young, Schneider & Shectman 1981), V1776 Cyg (Garnavich et al. 1990; Hoard 1998), UU Aqr (Hoard et al. 1998b), LS Peg (Szkody et al. 1997), V442 Oph (Hoard 1998), BT Mon (Smith, Dhillon & Marsh 1998) and DW UMa (Smith, Dhillon & Marsh 2001). Most recently, Hynes et al. (2001) detected several SW Sex type features in the X-ray binary XTE J2123-058, where the compact object is a neutron star. Recent
reviews of the phenomenon are given in Horne (1999) and Hellier (2000), each of which promotes a different model. Smith, Dhillon & Marsh (2001) provides another valuable review, presenting a detailed study of SW Sex and DW UMa and considering several different models. The anomalous features of the SW Sex stars, as described in Horne (1999), are outlined below.

- Broad single-peaked emission lines, rather than the double-peaked lines expected for emission from the disc of a high-inclination system. The radial velocity of the peak lags behind that expected for the white dwarf. In Doppler maps, the strongest emission region (corresponding to the single peak) is seen at low velocity in the lower-left quadrant (negative $V_x$ and $V_y$).

- Emission line eclipses are much shallower than the continuum eclipse.

- Absorption features are seen around phase 0.5.

- Continuum eclipses are V-shaped, implying a cooler inner disc than in steady-state models.

In addition, Smith, Dhillon & Marsh (2001) also consider the presence of the high excitation He II 4686 Å line and that the system be a nova-like with period 3–4 hours to be defining characteristics. Not all SW Sex systems show all of these features, e.g. WX Ari which lacks the absorption features, while BT Mon has an orbital period of 8 hours, and XTE J2123-058 is not even a CV! We discuss these features throughout this chapter, and consider the various models for the SW Sex phenomenon in Section 4.10.

4.2 Observations

The high resolution observations presented in this chapter were obtained by Dr. C. A. Haswell and A. Shambrook in January 1995 on the 4m Blanco telescope at the Cerro Tololo Interamerican Observatory (CTIO) using the 4m R-C spectrograph. This time-resolved spectroscopy covers spectral range 6300 to 6900 Å, covering both the Hα and He I 6678 Å lines. The observations were simultaneous with the January 1995 R-band photometry presented in the previous chapter. The simultaneous photometry made it possible to correct the reduced spectra for slit losses, as described below. 441 spectra were collected with integration times varying between 120 s and 480 s. We also present some lower resolution spectroscopy of V348 Pup taken by Dr. T. Abbott during technical time.
4.3 Data Reduction

Table 4.1: The Observations. $\Delta \lambda$ is the FWHM resolution of the spectra measured from arc lines. $N$ is the number of spectra.

<table>
<thead>
<tr>
<th>Date</th>
<th>Tel.</th>
<th>Orbits</th>
<th>N</th>
<th>Exp. time (s)</th>
<th>$\lambda$ range (Å)</th>
<th>$\Delta \lambda$ (Å)</th>
<th>Observer(s)</th>
</tr>
</thead>
<tbody>
<tr>
<td>3rd Jan 95</td>
<td>CTIO</td>
<td>1.9</td>
<td>112</td>
<td>120</td>
<td>6300–6900</td>
<td>1.2</td>
<td>Haswell, Shambrook</td>
</tr>
<tr>
<td>4th Jan 95</td>
<td>CTIO</td>
<td>0.9</td>
<td>56</td>
<td>120</td>
<td>6300–6900</td>
<td>1.2</td>
<td>Haswell, Shambrook</td>
</tr>
<tr>
<td>5th Jan 95</td>
<td>CTIO</td>
<td>3.0</td>
<td>167</td>
<td>120</td>
<td>6300–6900</td>
<td>1.2</td>
<td>Haswell, Shambrook</td>
</tr>
<tr>
<td>6th Jan 95</td>
<td>CTIO</td>
<td>2.9</td>
<td>106</td>
<td>87 @ 120</td>
<td>6300–6900</td>
<td>1.2</td>
<td>Haswell, Shambrook</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>19 @ 480</td>
<td></td>
<td></td>
</tr>
<tr>
<td>6th Mar 99</td>
<td>NOT</td>
<td>2.1</td>
<td>18</td>
<td>2 @ 600</td>
<td>3180–5550</td>
<td>8.</td>
<td>Abbott</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>16 @ 400</td>
<td></td>
<td></td>
</tr>
<tr>
<td>7th Mar 99</td>
<td>NOT</td>
<td>2</td>
<td>400</td>
<td></td>
<td>3180–5550</td>
<td>8.</td>
<td>Abbott</td>
</tr>
</tbody>
</table>

using the ALFOSC spectrograph on the 2.5m Nordic Optical Telescope. The observations are summarized in Table 4.1.

4.3 Data Reduction

The NOT dataset was reduced using IRAF to apply the standard techniques described in Chapter 2. Flux standard exposures were used to correct the spectra for atmospheric extinction and instrumental response, but no correction for slit losses could be made. The weather conditions were poor leading to considerable slit losses (seen in the lightcurve of the spectra). Because of the poor seeing, a 1.3" slit was used producing the low 8Å resolution. Each spectrum was therefore normalized by fitting the continuum with a 4th order polynomial and dividing by this. Only these normalized spectra are discussed.

The CTIO detector was a 2×2 array of 4 CCDs read using the quad-readout mode. By using the ARED IRAF package, standard techniques were applied for the CCD reductions. These were long slit observations, and distortions meant that the dispersion solution for the spectra varied slightly as a function of position, leading to a slightly bowed appearance of the thorium-argon arc lamp lines. To correct for this effect, the lines in each arc exposure were identified and a dispersion solution was obtained as a function of spatial direction.
4.3 Data Reduction

Groups of ten lines were averaged in the spatial direction for each dispersion solution. By fitting the resulting solutions with 6th order Chebyshev polynomials in both directions, a geometrical correction was applied to all object images which transformed them so that all rows in the image had the same dispersion solution. Note that such transformation solutions were obtained separately for each arc lamp spectrum, and the transformations applied to each object image were obtained from arcs taken at a similar time. By also applying this correction to the arc lamp spectra, it was possible to determine how successful this procedure was - the arc lamp lines were perfectly straight. Applying these corrections also had the effect of wavelength calibrating the images before the spectra were extracted. Standard procedures were then followed to trace and extract the spectra, and correct for atmospheric extinction and instrumental response.

The flux in each spectrum suffered from slit losses. These slit losses occur when the proportion of flux from the star reaching the detector is not constant, as a result of seeing variations, focus variations and imperfect pointing of the telescope. It was possible to correct for this by using the simultaneous R band photometry. Software was written in IDL to carry out this procedure. The method is described below.

The simultaneous photometry provides us with $R_{phot}$ - the R band magnitude of V348 Pup throughout the observations. Since the integrations for the photometry will not coincide exactly with those for the spectroscopy, the lightcurve is averaged over the integration for each spectrum to determine an appropriate $R_{phot}$ for that spectrum. In order to use this to correct the spectra, we need to calculate an equivalent R magnitude for each spectrum. This value, $R_{spec}$, may be calculated using the standard definition which uses the ratio of the total flux in the R band for the spectrum to the equivalent R band flux in Vega.

$$R_{spec} = -2.5 \log_{10} \frac{\text{Flux}_{spec}}{\text{Flux}_{Vega}}$$

where we have

$$\frac{\text{Flux}_{spec}}{\text{Flux}_{Vega}} = \frac{\int_R \tau_{\lambda} f_{\lambda, spec} d\lambda}{\int_R \tau_{\lambda} f_{\lambda, Vega} d\lambda}.$$

$f_{\lambda} d\lambda$ is the flux between wavelength $\lambda$ and $\lambda + d\lambda$, while $\tau_{\lambda}$ is the transmission in the R band as a function of wavelength $\lambda$. To correct the spectrum for slit losses we need to multiply the spectral flux by a constant factor $\alpha$ such that the corrected spectrum has an R magnitude which is equal to $R_{phot}$. This simply requires us to use

$$\alpha = 10^{0.4 \Delta R}$$

where $\Delta R = R_{spec} - R_{phot}$. 
The spectral range of the R band is 5500 Å to 8700 Å whereas the spectra cover 6300 Å to 6800 Å. It is therefore necessary to extrapolate the continuum to cover the full range of the band. A straight line provides a good fit to the continuum in the flux calibrated spectra and extrapolating this proves an effective way of calculating $R_{\text{spectrum}}$. Figure 4.3 shows an example extrapolated spectrum multiplied by the R band transmission function, $\tau_\lambda$. There were a few short periods of time where the simultaneous photometry began after the spectroscopy or finished early. For these instances $R_{\text{phot}}$ was calculated from the appropriate phase of the average orbital light curve for that night.

4.4 Average Spectra

Figure 4.2 shows average spectra for each night of the CTIO observations while Figure 4.3 shows the average NOT spectrum. Note that the NOT spectrum is normalized, while the CTIO data is correctly flux calibrated. These average spectra are uncorrected for orbital motion.

The CTIO spectra show broad Hα and He I 6678Å emission lines, while the NOT spectrum also shows broad Balmer and He I emission, along with the high excitation He II 4686Å and C III/N III Bowen blend around 4640Å. Weaker He II emission lines at 4200Å, 4542Å and 5412Å are also visible in the NOT spectrum, along with C II 4267Å emission, and possibly Ca II K absorption at 3934Å. The Balmer, Bowen and He II lines are all
Figure 4.2: Average CTIO spectra of V348 Pup.
single-peaked, while the profile of \textsc{He}I in the CTIO data has a flat top, and looks almost double-peaked on the 4th January. The NOT spectrum looks very much like the Tuohy et al. (1990) spectrum of V348 Pup, except that in the more recent NOT observations the \textsc{He}II emission is stronger. As described in Section 1.7, disc emission from a high inclination system should produce double-peaked emission line profiles due to the Doppler shift of emission from the approaching and receding sides of the disc. From the previous chapter we know that V348 Pup is a high-inclination disc-accreting system. These single-peaked emission lines in V348 Pup are very similar in behaviour to the SW Sex systems, both in the shape and widths of the lines, and also in the relative strengths of the lines.

Table 4.2 shows the equivalent width and FWHM of the most prominent lines. Each FWHM was measured by fitting the steep blue and red sides of each line between 0.3 and 0.7 times the maximum flux with straight lines and interpolating to find the width at half maximum flux. The equivalent width was calculated by integrating the area under the continuum-normalized line profile. The blended C\textsc{III}/N\textsc{III} and \textsc{He}II features were
measured by fitting a pair of gaussians simultaneously to the two features and using the FWHM and area from the fitted profiles. The FWHMs are all very broad, between \(~1350\) and \(1860\) km s\(^{-1}\), compared to \(~1000\) to \(1500\) km s\(^{-1}\) in DW UMa (Smith, Dhillon & Marsh 2001) and \(~1000\) to \(1600\) km s\(^{-1}\) in BH Lyn (Dhillon et al. 1992). The variation of equivalent width from line to line is much like that in the SW Sex systems, although the strength of the high excitation features compared to the Balmer lines is stronger in V348 Pup than in the SW Sex systems apart from BT Mon (Smith, Dhillon & Marsh 1998). C\(\text{II}\) emission is seen in many of the SW Sex systems, e.g. PX And (Still, Dhillon & Jones 1995) and SW Sex (Dhillon, Marsh & Jones 1997).

Table 4.2: Widths of prominent lines in the average spectra V348 Pup.

<table>
<thead>
<tr>
<th>Line</th>
<th>Telescope</th>
<th>FWHM km s(^{-1})</th>
<th>EW Å</th>
</tr>
</thead>
<tbody>
<tr>
<td>H(\alpha)</td>
<td>CTIO</td>
<td>1460±100</td>
<td>57.8±0.1</td>
</tr>
<tr>
<td>H(\beta)</td>
<td>NOT</td>
<td>1540±70</td>
<td>31.2±2</td>
</tr>
<tr>
<td>H(\gamma)</td>
<td>NOT</td>
<td>1630±140</td>
<td>30.0±2</td>
</tr>
<tr>
<td>H(\delta)</td>
<td>NOT</td>
<td>1860±200</td>
<td>23.6±2</td>
</tr>
<tr>
<td>H(\epsilon)</td>
<td>NOT</td>
<td>1780±170</td>
<td>15.7±2</td>
</tr>
<tr>
<td>He I 6678</td>
<td>CTIO</td>
<td>1350±170</td>
<td>6.5±0.1</td>
</tr>
<tr>
<td>He I 4922</td>
<td>NOT</td>
<td>1630±460</td>
<td>3.8±1</td>
</tr>
<tr>
<td>He II</td>
<td>NOT</td>
<td>1350±50</td>
<td>36.0±2</td>
</tr>
<tr>
<td>Bowen</td>
<td>NOT</td>
<td>1670±150</td>
<td>13.9±2</td>
</tr>
</tbody>
</table>

There is some detailed structure around the peaks of the H\(\alpha\) lines, with a clear narrow absorption dip visible on the blue side of H\(\alpha\) seen only on the 5th January 1995. There is a strong absorption dip on the red side of He I 4471Å in Figure 4.3. By fitting a gaussian to the He I emission, then subtracting the fit and fitting another gaussian to the absorption, the width and wavelength of the absorption were measured. Figure 4.4 shows the fits (continuous curve) plotted over the average profile. This gives a wavelength for the absorption of 4486.9±1.2Å, with a FWHM of 6±3Å and an equivalent width of 1.4±0.9Å. This is red-shifted by 1035 km s\(^{-1}\) from the rest wavelength of the He I line. A very similar
red-shifted absorption was seen in the same line in BH Lyn (Hoard et al. 1998a) and in He\textsc{i} 5876Å in PX And (Thorstensen et al. 1991) while very strong He\textsc{i} absorption was seen in V795 Her (Casares et al. 1996).

4.5 Lightcurves

Figure 4.5 shows the lightcurves of the continuum and the integrated flux under the continuum-subtracted emission lines profiles for the CTIO data.

The deep eclipses are clearly seen in the continuum, while the H\textalpha\ eclipses are weaker, with evidence of further weakening of H\textalpha\ around phase 0.5. There appears to be similar behaviour in He\textsc{i} 6678Å, but in both emission line lightcurves stochastic variations and noise make it difficult to study the curves in detail. Therefore the three lightcurves were folded onto a single orbit and binned into 100 phase bins, producing the average orbital curves shown in Figure 4.6.

The continuum eclipse has a depth of about 75%, and is V-shaped as in many SW Sex systems, e.g. DW UMa (Smith, Dhillon & Marsh 2001) and PX And (Thorstensen et al. 1991). The eclipses in the emission lines are shallower and not V-shaped. The H\textalpha\ flux is strongest before eclipse around phase 0.8, and faintest around phases 0.5–0.6. He\textsc{i} displays the same variations in flux, but the faint region around phase 0.5 is almost as
Figure 4.5: Lightcurves of the continuum and the H\(\alpha\) and HeI 6678Å emission lines.
Figure 4.6: Phase folded and binned lightcurves of the continuum and the Hα and He I 6678Å emission lines.

faint as mid-eclipse. Many of the SW Sex stars show such contrast between the continuum and emission line lightcurves and eclipses, with the lightcurves of BT Mon (Smith, Dhillon & Marsh 1998) looking very much like these V348 Pup curves.

4.6 Line profile variations

Figures 4.7 and 4.8 show the line profile variations throughout the orbit, produced by averaging the spectra in 10 orbital phase bins. The data from 5th January 1995 were not included in creating Figure 4.7 because of the transient blue-shifted absorption feature seen on that night, but we note that apart from that feature, the behaviour of the 5th January profiles was exactly the same as described below.
Figure 4.7: Phase binned CTIO spectra showing the evolution of the line profile.
4.6 Line profile variations

Looking at Figure 4.7, we see significant changes in the line profile throughout the orbit. While Hα remains single-peaked at all phases, between about phase 0.3 and 0.7 a strong broad reduction in the red-shifted emission occurs, being strongest around phase 0.4–0.6. There also another faint absorption feature around phase 0.9–0.1 close to the rest wavelength of Hα. An average of the spectra in this range reveals this feature to be red-shifted by about 80 km s\(^{-1}\). He\(\text{l}\) 6678 Å, much weaker than Hα, is single-peaked in phase range 0.0–0.3 and 0.7–1.0, but becomes more complicated, appearing double-peaked around phase 0.3–0.7. The lines are faintest around phases 0.4–0.6, as seen in the lightcurves.

The lower S/N and spectral resolution line profiles in Figure 4.8 show the same behaviour in the Balmer and He\(\text{l}\) lines as seen in the CTIO data, with the broad red weakening around phase 0.5 in the Balmer lines, and double peaks in He\(\text{l}\) at that phase. The faint blue-shifted absorption is not seen, but this could be due to lower resolution of the NOT data. He\(\text{II}\) appears to show none of the variations of the lower excitation lines. The sharp red-shifted absorption in He\(\text{l}\) 4471 Å is significantly stronger around phase 0.3–0.6 than earlier in the orbit, and appears weakest around phase 0.8–0.9. There are no systematic orbital variations apparent in the Ca\(\text{II}\) absorption.

These variations in emission line profiles are almost identical to those seen in DW UMa and very similar to those in SW Sex reported in Smith, Dhillon & Marsh (2001). Such emission line variations are seen in the other SW Sex stars too, e.g. V795 Her (Casares et al. 1996). Hoard (1998) also presents line profile variations in DW UMa, similar to those in Smith, Dhillon & Marsh (2001) and these V348 Pup data, although the 'absorption' features are stronger in the Hoard (1998) spectra than in the Smith, Dhillon & Marsh (2001) DW UMa data, showing that the detailed behaviour varies over time.

The maximum strength of the He\(\text{l}\) 4471 Å absorption feature occurs around phase 0.5, in broad agreement with the SW Sex systems, e.g. PX And (Thorstensen et al. 1991). It should be stressed that only those features where the flux goes below the continuum can definitely be identified as absorption lines. Dips in the emission line profiles could result from genuine absorption lines, or from occultations of, or intrinsic changes in, the emission from the accretion flow.
4.6 Line profile variations

Figure 4.8: Phase binned NOT spectra showing the evolution of the line profile.
4.7 Radial velocities

Figure 4.9 shows the heliocentric velocities of the prominent emission lines, measured using the same fitting methods used to measure the FWHMs in Section 4.4. Red-shifted velocities are positive. Black dots show the measurements, with error bars in grey and sinusoidal fits of the form \( V = \gamma - V_1 \sin 2\pi(\phi - \phi_0) \). Dashed purple lines show the value of \( \gamma \). The values of \( \gamma \), \( V_1 \) and \( \phi_0 \) for each line are shown Table 4.3. The fitting routine was unable to find a satisfactory sinusoidal fit to the HeI 6678Å emission line measured from the CTIO data. This is because the low S/N and complicated changes in the HeI 6678Å line profile from single to double-peaked prevent consistent measurement of the
radial velocity throughout the orbit.

The Hα curve is clearly the most reliable radial velocity curve, coming from 441 spectra of high resolution and high signal-to-noise. There is a very clear roughly-sinusoidal variation in these velocities, with a sinusoid providing a good fit except at the two phases of maximum excursion - phases 0.4–0.5 and 0.9–1.1, where the measured velocities are too high for the fitted curve. The line profiles are most complicated around these two phases. The red weakening of Hα is strongest around phase 0.5, and the eclipse around phase -0.1–0.1 will affect the measurements, since we cannot expect the accretion flow to be eclipsed equally at all velocities throughout eclipse. The maximum velocity of the Hα fit is at phase 0.92, which is inconsistent with the white dwarf velocity, which should be maximum at phase 0.75. Such emission line velocity phase lags of 0.1–0.2 behind the white dwarf orbit are one of the defining characteristics of SW Sex stars Horne (1999). In addition, the projected amplitude of the white dwarf velocity should be about 100 km s⁻¹¹, while the measured value is 337 km s⁻¹. Tuohy et al. (1990) find the same Hα velocity amplitude in V348 Pup, with the same phasing. However they find \( \gamma = 99 \pm 13 \) km s⁻¹ while the Hα value from this study is \( \gamma = 44.9 \pm 0.3 \) km s⁻¹. However Tuohy et al. (1990) do find \( \gamma = 45 \pm 12 \) km s⁻¹ for He II. The measurements from the NOT data give \( \gamma \) between 166 and 272 km s⁻¹. This discrepancy could be explained simply due to the low resolution of those data (FWHM ~ 480 km s⁻¹). The Hα measurement is our most reliable estimate of \( \gamma \), although the discussion above reveals how it cannot be taken as a determination of the systemic velocity of V348 Pup. The consistent results from the other radial velocity measurements listed in Table 4.3 are that the emission lines have too large a velocity amplitude and are at the wrong phase to correspond to emission from the white dwarf or boundary layer, or any other known component of disc-accreting CVs.

4.8 Trailed spectra

Figure 4.10 shows trailed spectra of the CTIO Hα and Hβ emission lines, produced by binning each night's data into 60 phase bins per orbit and 120 velocity bins between -1800 and 1800 km s⁻¹. These plots present a more compact way of studying the line profile variations, enabling us to view the full evolution of the line profiles without folding the data onto a single orbit.

The Hα trail shows the broad single-peaked S-wave seen in Section 4.6, clearly following

\(^1\)using the \( M_{wd} = 0.64 M_\odot \) and \( q = 0.31 \) estimated for V348 Pup in the previous chapter (Section 3.10)
Table 4.3: Radial velocity fits to prominent emission lines in V348 Pup, where $V = \gamma - V_1 \sin 2\pi(\phi - \phi_0)$.

<table>
<thead>
<tr>
<th>Line</th>
<th>$\gamma$ km$^{-1}$</th>
<th>$V_1$ km$^{-1}$</th>
<th>$\phi_0$</th>
</tr>
</thead>
<tbody>
<tr>
<td>H$\alpha$</td>
<td>44.9±0.3</td>
<td>336.9±0.5</td>
<td>0.1784±0.0002</td>
</tr>
<tr>
<td>H$\beta$</td>
<td>198.0±1.8</td>
<td>204.0±2.6</td>
<td>0.121±0.002</td>
</tr>
<tr>
<td>H$\gamma$</td>
<td>230.2±2.3</td>
<td>131.0±3.3</td>
<td>0.067±0.004</td>
</tr>
<tr>
<td>H$\delta$</td>
<td>169.4±2.9</td>
<td>169.4±4.3</td>
<td>0.056±0.003</td>
</tr>
<tr>
<td>H$\epsilon$</td>
<td>271.7±2.9</td>
<td>201.0±4.7</td>
<td>-0.029±0.003</td>
</tr>
<tr>
<td>He I 4922</td>
<td>166.4±7.3</td>
<td>309.1±8.4</td>
<td>-0.021±0.005</td>
</tr>
<tr>
<td>He II</td>
<td>227.8±1.6</td>
<td>204.6±2.5</td>
<td>0.090±0.002</td>
</tr>
</tbody>
</table>
Figure 4.10: Phase-binned, velocity-binned and continuum-subtracted trailed spectra. Deep blue represents the continuum flux, while bright red corresponds to the maximum flux.
4.8 Trailed spectra

the radial velocity curve of Hα from the previous section. We can clearly see now how the behaviour in Hα is consistent over all four nights. The eclipse is clearly seen, as is the weakening of the emission around phase 0.5. The most significant difference between the behaviour on each night is the transient blue-shifted narrow absorption feature at velocity \( \sim -200 \text{ km s}^{-1} \) (known as a tramline (Smith & Dhillon 1998)) which is visible from the start of the observations on the 5th and persists for about two orbits at constant velocity before becoming broader and increasing in blue shift for about half an orbit before it disappearing. It looks like it disappears at about phase 3.1 on the 5th, but there is a gap in the data for about 0.2 orbits beginning just after this time, and it is possible that the feature persists into this gap. However, it has undoubtably disappeared after this gap, around phase 3.4. No such feature appears on any other night. By averaging the spectra when the tramline velocity is constant, and when it broadens and moves slightly further to the blue, and then fitting gaussians to the line profile and absorption feature, the velocity of the feature was measured. The average profiles (black) and fits (red) are shown in Figure 4.11. Up until phase 2.8, the velocity is \(-176\pm20 \text{ km s}^{-1}\) with a width of \(-110\pm40 \text{ km s}^{-1}\), and for the last half orbit it is \(-230\pm20 \text{ km s}^{-1}\) with a width of \(-160\pm40 \text{ km s}^{-1}\). Figure 4.11 also shows clearly that the absorption is significantly stronger after phase 2.8 than before.

Studying the He I 6678Å data reveals far more complicated emission structure, as seen in Section 4.6 where it was sometimes single-peaked and sometimes double-peaked. The lower signal-to-noise in He I makes it difficult to follow the structure clearly in this line (right hand panels in Figure 4.10), so we will instead look at the behaviour of He I in the phase folded trails described below. However, one important feature which can only be studied in this unfolded trail is the transient absorption feature. On the 5th Jan, we can see very faintly the same transient feature as was seen in the Hα data. It is most clearly seen where it crosses the brightest regions of the trail, around phases 1.75 and 2.75, and when it is broadest around the eclipse at phase 3.0. Average profiles of He I are shown in Figure 4.11, covering the same tramline phase ranges as the Hα profiles. Above each one is plotted the fitted absorption profile from the corresponding Hα line, making it possible to identify the absorption feature in He I, although it is very close to the noise level.

As already noted, the signal-to-noise in He I is low, so in the upper two panels of Figure 4.12 trailed spectra of both lines are shown, this time folded onto one orbit, thus

2Note that in the 2.8-3.3 phase range, the slightly red-shifted absorption feature also had to be fitted and subtracted to obtain a reliable result. This is visible as the dip between the peak of Hα and the tramline.
4.8 Trailed spectra

Figure 4.11: Average line profiles for 5th January 1995 showing the tramline absorption. Red curves in the left two panels show the line profile and absorptions fitted as a sum of two (upper left panel) or three (lower left panel) gaussians. The red curves on the right show the gaussian fitted to the tramlines in Hα for comparison with the noisier He I profiles.

increasing the signal-to-noise. The black line plotted on the eclipses is the mid-eclipse time, measured manually from the lightcurve of each velocity bin (but from a trail with 100 phase bins per orbit and 100 km s⁻¹ velocity bins). The lower panels show trails reconstructed from Doppler maps, and will be discussed in Section 4.9. The Hα trail is as described above, but due to the higher S/N in the folded trail, we can now also identify the faint narrow slightly red-shifted absorption dip just before eclipse noted in Section 4.6 (it is marked with a white circle in the top left panel of Figure 4.12). We also now see clearly the significant weakening of the red-shifted side of the line around phase 0.4-0.5. The mid-eclipse is seen to be earliest in the blue wing and latest in the red wing, with a disturbance around velocity zero due to the absorption feature.
Figure 4.12: Top panels show phase-folded, phase-binned, velocity-binned and continuum-subtracted trailed spectra. Lower panels show corresponding trails reconstructed from Doppler maps (see Section 4.9).
The He\textsc{i} trail now has sufficient signal-to-noise to make the behaviour clear. We see two broad bright single-peaked regions of emission from just after eclipse to around phase 0.25 and from about phase 0.65 to phase 0.85, the latter identifiable with the brightest part of the He\textsc{i} orbital lightcurve. The bright region earlier in the orbit is blue-shifted, while the later bright region is red-shifted. Again, the mid-eclipse is earliest in the blue wing and latest in the red, telling us that material flowing towards us is eclipsed first, and receding material last. Between phases 0.3 and 0.6 the emission is double peaked, with the peaks separated by about 1000 km s\textsuperscript{-1}. The blue peak is broader than the red peak, suggesting that the double-peaks do not arise from disc emission, but rather from two S-waves, with the broader S-wave corresponding to emission from the same region as H\alpha. This complicated He\textsc{i} structure explains why the radial velocity measurement of this line could not be fitted with a sine wave.

As discussed already, this variation in the line profiles is like that seen in the SW Sex stars, with He\textsc{i} being most like that in DW UMa presented in Smith, Dhillon & Marsh (2001). The constant velocity blue-shifted feature we see on the 5th January is like the so-called ‘tramlines’ first seen in DW UMa (Smith & Dhillon 1998). In DW UMa three tramlines were seen, two blue-shifted and one red-shifted, although the broadening and increased blue shift we detect as the tramline in V348 Pup disappears was not seen in DW UMa. These tramlines are a new feature which any model of SW Sex stars must explain. We will discuss this further later in this chapter.

4.9 Doppler tomography

Figure 4.13 shows Doppler maps for each line for each night of the CTIO run. These maps were produced using the maximum entropy method (see Chapter 2 and Marsh & Horne (1988)). Spectra between orbital phases -0.1 and 0.1 were not used in producing the maps because Doppler tomography assumes that all regions in the map have constant visibility throughout the orbit, which is not true during eclipse. The geometry over plotted is a ballistic accretion stream and the Keplerian velocity along the stream, as described in Chapter 2. All Doppler map plots and modeling in this chapter assumes the orbital parameters $q = 0.31$, an orbital inclination of 81.1\degree, and $M_{wd} = 0.64M_\odot$ (estimated for V348 Pup in the previous chapter (Sections 3.5 & 3.10)). A systemic velocity of 44.9 km s\textsuperscript{-1} was assumed, as estimated from H\alpha in Section 4.7.

The H\alpha maps all look very similar, dominated by a broad region of emission centred in
Figure 4.13: Maximum entropy Doppler maps for each night.
the lower left quadrant of the maps. The hotspot, where the accretion stream impacts the
disc, should appear somewhere between the two arcs in the top left quadrant, as material
initially at the stream velocity merges with a Keplerian disc. The emission region in Hα is
not located at the hotspot, and indeed no emission feature at all is seen from the hotspot
velocity. This Hα emission corresponds to no known component of the binary, having too
low a velocity for even the outer edge of a disc at the tidal cut-off radius (the white circle
in the maps). There is faint low velocity emission in the right hand half of the maps, which
in the 3rd January map forms an arc of emission following the lower right section of the
white circle. In the 4th January map it appears as a separate region of emission centred
around \( V_y = 0 \). There is a sharp low-velocity arc in the top half of the 5th January map
which results from the transient blue absorption feature described earlier, and does not
indicate any such arc-like feature in the velocity-distribution of the emitting material.

HeI shows the same strong emission in the lower left quadrant of the maps, but this
time the region of emission on the right of the maps is also strong, being comparable in
brightness on the 4th and 6th January maps, and stronger on the 5th. The emission in
the lower left confirms the suspicion of the previous section that the HeI emission was a
sum of two S-waves, with one corresponding to the same emission source as Hα. An arc
of emission such as that seen in Hα on the 3rd January is very clearly seen on the 5th,
while the maps from the 3rd and 6th show evidence of a ring of emission whose velocity
is far too low to come from a Keplerian disc.

This is exactly the sort of anomalous structure seen in Doppler maps of the SW Sex
stars, e.g BH Lyn (Hoard et al. 1998a), PX And (Hellier & Robinson 1994), and the
Balmer emission in DW UMa (Hoard 1998; Smith, Dhillon & Marsh 2001). The Hβ
Doppler map of BH Lyn from (Hoard et al. 1998a) is shown in Figure 4.14, with the same
region of emission in the lower left quadrant as in all the V348 Pup maps, and the same
low velocity arc/ring of emission on the right.

Note that the appearance of the emission on 4th January as two separate blobs, partic­
ularly in HeI, appears to be a genuine difference in behaviour and not just a result of the
lower phase coverage on this night (phase ranges 0.16–0.27 were not observed). This was
tested by producing a Doppler map using the data from all four nights of observations, but
leaving out those phases missing from the observations on the 4th (as well as the eclipse
phases). Leaving out the missing phases produced a map which looked no different from
a map including these phases.
The Doppler maps produced combining the data from all four nights, leaving out only the eclipse phases, are shown in Figure 4.15. These are average Doppler maps for the run, and as such show the same detail as described above. The faint ring of emission in He I is particularly clear thanks to the high signal-to-noise of these maps. Trailed spectra reconstructed by forward projecting these maps are shown in the lower two panels of Figure 4.12. Comparison of these reconstructed trails with the phase-folded observed trails in the top panel demonstrates how well these Doppler maps represent the emission distribution in the system. The Doppler maps contain no spatial information, and thus cannot and do not attempt to reproduce the eclipse of the line profile or the phase dependent absorption/occultation features. The eclipse phases are left out of the reconstructed trails. Apart from this, we see that the structure of the observed trails is very well reproduced by the Doppler maps - the variations in line profile, including the change between double-peaked and single-peaked emission in He I are all reproduced.

4.10 Discussion

4.10.1 V348 Pup: A new SW Sex star

These spectroscopic observations of V348 Pup reveal most of the defining characteristics of the SW Sex stars. V348 Pup shows broad, single-peaked emission lines whose radial velocity curves are delayed relative to the white dwarf velocity, corresponding to a region of emission in the lower left quadrant of the Doppler maps. There is absorption in He I which is strongest around phase 0.5, and a transient constant-velocity blue-shifted absorption
tramline which appears only on one night. The lightcurves of Hα and He I 6678Å reveal shallower eclipses than those of the continuum, with weaker emission around phase 0.5.

4.10.2 Models for the SW Sex phenomenon

Having demonstrated that V348 Pup should be considered another member of the SW Sex subclass of nova-like CVs, I now briefly outline the various models for this phenomenon, before considering one model in more detail. A valuable recent review of several models is given in Smith, Dhillon & Marsh (2001).

Accretion disc wind

The first attempt to explain the peculiar behaviour of SW Sex was made by Honeycutt, Schlegel & Kaitchuck (1986). They suggested a radiation-driven bipolar accretion disc wind, which could produce single-peaked emission lines, including narrow high-excitation lines, with shallower eclipses than in the disc continuum. However this model does not explain the phase-lag in the radial velocities of the emission lines, the low excitation emission in the lower left quadrant of Doppler maps, or the tramline absorptions. There is also no convincing explanation for the phase 0.5 absorption features from this model, with the absorbing material needing a similar vertical extent to the wind.
4.10 Discussion

*Magnetically channeled accretion flow*

Williams (1989) suggested that nova-likes exhibiting single-peaked emission line profiles might be magnetic systems without accretion discs, where the accretion stream is channeled directly onto the magnetic white dwarf by field lines. Recent discoveries of modulated circular polarisation in the SW Sex candidate LS Peg and in SW Sex itself (Rodríguez-Gil et al. 2001), and periodic flaring in emission lines in DW UMa (Smith, Dhillon & Marsh 2001) and BT Mon (Smith, Dhillon & Marsh 1998), behaviour typical of intermediate polars (e.g. FO Aqr, Marsh & Duck 1996), suggests that magnetic fields do play an important role in the behaviour of SW Sex systems. Rodríguez-Gil et al. (2001) propose that SW Sex systems are high accretion rate intermediate polars, with the single peaked lines coming from the shocked regions where material overflowing a magnetically truncated disc meets with the magnetosphere. However there has been no detailed modeling showing how or if this behaviour can explain the structure seen in Doppler maps of SW Sex systems, the phase 0.5 absorptions or the tramlines. Such modeling would be valuable.

*The stream-disc impact model*

Recently Groot, Rutten & van Paradijs (2001) suggested an alternative model after a detailed study of SW Sex in a rare low accretion state. They proposed that the single-peaked emission comes from above the plane of the disc in the stream-disc impact region. A hotspot photosphere produces transient absorption features around phase 0.82-0.94, and they suggest material overflowing the hotspot then absorbs continuum emission from the hotspot around phase 0.5, explaining the phase 0.5 absorption of SW Sex systems, even though this feature was not seen in their SW Sex observations. However the Doppler maps of V348 Pup and other SW Sex systems, which are dominated by emission from low velocities in the lower left quadrant of the map, are not consistent with emission from the hotspot. This different behaviour found by Groot, Rutten & van Paradijs (2001) probably results from differences between the low state of SW Sex they observed and the high state in which SW Sex systems are usually seen. This model provides no explanation for the tramlines.

*Stream overflow models*

One of the most popular explanations for the SW Sex phenomenon is that the accretion stream does not simply impact and merge with the disc at its edge, but instead some
fraction overflows the disc edge, finally impacting the disc closer to the white dwarf. Emission from an overflowing stream was first suggested by Shafter, Hessman & Zhang (1988), and has been considered in some detail by Hellier e.g. in the case of V1315 Aql (Hellier 1996), with a review in Hellier (2000). A more complicated stream-overflow model was suggested by Hoard et al. (1998a) for the case of UU Aqr. In the model described in Hellier (2000), the overflowing stream is fainter than the disc and so produces absorption of the disc emission. If the disc is flared, then this absorption is seen predominantly around phases 0.4-0.6, and explains the double-peaked nature of SW Sex emission lines around this phase. The region closer to the white dwarf where the stream re-impacts the disc is seen in emission producing a single-peaked S-wave component. This component is in the lower left quadrant of the Doppler map, but the velocity is much higher than that of the emission in observed maps, such as those of V348 Pup. To explain this discrepancy, it is necessary to assume a spread in stream-velocities after the initial impact with the edge of the disc, so that a significant proportion of the stream is at a lower velocity than the ballistic trajectory. The bright region in the lower left of the Doppler map is then the sum of emission from the low-velocity parts of the stream re-impact and emission from the disc itself. To match the observed maps, this model also requires the disc to be substantially sub-Keplerian, interpreting the low-velocity ring as emission from the outer edge of the disc. In the case of V348 Pup, where the ring has a velocity radius of 200-400 km s\(^{-1}\) (measured from the He I Doppler maps in Figure 4.15) this would require a disc whose velocities are sub-Keplerian by at least 25%. Various modifications are needed to this model to explain phase-dependent absorption features, such as as azimuthal variations in the disc flare (Hoard, Thorstensen & Szkody 2000).

\textit{The magnetic propeller model}

The final model to be discussed here involves both stream-overflow and magnetic accretion. In this model, proposed by Horne (1999), the accretion stream is considered as being made of diamagnetic blobs of plasma which, on encountering a rapidly rotating magnetic field anchored in the accretion disc, are accelerated and ejected from the system. This is a modification of the model used to explain the unusual behaviour of the CV AE Aqr, notably extreme flickering and single-peaked low-velocity line emission located in the bottom left quadrant of the Doppler map (Wynn, King & Horne 1997; Eracleous & Horne 1996). In AE Aqr, the blobs are thought to be ejected by the magnetic field of the rapidly
spinning \((P_{\text{spin}} = 33s)\) white dwarf. The trajectories of the blobs under the influence of this 'magnetic propeller' are modified so that they pass through the region of the Doppler map corresponding to the line emission. The trajectories of blobs with different densities cross in this velocity region, with the line emission and the flaring behaviour resulting from 'fireballs' as the plasma blobs collide (Horne 1999).

Horne (1999) suggested that this magnetic propeller effect could explain most, if not all, of the unusual phenomena seen in SW Sex systems. Since there is no evidence for rapidly spinning white dwarfs in SW Sex systems (Rodríguez-Gil et al. (2001) find \(P_{\text{spin}} \approx 30\) min for LS Peg), Horne (1999) proposed that in SW Sex systems, the magnetic field responsible for ejecting the blobs from the system is anchored in the disc. The scenario is then one in which a fraction of the accretion stream overflows the disc edge, encounters the spinning magnetic field anchored in the disc and is ejected from the system in a similar manner to the propeller in (probably disc-less) AE Aqr. In the rest of this chapter we produce a simple model for a disc-anchored propeller in V348 Pup and compute blob trajectories, trailed spectra and Doppler maps for comparison with our observations.

### 4.10.3 A disc-anchored magnetic propeller in V348 Pup?

To model a disc-anchored propeller in V348 Pup, we adopt the prescription for the magnetic flow first proposed by King (1993) and considered further by Wynn & King (1995). This model was used by Wynn, King & Horne (1997) to study the propeller in AE Aqr, and by Hynes et al. (2001) to model the SW Sex-like behaviour in the neutron star system XTE J2123-0.58, and to model the behaviour of several systems by Wynn (2001).

In this model, the magnetic acceleration of a diamagnetic blob is

\[
a_{\text{mag}} = -k \left[ \vec{v} - \vec{v_f} \right]_\perp
\]

where \(\vec{v_f}\) is the magnetic field velocity, \(\vec{v}\) is the velocity of the blob and \(\perp\) denotes the component perpendicular to the magnetic field. The \(k\) parameter contains the details of the interaction between the plasma and the magnetic field, and will depend on the field strength, \(B\), and the blob length, \(l\), and density, \(\rho_b\). From King (1993),

\[
k \propto B(r)^2 l(r)^2 c_A(r)^{-1}
\]

where \(c_A\) is the Alfvén speed in the interblob medium, the speed at which a disturbance in the magnetic field propagates along a field line. As in Wynn, King & Horne (1997), we assume a sufficiently tenuous interblob medium that \(c_A\) is constant at the speed of light.
For a magnetic field of $10^5\text{G}$, this requires a density $< 10^{-12}\text{g cm}^{-3}$. From Wynn & King (1995), assuming the temperature of the blobs changes along their path much more slowly than their density, $l \propto B^{-2/3}$. Therefore, from Equation 4.2 we have

$$k(r) \propto B(r)^{2/3}. \quad (4.3)$$

The studies of the white dwarf propeller in AE Aqr (Wynn, King & Horne 1997) and the disc-anchored propeller in XTE J2123-058 (Hynes et al. 2001) both assumed a simple magnetic dipole for the field structure. While this might be a reasonable approximation for the white dwarf anchored propeller, a disc-anchored field is likely to have a more irregular structure, with loops of field of various strengths anchored in the disc at various radii, as discussed in Horne (1999) and Hynes et al. (2001). We therefore adopt an alternative, very simple, model for a disc-anchored field. We consider a magnetic field with only a vertical component (i.e. perpendicular to the orbital plane), with the angular velocity of the field at a radius $r$ from the white dwarf having the Keplerian value corresponding to radius $\alpha r$, i.e.

$$\omega_f(r) = \sqrt{GM_{wd}/\alpha^3 r^3}. \quad (4.4)$$

With $0 < \alpha \leq 1$ this provides a simple parametrization allowing a blob at radius $r$ to encounter a field anchored in the disc at radius $\alpha r$.

The magnetic dynamo effect is thought to generate magnetic fields in accretion discs and contribute to the disc viscosity (see Section 1.4). The dynamo is generated by the velocity shear in the disc (Tout & Pringle 1992). Therefore, in our very crude approach we assume that the magnetic field is proportional to the magnitude of the disc shear, giving

$$B(r) \propto \left| \frac{dv_{kep}}{dr} \right| = \left| \frac{d}{dr} \left( \sqrt{GM_{wd}/r} \right) \right| \propto r^{-3/2}. \quad (4.4)$$

Substituting this into Equation 4.3 gives the radial dependence of the magnetic drag parameter,

$$k(r) \propto r^{-1}. \quad (4.4)$$

For a white dwarf anchored dipole propeller in AE Aqr Wynn, King & Horne (1997) considered the cases $k \propto r^{-3}$ and $k \propto r^{-2}$, while Hynes et al. (2001) also consider $k \propto r^{-2}$ for the case of XTE J2123-058.

We are now in a position to calculate the trajectories of blobs in the system, taking into account the magnetic drag given in Equation 4.1 with $k$ varying as in Equation
4.10 Discussion

4.4, \( k(r) = k_0(r_0/r) \), so that \( k = k_0 \) at \( r = r_0 \), where we use \( r_0 = 10^8 \text{m} \) (as used in Wynn, King & Horne (1997)). This \( r_0 \) is about 1/3 of the primary Roche lobe radius in V348 Pup. Trailed spectra were produced by placing blobs at uniform time intervals along trajectories, where different trajectories were calculated with a uniform spread of values of \( k_0 \) between \( k_a \) and \( k_b \). The magnetic field is switched off (by setting \( k \) to 0) outside radius \( r_{\text{mag}} \). \( r_{\text{mag}} = 0.4a \) (\( = 0.83R_L \)) is used, consistent with the V348 Pup disc sizes measured in Chapter 33. The blobs are treated as spheres of radius \( r_{\text{blob}} \), and a blob is treated as being invisible if its line of sight to the observer passes through any other blob (or the donor star). This is effectively assuming that the blob region is optically thick in the line. This enables us to produce trailed spectra and emission line lightcurves for a spread of propeller trajectories, accounting for eclipse by the donor and of blobs by each other. The intrinsic line profiles of the blobs are treated as gaussians, with thermally broadened widths at temperature 3300K, the photospheric temperature of the donor, \( T_{\text{phot}} \). \( T_{\text{phot}} \) comes from the stellar models by Baraffe et al. (1998). The blob radii were set at 0.001\( a \) to produce appreciable occultation effects by blobs. Note that we stated earlier that the blob size will change along the trajectory, but this constant blob size fixed here is only used for calculating occultation by blobs. Similarly, the constant blob temperature assumed here is only used to estimate linewidths. The brightness of each blob was assumed to be proportional to the magnetic dissipation rate, \( S_{\text{blob}} \propto k[\vec{v} - \vec{v}_j]^2 \). This was speculated in Wynn, King & Horne (1997) as a possible form for the emission line brightness in AE Aqr. We also consider the case of emission from shocks where trajectories cross, although this is not included in our model trailed spectra. The material was assumed to escape through the \( L_1 \) point at speed \( C_s \), the photospheric sound speed of the donor (Warner 1995). Each trajectory had a random initial speed, of order \( C_s \), estimated using

\[
C_s \approx 10^4 ms^{-1} \sqrt{\frac{T_{\text{phot}}}{10^4 K}}.
\]  

(4.5)

The angle between the initial stream velocity and the \(-x\) direction is given a normal distribution, with average value 0° and standard deviation 90°. This gives a reasonable spread of initial velocities for the stream as it escapes from the Roche lobe of the donor through \( L_1 \). The trailed spectra were convolved with a gaussian of FWHM 50 km s\(^{-1}\) to simulate the effect of instrumental broadening.

We now present the results from using two different sets of parameters with interesting

\(^3\)Note that we would actually expect the disc-anchored field to overshoot the disc edge so that \( r_{\text{mag}} = r_{\text{disc}}/\alpha \), but here the magnetic drag is assumed to switch off at the disc edge, independent of \( \alpha \).
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results. Figures 4.16 and 4.18 show the results of the model for two different sets of parameters. Each of these figures has four panels. The top left panel shows the trailed spectrum, with deep blue showing the minimum (zero) brightness and red showing the maximum. The top right shows a spatial plot of the system, with the donor star in red, and five stream trajectories with values of $k_0$ uniformly spaced between between $k_a$ and $k_b$. The colouring along each stream trajectory reflects the brightness of the emission line at each point, assuming it is proportional to the magnetic dissipation. The purple blobs mark the region where the stream trajectories cross. The middle right panel is a Doppler map of the computed trail, produced using the Fourier-filtered back-projection technique, leaving out phases -0.1 to 0.1 (the eclipse). The bottom panel shows the orbital lightcurve of the emission, including the occultation effects.

Field anchored at radius 0.9$r$

The first model, shown in Figure 4.16, has $\alpha = 0.9$, i.e. the magnetic field at a radius $r$ from the white dwarf is anchored at radius 0.9$r$. The value of $k_0$ ranges from $0.2 \times 10^{-3}$ to $1.5 \times 10^{-3}$. Each stream trajectory corresponds to the path of a blob lasting 0.4 orbital periods. We do not attempt to explain the mechanism by which the stream overflows the disc, and indeed these computed trajectories are in the orbital plane. Modeling of stream-overflow, such as that of Lubow (1989), reveals that the stream is not likely to get much past closest approach to the white dwarf before impacting with the disc. Horne (1999) suggested that the loops of magnetic field anchored in the disc would both accelerate the diamagnetic blobs tangentially and lift them up away from the disc, allowing the overflow to last longer over the disc before impacting it or being ejected from the system. The model here only includes a vertical component of magnetic field, and so from Equation 4.1 it will not simulate any such lifting of the stream by the magnetic field. Taking this into account, it is reasonable to expect the blob trajectories to last for the 0.4 orbits shown here. However, noting that these trajectories continue to orbit the white dwarf in the disc region for several orbits after those shown here, it seems likely that the blobs would eventually crash down onto the disc before gaining enough energy from the magnetic field to leave the system.

Inspecting the spatial plot, we see that there are two regions of emission, the first one where the stream first encounters the magnetic field and so the difference between stream and field velocities produces substantial magnetic dissipation, and the second shortly after
Figure 4.16: Model for disc propeller anchored at 0.9r.
the closest approach of the stream to the white dwarf. The stream trajectories all cross after the stream has made about three-quarters of an orbit of the white dwarf, shortly after the second region of high magnetic dissipation. The trailed spectrum of these trajectories is complicated, with 'S-waves' from each region of emission having thickness variations as large as their amplitudes. This results from the long thin shape of the emission regions in velocity space being projected at different orientations throughout the orbit. The peaks are blurred and least distinct around phases 0.3 to 0.4 and about 0.8 to 0.9, with these two region of blurred emission being blue- and red-shifted respectively. The trough between the peaks is most prominent around phases 0 to 0.2 and 0.5 to 0.7. The eclipse begins in the blue and ends in the red, with a total phase difference in mid-eclipse time between the two sides of the line of about 0.1. Comparing with the observed trails in Figure 4.12, we see that there is no resemblance between this synthetic trail and the Hα trail, which has a single broad S-wave. However there are some similarities with the behaviour of the He I trail. The difference in eclipse time between the red and blue side of the line is very similar, and the change between broad single-peaked emission and double-peaked emission in the observed maps is a little like the change between distinct single peaks and blurred peaks with a shallower trough between them in the computed map, including the red and blue-shift of the two single peaked regions. However the behaviour in the computed trail is delayed by about 0.2 in phase relative to the observed trails.

Studying the Doppler map of the synthetic data reveals two regions of emission on the left and right of the map, something like that in the observed He I maps but not the Hα maps (Figure 4.15). We can clearly see now that the main differences between the computed and observed He I maps are that the regions of emission in the observed map are too sharp, have velocities which are too high, and are rotated clockwise about the white dwarf velocity relative to the observed map, corresponding to the phase difference between the computed and observed trails. If the velocity distribution of the stream after the first encounter with the disc spread to lower velocities, as suggested in the stream-overflow model of Hellier (1996), the emission region on the left of the computed map would extend to lower velocities, matching the observed map better. The models of the stream-disc impact by Armitage & Livio (1998) do suggest a greater spread of velocities after the initial stream-disc impact, so this is plausible.

We also note from the Doppler map that the velocity of the stream cross-over region coincides with the bright blob on the right of the observed He I Doppler maps, also seen
weakly in the Hα maps. Figure 4.17 shows the observed maps with the model blob trajectories overplotted. So if the emission actually arises where high-velocity blobs impact with lower velocity blobs at the cross-over region, as suggested for the propeller in AE Aqr (Horne 1999), then this disc-anchored propeller provides a possible explanation for the emission on the right of the Doppler map.

Figure 4.17: $\alpha = 0.9$ model trajectories (white) and trajectory cross-over velocities (purple circles) plotted over the observed Doppler maps.

Finally, we note that the model lightcurve shows some significant similarities with the observed emission line lightcurves of both Hα and He I (Figure 4.6), with the emission line flux weakening around phase 0.5-0.6, and the post-eclipse eclipse flux being weaker than the pre-eclipse flux. Such behaviour is seen in the observations of both emission lines, although the phase 0.5-0.6 dip is more pronounced in the observations, being strongest in He I. The cause of the dip around phase 0.5-0.6 is the obscuration of the brightest region of the accretion flow where it first encounters the disc by the blobs slightly further downstream.

Field anchored at radius 0.5r

The second model we use assumes the magnetic field to be anchored much further in towards the white dwarf: a blob at radius $r$ sees a magnetic field anchored at radius 0.5$r$ (corresponding to $\alpha = 0.5$). With this prescription we would expect to see behaviour more
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like the rapidly rotating dipole models for the propeller, since a field anchored further in
the disc will rotate significantly faster than the local disc and stream velocity, and therefore
would be expected to propel the blobs quickly out of the system. It was found that values
of \( k_a = 0.35 \times 10^{-3} \) and \( k_b = 0.7 \times 10^{-3} \) respectively with trajectories lasting \( 0.8P_{\text{orb}} \)
produce blob stream trajectories like those suggested for AE Aqr (Wynn, King & Horne
1997) and the neutron star system XTE J2123-058 (Hynes et al. 2001). Figure 4.18
shows the computed model. The blobs swoop at high velocity past the white dwarf and
are quickly accelerated out of the system, with high velocity blobs colliding with slower
blobs around one to two orbital separations to the left of the white dwarf. The blobs
then decelerate, moving outwards at constant velocities of 200–500 km s\(^{-1}\). This stage
corresponds to the constant velocity loops within the white circle in the Doppler map.

The model trail again shows a complicated sum of S-waves of variable width, and a
blue-red eclipse. The emission is again broadest around phase 0.4 and about 0.8, with
these two regions being blue- and red-shifted respectively. However, the troughs between
the peaks are deep for most of the orbit, unlike the \( \alpha = 0.9 \) model where they were filled
in by blurred crossing S-waves for much of the time. Inspecting the model Doppler map
we see an arc of emission on the left hand side of the map, bright where the trajectories
are close together as the blobs first encounter the disc, and bright again at the point in the
trajectory with lowest \( V_y \), as the blobs coast past the white dwarf at maximum velocity.
The behaviour bears no resemblance to the broad single emission we observed in H\( \alpha \) and
is rotated clockwise by 90° relative to the He I map. If the blob trajectories are a good
representation of the true accretion flow, the magnetic dissipation model for the hotspot
emission is not. The predicted lightcurve does have the phase 0.5 dip again, although this
is because the dominant emission region along the trajectories using this model is still the
region between first encounter with the disc and closest approach to the white dwarf, as
in the \( \alpha = 0.9 \) model.

We therefore turn our attention once more to the blob-collision model for the emission
(Horne 1999). We see in Figure 4.19 that the region where the blobs collide corresponds
exactly to an arc through the middle of the region of emission in the lower left quadrant
of the H\( \alpha \) and He I Doppler maps. As discussed in Hynes et al. (2001), if the region in
which the blobs collide is optically thick in the lines, then the emission will come mainly
from the inside of the collision region. This emission would then be obscured by the outer
edge of the region in the phase range about 0.25 to 0.7, just when we observe the dip
Figure 4.18: Model for disc propeller anchored at 0.5r.
Figure 4.19: $\alpha = 0.5$ model trajectories (white) and trajectory cross-over velocities (purple circles) plotted over the observed Doppler maps.

in the emission line lightcurves. This model can therefore explain quite successfully the single-peaked S-wave emission in Hα including its lightcurves, and can also explain the corresponding component of the He I emission.

It should be noted that for the blob collision region to be eclipsed by the donor star, a higher inclination is required. This can be seen from Figure 3.6, where the “shadow” of the donor star does not reach far enough out of the primary Roche lobe to eclipse the blob collision region. For an inclination of 90°, a mass ratio $q = 0.16$ is required to explain the eclipse width measured in Section 3.5. For the superhump period excess, $\epsilon$, of V348 Pup, a mass ratio of about 0.16 is consistent with the $\epsilon$-$q$ relation from models of eccentric discs presented in Murray (2000). This would not affect the results of Chapter 3, because the eclipsed region of the primary Roche lobe is insensitive to this change in $i$ and $q$ as long as $q$ is chosen to maintain the observed eclipse width at half flux for the new value of $i$. Note that using Equation 3.3 to estimate the donor mass, $q = 0.16$ implies a white dwarf mass of 1.23M$_\odot$, twice the typical white dwarf mass in CVs.

The material leaving the system provides a convincing explanation for the transient blue-shifted tramline seen in V348 Pup and the blue-shifted lines in DW UMa (Smith, Dhillon & Marsh 2001). This material is coasting outward at a constant velocity of less than about 500 km s$^{-1}$. A particularly dense clump of such material might cross our line of sight to the system, leading to a blue-shifted constant velocity absorption feature. Bearing
in mind that the expelled material is moving radially outwards (the apparent clockwise motion in the spatial plot and Doppler map is because these are both measured in the anticlockwise moving corotating frame), it is plausible that such a clump of absorbing material would remain in our line of sight for a couple of orbits. We might even have fun and speculate that the sudden broadening and blue-shift of the tramline in V348 Pup shortly before it disappears is the result of a faster moving clump of gas colliding with the first one, briefly providing a larger absorbing clump of material with a larger velocity dispersion moving more quickly towards us, before both clumps move out of the line of sight. This would explain both the disappearance of the tramline, and the broadening, increased blue-shift and deepening of the tramline just before it disappears.

Horne (1999) also suggests that the low-ionization absorption lines seen around phase 0.5 in SW Sex stars are caused by the exit stream passing in front of the white dwarf. The only phase 0.5 absorption we see in V348 Pup is the red-shifted He I 4471 Å emission which is strongest around phases 0.3 to 0.6, which does correspond to phases when the exit stream is in front of the white dwarf. However there is a significant problem with this explanation for the He I absorption: the absorption has a red-shift of about 1000 km s\(^{-1}\), while the exit stream in this range has a would lead to absorption with a blue-shift of around a few hundred to ~1000 km s\(^{-1}\). That this is not He I absorption, but absorption from some unidentified line is unlikely, given that such a feature was seen in a different He I line (5876 Å) in PX And (Thorstensen et al. 1991). However this may still be a plausible explanation for other phase 0.5 absorption features in other systems.

The low velocity ejected material has the same range of velocities as the ring of emission seen in some systems, such as in the average He I Doppler map of V348 Pup. This ring in the map does not require the presence of a ring of material. A single clump of emitting material with a constant velocity component towards or away from the observer would produce a ring of emission in a Doppler map if it persisted for at least an orbital period. A clump of the expelled material seen in emission could cause exactly such a feature in the Doppler maps. Work is being done on modeling the time-dependent emission from the fireballs resulting from blob collisions in AE Aqr, which suggests that they can remain in emission long enough to explain this low-velocity ring in the Doppler maps (K. Horne, private communication).
4.11 Conclusions

This simple modeling with $\alpha = 0.5$ shows that magnetic fields anchored in the accretion disc can produce the same propeller effect as a rapidly rotating dipole like that invoked to explain the behaviour of AE Aqr (Wynn, King & Horne 1997), as suggested by Horne (1999), and that it can explain most of the peculiar features of the SW Sex stars if the emission is dominated by the collisions of the diamagnetic blobs. Assuming the emission follows the magnetic dissipation rate predicts behaviour which shows some similarities with observations, but the collision model for the emission mechanism seems much more promising.

The magnetic field structure is likely to be far more complicated than that assumed here, and the suggestion of Horne (1999) that material flowing higher above the disc plane might encounter fields anchored at smaller radii than material closer to the disc plane would suggest that we might see a combination of the behaviour seen here for the models with $\alpha = 0.5$ and $\alpha = 0.9$. That would then enable us to explain both emission regions in the Doppler maps as coming from fireballs as blobs collide, but from blobs flowing at different heights above the disc. In addition, this would lead to the observed blue-red eclipse, since around phase -0.1 to 0.1, the region of the trajectories just after collision in the $\alpha = 0.9$ model would be blue-shifted and eclipsed earlier than the red-shifted cross-over region for the $\alpha = 0.5$ model. We would expect the energy released when the blobs collide to be dependent on the velocity dispersion of the trajectories in the stream cross-over region, which from Figure 4.18 is $\sim 600 \text{ km} \text{s}^{-1}$ and from Figure 4.16 is $\sim 300 \text{ km} \text{s}^{-1}$. We would therefore expect the emission region on the right of the Doppler maps to be cooler than that on the left, possibly explaining why the emission in the right hand region is stronger relative to the left hand region for the lower-energy He I transition than for Hα. Using the $k_0=0.35-0.7\times10^{-3}$ as applied to the $\alpha = 0.5$ model, trajectories for $\alpha=0.6$ produce much the same propeller effect as $\alpha=0.5$, but the streams cross further to the right in the Doppler map (around the bottom edge of the primary Roche lobe in velocity space). Values of $\alpha=0.7$ and 0.8 yield more complicated trajectories, but with very low velocity dispersions at the first stream cross-over region, and so predict no significant emission from blob collisions. Using the larger spread of $k$ values ($k_0=0.2-1.5\times10^{-3}$ as applied to the $\alpha = 0.9$ model) with $\alpha =0.6$, 0.7 and 0.8 increases the velocity dispersion at the cross-over regions, but still does not suggest any new emission region not already discussed in the previous sections. This suggests that even for a spread of effective values
of $\alpha$, the likely regions of emission from ‘fireballs’ as blobs collide are those seen in the observed maps, although using the larger spread of $k_0$ with $\alpha=0.5$ does lead to the arc of stream cross-over regions extending to higher $V_y$ towards the donor velocity.

The analysis above was repeated using the relation $k(r) \propto r^{-2}$, corresponding to the magnetic dipole relation $|B| \propto r^{-3}$, as used by Wynn, King & Horne (1997) and Hynes et al. (2001). The behaviour of the trajectories and locations of cross-over regions was very similar, making our conclusions robust to using this alternative model for the magnetic flux density.

Horne (1999) notes how the ejected material carries away angular momentum from the inner disc without causing any local emission, leading to a cooler inner disc than predicted by steady-state disc models, explaining the V-shaped continuum eclipses of SW Systems.

There are still problems with this model, notably the lack of an explanation for the 1000 km s$^{-1}$ red-shifted phase 0.5 narrow He I absorption, and the slightly red-shifted phase 0.9–0.1 narrow H$\alpha$ absorption. Nevertheless, the disc-anchored propeller provides the most complete explanation so far for the SW Sex phenomenon.
Chapter 5

Late superhumps and the stream-disc impact in IY UMa

5.1 Introduction

IY UMa was identified as a dwarf nova type cataclysmic variable in January 2000 (Uemura et al. 2000b) when it went into superoutburst showing strong superhumps and deep eclipses in its lightcurve.

A detailed study of the superoutburst of IY UMa was carried out by Patterson et al. (2000) (hereafter called P2000). Common superhumps were seen at the height of superoutburst, and were replaced by strong late superhumps for about one week after the steep decline. IY UMa has orbital period $P_{\text{orb}} = 1.77$ hours. The system parameters were estimated in Patterson et al. (2000) as $M_{\text{wd}} = 0.86 \pm 0.11M_\odot$, $M_{\text{donor}} = 0.12 \pm 0.02M_\odot$ and orbital inclination $i = 86.8 \pm 1.5$. This high inclination means that we are viewing the disc almost edge on so that it contributes only a small fraction of the optical luminosity during quiescence. Indeed the white dwarf and stream-disc impact (bright spot) are the dominant light sources during quiescence (see Fig 5.1). The following work is based on that published in Rolfe, Haswell & Patterson (2001).

5.2 Observations

The data analysed here comprise 7 nights of V-band photometry covering the late superhump era, HJD 2451572–79. This is a subset of the superoutburst data presented in P2000. There are 22 eclipses observed with typical time resolution 20 to 30 seconds. Figure 5.1 shows a representative orbital lightcurve, where the strong orbital hump is visible
5.3 Ephemerides

Throughout the study of IY UMa in this chapter, the orbital phase is calculated using the white dwarf mid-eclipse ephemeris \( T_{\text{mid}} = \text{HJD} 2451570.85376(2) + 0.07390906(7)E \) and late superhump phase is calculated using the late superhump maximum ephemeris \( T_{\text{late}} = \text{HJD} 2451571.731(3) + 0.07558(6)E \). Both ephemerides are from P2000.

Figure 5.1: An example orbital lightcurve of IY UMa during the late superhump era which clearly shows the white dwarf and hotspot eclipse ingresses and egresses. This orbit has mid-eclipse at HJD 2451578.69.

as the stream disc impact region (bright spot) comes into view before eclipse. At the time of this particular orbit, the bright spot was the dominant light source in the system. The sharp white dwarf and hotspot eclipses are easily identified, with the white dwarf being eclipsed first, followed by the hotspot 3 minutes later. The white dwarf emerges first from eclipse 7 minutes after its ingress, followed by the hotspot eclipse egress after a further 6 minutes.

The entire dataset being analysed is shown in Figure 5.10, and is discussed in Section 5.5.1.
5.4 Tracing the eccentric disc shape

As in the study of V348 Pup eclipses (Chapter 3), the eclipses in IY UMa provide detailed information on the spatial distribution of emission in the system. In IY UMa, however, the sharply defined eclipses of the prominent hotspot tightly constrain its location. Eclipses throughout the disc precession cycle constrain the shape of the disc, and this information was fully exploited as described below.

5.4.1 Shadow method

Assuming the surface of the donor star is described by its Roche potential surface, and if the mass ratio, $q$, and orbital inclination, $i$, are known, the region of the orbital plane eclipsed by the donor at any particular orbital phase can be calculated. Figure 5.2 shows how these "shadows" cast by the donor star at phases corresponding to the start and end of hotspot eclipse ingress and egress constrain the location of the hotspot in the orbital plane to a small diamond shaped region. This method was used by Wood et al. (1986b). The hotspot is the point where the accretion stream impacts the edge of the disc, so determining the hotspot location provides a constraint on the location of the disc edge. Assuming the disc precesses at the beat period between the orbital and superhump periods (Equation 1.7), the geometry shown in Figure 5.3 was used to trace out the outline of the disc edge using those eclipses where both ingress and egress of hotspot eclipse could be measured. It should be noted that the hotspot may lie within the outermost orbits of disc material, but this method should still provide reliable constraints on the disc shape and its minimum size.

5.4.2 Stream trajectory method

For some eclipses, only the egress of hotspot eclipse could be measured, and so the technique described above could not be employed for those eclipses. However, there is an alternative technique for determining the location of the hotspot which can be employed in this situation (Hessman et al. 1992). Since the hotspot lies on the accretion stream, and the times of hotspot egress depend on the distance of the hotspot from the line of centres of the two stars, a ballistic trajectory for the accretion stream can be used to infer the position of the hotspot along the stream. Hence the location of the hotspot and the disc edge can be determined. Figure 5.4 shows the variation of eclipse egress phases along a ballistic stream trajectory.
Figure 5.2: "Shadows" cast by the donor star at the beginning and end of hotspot ingress (observer looking in directions 1 and 2 respectively) and at start and end of hotspot egress (directions 3 and 4) constrain the location of the hotspot in the orbital plane to a small diamond shape.
Figure 5.3: The geometry of the system used to study the hotspot eclipses. This figure shows the situation at orbital phase 0.11, shortly after a hotspot eclipse. The disc coordinates \((X_{\text{disc}}, Y_{\text{disc}})\) are centred on the white dwarf, with orientation determined from the disc precession phase (see Section 5.4.3). The stream impacts the disc edge (thick dot-dashed outline) at the hotspot, at angle \(\alpha_{\text{orb}}\) to the line of centres of the two stars, which itself makes an angle \(\theta_{\text{orb}}\) with the line of sight. The X-axis of the disc makes an angle \(\theta_{\text{disc}}\) with the line of sight, and so the position of the stream-disc impact in the precessing disc frame makes an angle \(\alpha_{\text{disc}} = \alpha_{\text{orb}} + \theta_{\text{orb}} - \theta_{\text{disc}}\) with the disc X-axis. \(\beta\) is approximately the angle between the stream and disc velocities at the point of impact.
5.4 Tracing the eccentric disc shape

Figure 5.4: Orbital phase of eclipse egress for points on a ballistic stream trajectory calculated assuming the shape of the donor star to be described by its Roche potential surface. The solid curve represents a stream started from the L1 point with \((V_x, V_y) = (-C_s, 0)\) (see Section 5.4.2) while the dotted curves are streams started with \((V_x, V_y) = (-C_s, \pm C_s)\). The crosses are the measured egress phases at their inferred positions along the stream.
5.4 Tracing the eccentric disc shape

The uncertainty in the time of each point of contact translates mostly to an error in the radial distance of the hotspot from the white dwarf for an assumed $q$. Azimuthal error was estimated by considering three slightly different stream trajectories with initial velocities at the inner Lagrange point (with axes as shown in Figure 5.6) of $(V_x, V_y) = (-Cs, 0)$, $(-Cs, Cs)$ and $(-Cs, -Cs)$. $Cs$ is the sound speed at the L1 point, calculated as in Section 4.10.3 using Equation 4.5. The different hotspot positions coming from these three slightly different trajectories provide small near-azimuthal error bars on the hotspot locations deduced using this method.

5.4.3 Combining the methods

The disc radius determined using the stream trajectory method depends on whether the start, middle or end of egress is used. For those eclipses where both ingress and egress were measured, using the mid-egress times to locate the hotspot on the accretion stream gives disc radii consistently greater than with the robust "shadow" method whose results completely constrain the location of the hotspot in the orbital plane (Figure 5.5). The average discrepancy is 0.015$a$. This is small compared to the uncertainty in each radius determination and does not affect the results of this work, but for consistency 0.015$a$ was subtracted from all radius values determined using the stream trajectory method. Summarizing, the shadow method was used to determine hotspot location and disc radius for those eclipses where both hotspot ingress and egress are measurable, and the stream trajectory method with mid-egress time was employed for the remaining eclipses (with the -0.015$a$ correction).

The beat phase between orbital and superhump phase, interpreted as disc precession phase $\phi_{prec}$, is defined as zero when superhump maximum and mid-eclipse coincide. The angle $\theta_{disc}$ (see Figure 5.3) is therefore $2\pi \phi_{prec}$.

5.4.4 Eclipse measurements

Table 4 in P2000 gives measurements of the hotspot egress timings for most eclipses analysed here. Measurements of hotspot ingress time and some additional hotspot egress times were measured by eye. The error in each ingress/egress time is taken as the corresponding ingress/egress duration. The additional egress measurements are less accurate than those published in P2000 due to the lower time resolution of these eclipses. This is reflected in the larger error bars associated with these values. All measurements used are given in
Figure 5.5: Disc radius determinations for those eclipses where both ingress and egress of hotspot eclipse were measured. Triangles show radius determinations using the stream method with eclipse egress, while diamonds show the result of using the shadow method. Determinations using the former method are on average 0.015\(a\) larger than using the latter.
5.4 Tracing the eccentric disc shape

Table 5.1: Hotspot time measurements. Ingress and egress times are HJD-2451500. Mid-eclipse times calculated from P2000 ephemeris.

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Table 5.1.

5.4.5 Results

The average disc radius $R_{\text{disc}}$ in units of orbital separation $a$ is $< R_{\text{disc}}/a > = 0.31$ with a spread of 0.07. The disc shape is shown in Figure 5.6 and looks quite circular, but is clearly not centred on the white dwarf. The shape is similar to that found in OY Car by
Figure 5.6: The disc shape traced out from the hotspot ingress and egress times. The disc orientation shown is that which would occur at superhump maximum. Thin regions with large radial extent were determined using the stream trajectory method (Section 5.4.2). The other regions, which are azimuthally extended (due to the relative motion of hotspot and disc frame during hotspot eclipse) were determined using the “shadow” method (Section 5.4.1). The dot-dashed outline is the fitted disc shape and the dotted circle has radius $r_{\text{circ}}$ (the circularization radius for mass ratio $q = 0.14$) and is plotted to make the non-axisymmetry of the disc easier to see.

Hessman et al. (1992), where the stream trajectory method was employed using hotspot eclipse ingresses.

A good fit to the disc shape was achieved using an ellipse with one focus at the white dwarf, semi-major axis $a_{\text{disc}}$ and eccentricity $e$, described by

$$R_{\text{disc}} = \frac{a_{\text{disc}}(1 - e^2)}{1 - e \cos(\alpha_{\text{disc}} - \alpha_0)}.$$ 

$\alpha_0$ is the angle in the precessing disc frame corresponding to maximum disc radius. The resulting fit has $a_{\text{disc}} = 0.33a \pm 0.06a$, $e = 0.29 \pm 0.17$ and $\alpha_0 = 117 \pm 37^\circ$ and is shown in Figure 5.7. The errors in the timing determinations and hence deduced geometry (particu-
5.5 Discussion

The disc radius as a function of $\alpha_{\text{disc}}$ (see definition in Figure 5.3). Points marked with triangles were determined using the "shadow" method (Section 5.4.1), while points marked with stars were determined using the stream trajectory method (Section 5.4.2) and hotspot eclipse egress.

Figure 5.7: The disc radius as a function of $\alpha_{\text{disc}}$ (see definition in Figure 5.3). Points marked with triangles were determined using the "shadow" method (Section 5.4.1), while points marked with stars were determined using the stream trajectory method (Section 5.4.2) and hotspot eclipse egress.

larly when using the stream trajectory method) are rather pessimistic; the determinations of $a_{\text{disc}}$, $e$ and $\alpha_0$ are robust to adopting instead more optimistic error estimates.

The disc is within the tidal cut-off radius $r \sim 0.9R_L = 0.50a$. Within the errors, the largest radius part reaches the location of the 3:1 resonance at $r \sim 0.46a$ (Osaki 1996), providing a plausible explanation for the maintenance of disc eccentricity at this stage after the superoutburst. Since the location of the resonance remains populated, at least where the disc radius is largest, it is to be expected that the disc eccentricity will persist for some time after the superoutburst, i.e. at least $\sim 1$ week in this case. The smallest radius coincides with the circularization radius.

5.5 Discussion

The disc shape in IY UMa is similar to that found for OY Car by Hessman et al. (1992), except that in IY UMa the minimum radius is at disc azimuth $\alpha_{\text{disc}} = -63^\circ$ while in OY Car it is around $\alpha_{\text{disc}} = 0$. This means the smallest radius of the disc is lined up with the donor star at superhump phase -0.18 in IY UMa while this occurs close to
Figure 5.8: (a) The histogram shows the peak bright spot modulation, obtained by subtracting the normalized average orbital curve from the normalized data and then extracting the data for the orbital phase range 0.85–0.95 and folding on superhump phase. (b) shows the measured hotspot intensity $\Delta I_{hse}$ (crosses). The dashed curve in (a) and (b) is the kinetic energy of the stream at the hotspot, while the dotted curve is the kinetic energy of relative motion of the stream and disc dissipated at the hotspot. The solid curves are the predicted hotspot peak and egress intensities using our simple 3D model for the hotspot. See Section 5.5.1. The latter three curves were predicted from the fitted disc shape and scaled in flux to have the same maximum as the measured data. (c) shows the angle $\beta$ between the stream and disc velocities at the impact point.

superhump phase 0 in OY Car. This has implications for the viability of the hotspot as the source of the superhump light during the late superhump era. Superhumps caused by the variation in hotspot brightness as the stream impacts the disc at varying depth in the white dwarf potential well were first suggested by Vogt (1982). This theory has generally been rejected as an explanation for the common superhumps occurring during the decline from superoutburst maximum in SU UMa stars, but the possible link between this model and late superhumps (which appear during very late decline of the superoutburst) has been suggested several times, e.g. by Osaki (1985), Whitehurst (1988b) and by Rolfe, Haswell & Patterson (2000) after a study of persistent superhumps in the nova-like V348
5.5 Discussion

Pup. See Section 3.11 for more discussion.

5.5.1 Hotspot brightness models

The histogram in Figure 5.8a shows the peak bright spot modulation, obtained by subtracting the normalized average orbital curve from the normalized data and then extracting the data for the orbital phase range 0.85–0.95 and folding on superhump phase. The crosses in Figure 5.8b show the hotspot flux uncovered on egress of the hotspot eclipse ($\Delta I_{hse}$ from P2000). The various other curves in Figure 5.8a & b show the hotspot brightness, $F_{hotspot}$, predicted by the disc shape fitted here, calculated using three models described below.

*The gravitational potential model (Model (i))*

The dashed curve in Figure 5.8a & b is calculated using $F_{hotspot} \propto 1/r$ where $r$ is the radius of the disc at the hotspot. This assumes the energy released at the hotspot varies as its depth in the white dwarf gravitational potential as first suggested in Vogt (1982), effectively considering the kinetic energy of the stream to be the gravitational potential energy released.

The phasing of the hotspot brightness predicted by this model does not agree with either of the two measures of hotspot variation in Figures 5.8a & b, with the peak in predicted brightness being about 0.25 earlier than the measured peak. The fractional amplitude in the predicted variation is about a third of that of the measured variation: in Figures 5.8a & b the predicted curve has simply been scaled to have the same maximum value as the observed hotspot curves.

Taking the stream kinetic energy at impact to be $\propto 1/r$ neglects the gravitational potential and the kinetic energy at the L1 point, where the stream starts. Taking the velocity at L1 to be the photospheric sound speed $C_s \approx 5 \, \text{km} \, \text{s}^{-1}$ (Equation 4.5), we see that this is negligible compared to the typical velocity at the disc impact of $\sim 800 \, \text{km} \, \text{s}^{-1}$. However, the potential at L1 is not negligible compared to that at the disc edge. Using correctly calculated stream trajectories to determine the stream velocity at the impact roughly doubles the fractional amplitude improving agreement with the observed hotspot changes. Crucially, however, the phasing of the predicted hot spot brightness is unchanged, so this model still poorly describes the hotspot brightness.

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1Here fractional amplitude is defined as the ratio (maximum flux - minimum flux)/average flux.
5.5 Discussion

The available kinetic energy model (Model (ii))

In Figure 5.8c we also show the variation of the angle $\beta$ between the stream and disc velocities, $\vec{V}_{\text{stream}}$ and $\vec{V}_{\text{disc}}$ (described below), at the point of impact (approximately equal to $\beta$ in Figure 5.3). The hotspot emission should arise primarily from dissipation of kinetic energy as the infalling stream impacts the disc edge and merges with the Keplerian motion of the disc. The available kinetic energy increases as the angle $\beta$ between the velocity vectors of the stream and disc flows increases. Figure 5.8c clearly shows $\beta$ to be greatest around the observed maximum of the hotspot intensity, which coincides with superhump maximum.

This effect is taken into account in this model, where the maximum kinetic energy released at the impact point is calculated and depends on the directions and magnitudes of the stream and disc velocities. For this highly simplified model, the stream-disc impact is assumed to occur at a single point, and we consider the behaviour at any time in an instantaneous inertial frame moving with the disc flow velocity at the impact point. In this frame, the stream has velocity $\vec{V}_{\text{rel}} = \vec{V}_{\text{stream}} - \vec{V}_{\text{disc}}$, which can be resolved parallel and perpendicular to the direction of disc flow at the impact giving $V_{\text{rel}||}$ and $V_{\text{rel}\perp}$ respectively. Since the stream material must have the same velocity as the disc after the impact, the kinetic energy associated with stream flow perpendicular to the disc must be dissipated at the hotspot at a rate $\frac{1}{2} \dot{M} V_{\text{rel}\perp}^2$, where $\dot{M}$ is the mass flow rate of the stream. The case is slightly more complicated when considering the stream flow parallel to the disc flow. If $V_{\text{rel}||} > 0$, i.e. the stream is moving faster than the disc in this direction, then the excess kinetic energy must also be dissipated at rate $\frac{1}{2} \dot{M} V_{\text{rel}||}^2$. If, however $V_{\text{rel}||} < 0$, which is usually the case, then the stream flow must be accelerated in the direction of disc flow, and hence kinetic energy must be supplied by the disc at a rate $\frac{1}{2} \dot{M} V_{\text{rel}||}^2$. Thus the net rate of dissipation of energy at the hotspot in this simple model is

$$E = \frac{1}{2} \dot{M} V_{\text{rel}||}^2 + \frac{1}{2} \dot{M} V_{\text{rel}||}^2 \left| \frac{V_{\text{rel}||}}{V_{\text{rel}||}} \right|. \quad (5.1)$$

Note that in Rolfe, Haswell & Patterson (2001) the difference between cases for which $V_{\text{rel}||} > 0$ and $V_{\text{rel}||} < 0$ was not considered. Therefore the model described here is more realistic, but since in most cases $V_{\text{rel}\perp}^2$ dominates $V_{\text{rel}||}^2$, there is no significant difference in predicted results.
The disc velocity is approximated

\[ |\vec{V}_{\text{disc}}| = \sqrt{GM_{\text{wd}} \left( \frac{2}{r} - \frac{1}{a_{\text{disc}}} \right)} \]

at the impact point with direction parallel to the disc edge in the white dwarf frame. This is the correct velocity for an elliptical orbit around a point mass. Putting all these ingredients together, we use Equation 5.1 to calculate the energy dissipated at the hotspot. This model is plotted as the dotted curve in Figures 5.8a & b.

The shape, phasing and fractional amplitude of the curve for this model shows much better agreement with the hotspot brightness measured in Figure 5.8, although the predicted peak is still a little early. The predicted curve also agrees similarly well in phase and fractional amplitude with the measure of hotspot flux in Figure 5.8b, although the predicted rise in hotspot flux is too early, making the peak too wide.

It should be pointed out that, assuming elliptical orbits of material in the outer disc, Faulkner found analytically that the energy released at the stream-disc impact should be maximum when the stream hits the leading edge of the disc on the latus rectum closest to periastron (B. Warner, private communication). This prediction agrees almost exactly with the result found here for the disc orientation at late superhump maximum if the late superhump emission comes from the energy released at the hotspot.

The available kinetic energy model, including hotspot geometry (Model (iii))

The solid curves in Figures 5.8a & b were produced assuming the hotspot brightness behaves as described in the previous model, but with the additional consideration of the three dimensional structure of the hotspot. A simple 3D structure was assumed, shown in Figure 5.9. The hotspot has an elliptical cross section in the \( r - z \) plane with axis of size \( r_{\text{spot}} \) in the radial direction and \( h_{\text{spot}} \) in the vertical direction (see inset in Figure 5.9). \( r_{\text{spot}} \) and \( h_{\text{spot}} \) and the spot brightness decrease downstream from the initial impact as \( e^{-\theta^2/\Delta\theta_{\text{spot}}^2} \). Upstream of the impact the hotspot surface is rounded off with a hemisphere of uniform brightness equal to that at the initial impact. The angular extent of the hotspot region is set by \( \Delta\theta_{\text{spot}} = \Delta\theta_0 a_{\text{disc}}/r_0 \), where \( r_0 \) is the disc radius at the impact point. This keeps the arc length of the hotspot region roughly constant as the eccentric disc precesses. The disc thickness is \( 2H_{\text{disc}} \) and any region of the hotspot surface within\(^2\) the disc is considered to be obscured completely. The total flux coming from an area element on

\(^2\)A region is within the disc if it lies within the elliptical disc shape and has height \( z \) above the orbital plane satisfying \( |z| < H_{\text{disc}} \).
the hotspot surface is foreshortened according to its area projected in the direction of the observer. This simple parametrized model for the hotspot structure makes it possible to model the full hotspot lightcurve, taking into account the visibility of the hotspot at any orbital phase and disc precession phase. The model lightcurve is shown in Figure 5.11. $\Delta\theta_0 = 36^\circ$ is used, which is the average azimuthal extent of the hotspot regions determined from the eclipse timings as described in Section 5.4.1. Noting that no eclipse of the white dwarf by the hotspot region is seen, the maximum possible value of $h_{\text{spot}}$ for $i = 86.78 \pm 1.5$ is $0.013a \pm 0.006a$. Therefore $r_{\text{spot}}(\theta = 0) = h_{\text{spot}}(\theta = 0) = 0.013a$ was adopted for the point of impact ($\theta=0$). A value of $H_{\text{disc}} = 0.009a$ was used, adjusted to achieve a good match between the simulated and observed lightcurves discussed below.

This model does not differ from the previous model in its prediction for the peak flux
of the orbital hump (Figure 5.8a). This is to be expected since at the peak of the orbital hump we should be seeing the whole length of the bright spot clearly, so the visibility at this orbital phase should not be very sensitive to small changes in orientation through the disc precession cycle. The hotspot flux measured using the hotspot eclipse egress is different from the predicted variation in intrinsic hotspot flux. This is because at the orbital phase of hotspot egress, we are seeing the impact region roughly end-on, and at this orientation the fraction of the hotspot flux reaching the observer is sensitive to the exact orientation of the hotspot, which of course varies with $\phi_{\text{prec}}$. The result is that the predicted variation (solid curve in Figure 5.8b) rises to maximum later than the intrinsic variation, but decreases at the same time as the intrinsic variation. This produces a more flat-bottomed curve which agrees better in shape with the observed egress flux than models (i) and (ii). The flux around egress flux minimum is too low, but this is not surprising since this model for the structure and behaviour of the impact region is a very simple one. It is also possible that the hotspot emission is strongly anisotropic due to the complicated structure of shocks and contact discontinuities expected to be present at the impact.

The late superhump and hotspot lightcurve

Figure 5.10 shows the evolution of the IY UMa lightcurve throughout the late superhump era. The important features to note are how strongly the amplitude of the orbital hump varies, and how the late superhump is very weak, in many cases undetectable, when it does not occur near the orbital hump. The combined lightcurve of the late superhump and the orbital hump is not simply a sum of the two, it is more like a product of two modulations, with the combined hump being extremely strong when orbital hump and late superhump coincide, while both orbital hump and late superhump are weak when they are not coincident. This is exactly what we expect if the stream-disc impact is the source of the late superhump. The hot spot is modulated at the late superhump period, as the impact geometry of the stream and disc varies. Consequently, the orbital hump, which is just the emission from the hotspot coming into view, is modulated on the disc precession period, being strongest at late superhump maximum. The late superhump is very weak away from the orbital hump since away from the orbital hump the bright spot is on the far side of the disc and hence has a low visibility. Model hotspot lightcurves produced using model (iii) are shown in Figure 5.11. Ignoring the stochastic flickering in the observations (which is strongest in the second panel, HJD 72.7), and the deep white dwarf eclipse in
Figure 5.10: All 7 lightcurves from the late superhump dataset analysed in this paper. The date of the start of each lightcurve is shown in the top left of each plot as HJD-2451500. Arrows indicate times of late superhump maximum according to the ephemeris in P2000. Note the significant variation in the amplitude of the orbital hump over the 7 days.
Figure 5.11: The model bright spot lightcurve for the same time series as Figure 5.10. The model assumes the intrinsic hotspot brightness to vary as the kinetic energy of relative motion of stream and disc dissipated at impact, while a simple model for the 3D structure of the hotspot enables us to predict its visibility as a function of orbital phase and precession phase. See Section 5.5.1. Arrows as in Figure 5.10.
5.5 Discussion

the observations not included in the model, this simple model does a convincing job of reproducing the observed bright spot behaviour, although the peak of the orbital hump when it occurs close to late superhump maximum is sharper in the observations than in our model.

The intrinsic bright spot flux in the lightcurve in the third panel (HJD 73.7) appears to be lower than the observed flux, which makes it difficult to see how the shape of this curve agrees with observations. Figure 5.12 shows the model and observations for this curve, with the model scaled by a factor of 2. The good agreement between the shape of the model and the observations is clear, particularly the double humped nature of each orbit curve and the flattish top to the orbital hump before some of the eclipses. More detailed modeling is necessary to explain the low intrinsic brightness of the hotspot in this region of the disc precession cycle. The relative height of the hotspot peak above the disc needed a bit of fine-tuning to achieve this match between the shapes of the model and observed lightcurves, with $r_{\text{spot}}(\theta = 0)$ having to be less than about twice $h_{\text{spot}}(\theta = 0)$.

Figure 5.12: The model data from the 3rd panel of Figure 5.11 scaled in flux by a factor of 2 and shifted up by 1.4 flux units, along with the corresponding observed data. This shows how the shape of the model lightcurve is morphologically close to that of the observations, although the intrinsic hotspot brightness predicted by the model is too low at the disc precession phase.

This work assumes the shape of the disc to be fixed and constant in the precessing disc frame, while the observations of V348 Pup (Chapter 3) and SPH simulations e.g. Haswell et al. (2001) suggest the disc shape and size changes during the superhump cycle. The disc shape found here represents the disc radius encountered by the accretion stream as
a function of superhump phase, and any effect of changing shape should have little effect on the angle $\beta$ at which the stream impacts the disc since the shape is traced out between radius measurements which are closely spaced in superhump phase.

5.6 Conclusions

It is difficult to avoid the conclusion that the late superhump light in IY UMa is predominantly coming from the hotspot, whose brightness is modulated at the superhump period as a result of the varying impact geometry of the stream and the disc. The disc shape and orientation found in OY Car (Hessman et al. 1992) suggests that the late superhumps in that system may also originate from the hotspot. An analysis applying this modified hotspot model to the OY Car observations would be valuable. Murray (1996) performed SPH simulations which support the hotspot as the source of late superhump light, but radiative processes were not modeled. More detailed hydrodynamic simulations concentrating on the stream-disc impact region have been carried out (e.g. Armitage & Livio (1998)), but there has been no study of the hotspot in such detail for non-circular discs. Such work would be valuable in the light of these observations.

As originally proposed by Osaki (1985), the transition from the common superhumps which appear during the decline from maximum of the superoutburst to the late superhumps during the very late decline of the superoutburst is easy to explain if the common superhumps result from viscous dissipation as the disc is tidally stressed and the late superhumps arise from the varying hotspot brightness. The disc drops out of the hot outburst state at the end of the superoutburst while remaining tidally eccentric. At this point, the disc viscosity dramatically reduces and so the viscous dissipation due to tidal stressing which causes common superhumps will be similarly reduced, allowing the late superhumps which were drowned out during the superoutburst to be seen. In some systems the relative strength of the late superhumps and common superhumps is such that both are seen simultaneously towards the end of the superoutburst, e.g. in VW Hyi (Schoembs & Vogt 1980) and probably also V1159 Ori (Patterson et al. 1995).
Chapter 6

Red spectroscopy of IY UMa: The search for the donor

6.1 Introduction

Chapter 5 introduced us to the recently identified high inclination CV IY UMa through a photometric study of numerous eclipses, probing the disc shape and behaviour of the stream-disc impact as it approached quiescence after a superoutburst in January 2000. In this chapter we analyse time-resolved red spectroscopy of IY UMa, searching for evidence of the donor star.

The orbital parameters of IY UMa were estimated in Patterson et al. (2000) (hereafter P2000). These estimates result from combining various measurements of hotspot and white dwarf eclipse timing with theoretical models for white dwarf and donor star luminosities, radii and masses. It is important to employ direct methods to measure the orbital parameters wherever possible, avoiding reliance on uncertain models and providing the opportunity to test such models. A complete solution for the orbital parameters of a CV can be obtained if the orbital period, the orbital radial velocity amplitudes of the white dwarf and donor star and the white dwarf eclipse width can be measured. The measurement of white dwarf eclipse duration fully constrains orbital inclination as a function of mass ratio (as in Section 3.5), relying only on the assumption of a Roche lobe filling donor star. Thus we have $i(q)$. The observed donor velocity amplitude, $K_2$, is

$$K_2 = \frac{2\pi a \sin i}{(1 + q)P_{\text{orb}}}. $$

Combining this with Kepler's 3rd law (Equation 1.1) gives the mass function (Frank,
6.2 Observations

King & Raine 1992),

\[ f(M_1, q) = \frac{M_1 \sin^3 i(q)}{(1 + q)^2} = \frac{P_{\text{orb}} K_2^3}{2\pi G}. \]

\( f(M_1, q) \) is useful, because it provides a lower limit for compact object mass, \( M_1 \), since \( \sin^3 i \leq 1 \) and \( q > 0 \), and it can be determined by measuring \( P_{\text{orb}} \) and \( K_2 \). Knowing \( P_{\text{orb}} \) and \( i(q) \) for IY UMa from the white dwarf eclipse measurements in P2000, measurements of \( K_1 \) (the white dwarf velocity) and \( K_2 \) would completely constrain the system parameters, \( M_1, M_2, i \) and the orbital separation.

As explained above, measurement of the donor star velocity amplitude, \( K_2 \), provides an important constraint on the system parameters. A fit of donor spectral type to the orbital period of CVs (Equation 4 in Smith & Dhillon (1998)) suggests a spectral type of M4.2 to M6.3 for the donor in IY UMa. Such stars display various absorption features in the far optical and near infrared which have been used before to identify the spectral type of donors in CVs and measure \( K_2 \). The donors in the short period dwarf novae Z Cha and HT Cas were identified and measured this way (Wade & Horne (1988) and Marsh (1990)). A programme of observations was therefore carried out to search for donor star absorption features in the wavelength range 7000Å–8300Å.

6.2 Observations

The observations were taken with Dr. Tim Abbott during a five half-night observing run at the Nordic Optical Telescope in La Palma. The red/near-infrared observations discussed in this section were taken on the 3rd and 4th January 2001, with a few calibration images taken on the 7th. ALFOSC, the Andalucia Faint Object Spectrograph and Camera, was used with slit width 1.2" and grism 8 to cover wavelength range 5810Å to 8350Å. The spectra have a dispersion of 1.24Å per pixel and a FWHM resolution of about 5.2 pix (6.2Å). In this chapter, only the red end of the spectra, with \( \lambda \) greater than 7000Å, was used. The blue end of these spectra is discussed in the next chapter. 39 10-minute exposures of IY UMa and a nearby comparison were obtained, along with exposures of several M type dwarfs for use as template stars (listed in Table 6.1). Helium and Neon arc lamp exposures were taken before every change of target and regularly when tracking IY UMa for long periods. Observations of the flux standard HD 93521 were taken which also served as a telluric standard. Halogen lamp flats and sky flats were also taken.
6.3 Data reduction

Bias subtraction was carried out using CCDPROC using combined bias frames produced separately for each night. To reduce readout time only a 98×2048 pixel strip of the CCD containing the object and comparison spectra was read out. This meant that no overscan correction was applied. In the red the CCD used in ALFOSC suffers from fringing. This results from interference between multiple reflections in the CCD of the 0.8 μm light. This is illustrated in Figure 6.1, showing the combined normalized lamp flat from 3rd January and a cut in the dispersion direction showing column 50. The fringes can clearly be seen getting stronger towards the reddest end of the CCD, reaching a maximum amplitude of about 25 per cent. The structure of the fringe pattern might depend on how the slit is illuminated as well as the precise dispersion solution for each image. Despite the small changes in the dispersion solution throughout the run and the fact that the lamp is an extended source while the objects are point sources, using the normalized lamp flats to flatfield the images removed all detectable traces of the fringing pattern (see e.g. Figure 6.4). The sky flats were used to produce an illumination correction. Wavelength calibration was carried out using the He arc lamp exposures.

A well exposed spectrum of HD 93521 was normalized, and the atmospheric absorption (telluric) features used to produce a template spectrum (Figure 6.2). The O9Vp star HD 93521 was used as a telluric standard since its spectrum contains few features, enabling easy identification of the atmospheric absorptions. This template was then used to correct for telluric features using the TELLURIC package in IRAF, which divides each spectrum by the template spectrum, scaled according to the ratio of airmasses in the template and object spectrum. Figure 6.2 shows the residuals after correcting 5 standard star exposures.

Table 6.1: M dwarf template stars

<table>
<thead>
<tr>
<th>Object</th>
<th>Spectral type</th>
<th>Reference</th>
</tr>
</thead>
<tbody>
<tr>
<td>GJ 1154, Zkh 176</td>
<td>M5V</td>
<td>Zakhozhaj (1979)</td>
</tr>
<tr>
<td>GJ 228B, Zkh 84</td>
<td>M4V</td>
<td>Zakhozhaj (1979)</td>
</tr>
<tr>
<td>BD+28 2110</td>
<td>M3V</td>
<td>Upgren (1962)</td>
</tr>
<tr>
<td>1E 1548.7+2009</td>
<td>M2Ve</td>
<td>Stocke et al. (1991)</td>
</tr>
<tr>
<td>CTI 092053.7+280101</td>
<td>M1.5V</td>
<td>Kirkpatrick et al. (1994)</td>
</tr>
</tbody>
</table>
(shifted up by 1.2 in normalized flux units). Apart from the strongest feature around 7600 Å, the correction has worked quite well.
Figure 6.3: Average red spectra of IY UMa on 3rd & 4th January 2001. (a) shows average in rest frame of white dwarf, (b) in rest frame of donor, (c) in frame of observer. (d) shows average when donor eclipses disc and white dwarf, (e) when disc obscures donor. Shaded regions show error bars.

Unfortunately conditions were not photometric, and so no wide slit exposures of IY UMa and comparison or of the flux standard were taken using this red grism. Therefore, flux calibration of these red spectra was not possible. However, the skew mapping technique described below relies on normalized spectra and so this lack of flux calibration is not a serious impediment to this work. Each spectrum was normalized by fitting a 4th order polynomial to the continuum regions and dividing by this. This was done automatically using IDL.

6.4 Average spectra

Various different averages of the IY UMa spectra are shown in Figure 6.3 while the M dwarf spectra are shown in Figure 6.4. Each plotted spectrum has been gaussian smoothed with
FWMH of 4 pixels which is less than the instrumental FWHM resolution, and shifted in flux by a multiple of 0.5 units (Figure 6.3) or 0.8 units (Figure 6.4).

Figure 6.3c shows the average of all 39 IY UMa spectra. The most obvious features are the He\textsc{i} emission and core absorption at 7065\AA, the 7773\AA\ O\textsc{i} absorption, the broad absorption band around 7600\AA\ to 8200\AA\ and the absorption line around 8200\AA. The He\textsc{i} is from the disc, and is discussed in Chapter 7.

The broad 7600\AA–8200\AA\ feature matches the TiO absorption band seen in the M dwarf spectra (Figure 6.4), where it is strongest in the M5V dwarf. Figures 6.3d & e show average spectra of IY UMa during eclipse (orbital phases -0.1-0.1) and where the donor should be partially obscured by the disc (phase 0.42-0.58)\textsuperscript{1}. The absorption is clearly much stronger (by a factor \~2) when the donor eclipses the disc than when the disc obscures the donor, confirming the identification of this feature as the TiO band absorption in the donor.

\textsuperscript{1}Using maximum entropy eclipse mapping Stanishev et al. (2001) measured the disc radius in IY UMa to be 0.25\textit{a} during quiescence. A disc of this radius obscures some part of the donor between phases 0.42 and 0.58.
The 7773Å O I triplet has been seen in many CVs. A study of 65 CVs by Friend et al. (1988) reveals this feature in absorption in several systems during outburst, but in emission during quiescence. HT Cas also shows this feature in emission during quiescence (Marsh 1990). Friend et al. (1988) suggests that we are seeing line emission from an optically thin disc in quiescence, and absorption from an optically thick disc in outburst. Two high inclination systems, Z Cha (Wade & Horne 1988) and V2051 Oph (Friend et al. 1988), show a broad emission from the O I triplet, superimposed with a narrower absorption component. It is possible that a broad weak emission component is also present in these IY UMa spectra. Friend et al. (1988) suggests that the narrow absorption feature is absorption of white dwarf emission by the disc, something which we might expect in high inclination systems. That this feature arises from the disc and certainly not from the donor is confirmed by Figures 6.3a & b, where each individual spectrum was Doppler-shifted into the rest frame of the white dwarf or donor before adding together the spectra. The velocities of the donor and white dwarf were assumed to be $K_2 = 442\ \text{km s}^{-1}$ and $K_1 = 62\ \text{km s}^{-1}$ respectively, calculated using the P2000 orbital parameters. In the rest frame of the white dwarf the O I feature remains sharp, while in the donor star frame it splits into two absorption dips, telling us that the donor is not the source of this feature. In addition, the depth of this absorption is significantly stronger when the disc obscures the donor (Figure 6.3e) than when the donor eclipses the disc (Figure 6.3d).

The 8200Å absorption feature closely matches the 8183Å and 8194Å Na I doublet seen in all five dwarf template spectra. In the average spectrum (Figure 6.3c) it does not have the same doublet structure as seen in the template spectra, but when shifted to the donor rest frame (Figure 6.3b) the doublet structure becomes apparent, clearly matching the laboratory wavelengths. This feature is also more significant (considering the flux errors) during disc eclipse (Figure 6.3d) than when the disc passes in front of the donor (Figure 6.3e). This is a clear detection of the Na I doublet, a feature which has been exploited many times to measure $K_2$ in CVs, e.g. in Z Cha (Wade & Horne 1988) and several other systems (e.g. Friend et al. (1990a); Friend et al. (1990b); Friend et al. (1988); Marsh (1990)).

The TiO absorption band around 7000Å-7400Å seen in the M dwarfs also appears to be present, if weak, in IY UMa, and is again more prominent when the donor obscures the disc than when the disc obscures the donor. The K I lines seen at 7665Å and 7699Å are not clearly seen, although there are weak features close to the noise level which could
6.5 Skew mapping

be tentatively identified as the K I lines.

Weaker Na I and TiO absorption around phase 0.5 than at other phases has been seen before in dwarf novae, e.g. Z Cha (Wade & Horne 1988), HT Cas (Marsh 1990) and SS Cyg (Hessman et al. 1984), and was attributed to a non-uniform distribution of absorption across the surface of the donor. The suggested cause of this non-uniform distribution is the filling-in of absorption features on the inner face of the donor by irradiation from the disc and/or white dwarf. This might also be the cause of the weakened features around phase 0.5 in IY UMa, rather than eclipse of the donor by the disc.

The absorption feature at 7600 Å is probably a combination of residual atmospheric absorption and the 7600 Å feature seen in M dwarf spectra. The M2Ve dwarf spectrum shows evidence of over-correction of the telluric absorption around 7600 Å and 8230 Å. There is also some evidence of over-correction in the M5V template around 8230 Å.

We see no evidence of Paschen emission around 8200 Å seen in some CVs e.g. TT Ari and V603 Aql (Friend et al. 1988).

This simple analysis clearly shows that the donor star in IY UMa has been detected, showing all the expected features of a late M dwarf star. Without flux calibrated spectra, the relative strengths of the absorption features cannot be used to estimate more precisely the spectral type of the donor. The low signal-to-noise ratio prevents us from exploiting the Na I doublet to measure \( K_2 \) directly from the spectra, but the skew mapping technique (Smith, Cameron & Tucknott 1993) has been exploited before to locate the donor in velocity space when the S/N is low.

6.5 Skew mapping

Skew mapping, described in Smith, Cameron & Tucknott (1993), attempts to detect donor star absorption features, such as those described above, and determine the location of the donor in velocity space. This technique has been successfully employed for several systems (e.g. Smith, Cameron & Tucknott (1993), Smith, Dhillon & Marsh (1998)) and recently in VY Aqr (Littlefair et al. 2000). It is the best method for detecting the donor and determining its velocity when the absorption features are weak. This is because it co-adds the spectra to increase signal-to-noise, while also correcting each spectrum for the orbital motion of the donor, to avoid smearing out donor absorption features. This is exactly how Figure 6.3b was produced, using an assumed donor star velocity, which revealed the Na I absorption doublet structure. Skew mapping extends this technique, by
cross-correlating each such co-added spectrum with a template donor star spectrum, and
finding the combination of template star and donor velocity at which the cross-correlation
peak is strongest. This process is equivalent to cross-correlating the template spectrum
individually with each object spectrum to produce a trailed “cross-correlation spectrum”,
and then back-projecting this trail into velocity space (as in Doppler tomography (see
Section 2.2.2)), producing a cross-correlation map in velocity-space (the skew map) which
should show a peak at the velocity of the donor star. This approach was used to examine
the donor star absorption in IY UMa.

First, the normalized spectra of the template M dwarfs and IY UMa were continuum
subtracted (i.e. had a value 1 subtracted). They were then gaussian smoothed with
FWHM of 2.5 pixels to increase S/N without losing any spectral information (the instru­
mental resolution was \( \sim 5.2 \) pixels). Each IY UMa spectrum, \( F(\lambda) \), was cross correlated
with the template spectrum, \( T(\lambda) \), for a range of velocity shifts, \( v \), using correlation coeffi­
cient \( C_v \) defined as

\[
C_v = \frac{\sum_{\lambda=\lambda_{\text{min}}}^{\lambda_{\text{max}}} F(\lambda)T'(\lambda)}{\left(\sum_{\lambda=\lambda_{\text{min}}}^{\lambda_{\text{max}}} F(\lambda)^2 \sum_{\lambda=\lambda_{\text{min}}}^{\lambda_{\text{max}}} T'(\lambda)^2\right)^{\frac{1}{2}}}
\]

(6.1)

where \( T'(\lambda) = T(\frac{\lambda}{1+v/c}) \). This definition of \( T'(\lambda) \) simply has the effect of red-shifting
the template spectrum by velocity \( v \) before cross-correlating. The denominator of \( C_v \) is
chosen so that if the template and donor spectra are identical, the value of \( C_v \) at the peak
will be 1, independent of any simple scaling of the flux of either spectrum. The correlation
trails thus produced were then transformed to velocity-space using the Fourier-filtered
back-projection technique (Section 2.2.2) in MOLLY. The strength of the skew map at a
given velocity therefore shows how well the object spectra match a star of the template’s
spectral type orbiting at that velocity.

Only the NaI doublet wavelength range was used in the cross-correlation, since this
is the only sharp donor absorption feature detected, and will therefore provide all the
velocity information. Skew maps were also produced using the spectral range \( \lambda =7045–
8240\AA \) (with the O I and He I lines masked out) which uses all the detected donor star
features, but this simply smears out the donor star in the skew map, probably a result
of the broad TiO absorptions showing a strong correlation but low velocity-sensitivity.
The normalization of the spectrum around the NaI doublet was improved by fitting a
straight line to the continuum either side of the doublet. The skew map also depends upon
the relative systemic velocity of IY UMa and the template star. Neither of these is yet
known, although measurements of quiescent Hα emission lead to an estimate for IY UMa of $\gamma = 13.6 \pm 1 \text{ km s}^{-1}$ (see Section 7.5). By cross-correlating the M3V template whose systemic velocity is 1 km s$^{-1}$ (Evans 1967) with the other four templates, the systemic velocities of these stars were taken into account. The skew maps for each template were produced for a range of values$^2$ of the IY UMa systemic velocity, $\gamma_{\text{YUMa}}$, between -100 and 300 km s$^{-1}$, enabling a determination of the systemic velocity which produces the clearest peak to be made, necessary because the low resolution of our spectra limits the accuracy of our systemic velocity determinations. The values of $\gamma_{\text{YUMa}}$ discussed below are not determinations of the systemic velocity of IY UMa: the low spectral resolution and S/N of the observations prevents such a measurement being made.

The orbital phase range 0.42 to 0.58 was not included when producing the back projections to avoid the weakened/obscured absorption lines around this phase interfering with the results. The results from using the entire phase range are considered at the end of this section.

Figure 6.5: Cross-correlation trails for each template

$^2$This was identified as a suitable range for $\gamma_{\text{YUMa}}$ after initially studying the range -750-750 km s$^{-1}$. 
Figure 6.5 shows the correlation trails for each template. These trails were folded and binned into 20 orbital phase bins, and into 100 velocity bins in the range -4000 km s\(^{-1}\) to 4000 km s\(^{-1}\). All five trails look almost identical, which is unsurprising since at this low resolution and signal-to-noise, differences in structure in the NaI doublet between the spectral types are not going to be distinguishable in the IY UMa observations. The trails show an S-wave, corresponding to a localized region in velocity space at which the IY UMa spectrum best matches the sodium feature. The peak-peak amplitude of this feature is about 1000 km s\(^{-1}\), and is at minimum excursion around orbital phase 0. This phasing and amplitude is in agreement with the predicted orbital motion of the donor (plotted in black in Figure 6.5) using the P2000 orbital parameters: line-of-sight donor velocity semi-amplitude 442 km s\(^{-1}\). The velocities are heliocentric and corrected for the template velocities measured as described above, so that the S-wave should be centred at the systemic velocity of IY UMa.

These trails were transformed to velocity space as described. The resulting skew maps (apart from the left hand M2Ve map in Figure 6.6, discussed below) have a clear maximum close to the predicted location of the donor star using the P2000 parameters. The exact location of the maximum depends on the assumed IY UMa systemic velocity, \(\gamma_{\text{IY UMa}}\). Two skew maps for each template are shown in Figure 6.6. Those in the left column use the value of \(\gamma_{\text{IY UMa}}\) which produces the strongest peak (i.e. max correlation), while those on the right correspond to the \(\gamma_{\text{IY UMa}}\) which produces a peak with no x-component of velocity (as expected for the donor). These values of \(\gamma_{\text{IY UMa}}\) were determined as described below.

The velocity and width of the peak was determined for each image by fitting a 2-dimensional gaussian of the form

\[
C_{\tilde{y}} = C_0 + C_1 \exp \left( -\frac{(v_x - v_{x,\text{max}})^2}{2\Delta v_x^2} - \frac{(v_y - v_{y,\text{max}})^2}{2\Delta v_y^2} \right)
\]

where \(C_{\tilde{y}}\) is the value of the skew map at velocity \((v_x, v_y)\). Figure 6.7 shows the variation in position and shape of the skew map peak as a function of \(\gamma_{\text{IY UMa}}\). The phase shift is the angular offset of the peak from \(v_x=0\) measured about the centre of mass, i.e. \(\frac{1}{2\pi} \arctan \frac{v_{x,\text{max}}}{v_{y,\text{max}}}\). The velocity amplitude is \(V = \sqrt{v_{x,\text{max}}^2 + v_{y,\text{max}}^2}\). The peak width is measured using \(\Delta v_r = \sqrt{\Delta v_x^2 + \Delta v_y^2}\). The correlation of the peak is taken as \(C_0 + C_1\).

As \(\gamma_{\text{IY UMa}}\) increases from -100 km s\(^{-1}\) to 300 km s\(^{-1}\), the peak moves from a velocity of about 600–700 km s\(^{-1}\) and phase shift of 0.01 through the expected donor location to a lower velocity around 350–450 km s\(^{-1}\). This result is true for all five templates. Only
Figure 6.6: IY UMa skew maps excluding phase range 0.42–0.58. Black markings are as described in Chapter 2.2.2. White cross is measured peak maximum.
the peak size \( \Delta v_r \) and the peak correlation vary significantly depending on the template used. The difference in \( \Delta v_r \) between templates should not be taken as a measure of how tightly each template constrains the location of the donor: it is more likely to result from differing seeing between the different template observations. However, the variation of \( \Delta v_r \) with \( \gamma_{\text{YUMa}} \) for each template is more useful, telling us which value of the IY UMa systemic velocity gives us the sharpest peak, and is therefore the most likely value for the systemic velocity. This shows the peak to be narrowest for \( \gamma_{\text{YUMa}} \) around 75 to 125 km s\(^{-1}\). The most important measure of which value of \( \gamma_{\text{YUMa}} \) is correct is that for which the peak correlation is strongest. This is around \( \gamma_{\text{YUMa}} = 90 \) to 150 km s\(^{-1}\) for the M1.5V, M3V, M4V and M5V templates. For the M2Ve template the correlation rises steadily as the systemic velocity increases. Studying higher values of \( \gamma_{\text{YUMa}} \) reveals a peak for the M2Ve template around 300 km s\(^{-1}\), but this corresponds to a large phase

Figure 6.7: Comparison of skew maps (excluding phase range 0.42-0.58) for each template.
shift of about 0.11, placing the peak a long way from the donor velocity (as can be seen
the left hand M2Ve map in Figure 6.6). The correlation for this template is lower at
all values of \( \gamma_{YUMa} \) than for the other templates, telling us that this template spectrum
provides the poorest match with IY UMa. The correlation of the remaining four templates
is similar, which considering the low signal-noise and resolution of these observations is
not surprising: detailed differences in the Na I doublet between the spectral types have
not been resolved.

The value of \( \gamma_{YUMa} \) with the highest peak correlation gives us one measure of the best
fitting skew map. These skew maps are shown in the left hand column of Figure 6.6. The
average phase shift for all five of these maps is 0.038, with a spread in values of 0.001. This
shift is much greater than accumulated error in the P2000 ephemeris at the time of these
observations, which is only 0.005 orbits. The 360 second exposure time of the IY UMa
spectra corresponds to a phase resolution of 0.06, so this phase shift is consistent with the
peak actually lying along the line of centres of the two stars, where we expect the donor.
This then gives us a donor velocity of \( K_2 = 412 \text{ km s}^{-1} \) with a spread of 5 \( \text{km s}^{-1} \). If this
measured phase shift is real, then we should also consider values of \( \gamma_{YUMa} \) for which the
phase shift is zero, placing the donor along the line of centres. The corresponding skew
maps are shown in the right hand column of Figure 6.6, and correspond to \( K_2 = 545 \text{ km s}^{-1} \)
with a spread of 11 \( \text{km s}^{-1} \).

As described earlier, these skew maps were produced omitting phase range 0.42–0.58.
As a check, the analysis was repeated using the entire phase range. Figures 6.8 and 6.9
show the skew maps and comparison of the peak as a function of \( \gamma_{YUMa} \) as before. The
peak in correlation, the minimum value of \( \Delta v_r \), and phase shift of zero occur for values of
\( \gamma_{YUMa} \) about 40–50 \( \text{km s}^{-1} \) greater than before. The phase shift at maximum correlation
is now 0.059 with a spread of 0.002, as large as the phase resolution of the data. The orbital
velocity of the donor for maximum correlation is now \( K_2 = 390 \text{ km s}^{-1} \) with a spread of
3 \( \text{km s}^{-1} \). The skew maps on the right of Figure 6.8, corresponding to phase shift zero,
show far less distinct peaks than in Figure 6.6. They yield a value of \( K_2 = 467 \text{ km s}^{-1} \)
with a spread of 17 \( \text{km s}^{-1} \), which is much smaller than the 545 \( \text{km s}^{-1} \) from the maps
excluding phases around 0.5, but closer to the 442 \( \text{km s}^{-1} \) predicted by the P2000 orbital
parameters. This is probably because the maps now include more absorption from the
inner face of the donor, where the orbital velocity is lower. However, the far messier peak
in the phase shift zero maps renders this measurement less reliable.
Figure 6.8: IY UMa skew maps covering entire phase range. Black markings are as described in Chapter 2.2.2. White cross is measured peak maximum.
6.6 Conclusions

The donor star has been unambiguously detected in IY UMa. It is a late type M dwarf, but the low signal-to-noise and resolution of these observations prevents an accurate determination of the spectral type. The donor velocity, $K_2$, lies in the range about 400–550 km s$^{-1}$, consistent with the value of 442 km s$^{-1}$ using the model-dependent orbital parameters estimated from photometric study in P2000. Observations with a larger telescope, providing higher spectral and temporal resolution and signal-to-noise would facilitate a direct and accurate determination of both the spectral type and velocity of the donor star in IY UMa.

Figure 6.9: Comparison of skew maps (covering entire phase range) for each template.

6.6 Conclusions
Chapter 7

Blue Spectroscopy of IY UMa

7.1 Introduction

In this chapter, analysis of extensive time-resolved spectroscopy of IY UMa taken during quiescence and the rise to outburst is presented, showing emission lines from both the accretion flow and the donor star, with absorption from the white dwarf or boundary layer. In the previous chapter, absorption features from the donor were detected. Clear detections of all such features in one system are rare, demonstrating again what an ideal laboratory for studies of accretion flow has been handed to astronomers with the discovery of IY UMa. Preliminary work on the March 19th observations was presented at the Astro-Tomography workshop in Brussels in July 2000, with an entry in the conference proceedings (Rolfe, Abbott & Haswell 2001).

7.2 Observations

The observations consist of 13 orbits of spectra from March 2000 and January 2001 taken using the ALFOSC spectrograph on the Nordic Optical Telescope using grisms 6, 7 and 8. These observations are summarized in Tables 7.1 and 7.2. This includes the spectra described in Section 6, but in this section only the blue end of those spectra is discussed. The March 2000 data were obtained during technical time while the January 2001 observing time was awarded by the NOT time allocation committee.

The January 2001 exposures of IY UMa included the nearby comparison star on the slit for correction of slit losses. Wide slit exposures of IY UMa and the comparison, and
Table 7.1: The Observations. $\Delta \lambda$ is the FWHM resolution of the spectra measured from arc lines. $N$ is the number of spectra.

<table>
<thead>
<tr>
<th>Date</th>
<th>Grism</th>
<th>Orbits</th>
<th>$N$</th>
<th>Exp. time (s)</th>
<th>$\lambda$ range (Å)</th>
<th>$\Delta \lambda$ (Å)</th>
<th>Observer(s)</th>
</tr>
</thead>
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<tr>
<td>18 Mar 2000</td>
<td>6</td>
<td>4.1</td>
<td>71</td>
<td>300</td>
<td>3180–5550</td>
<td>6</td>
<td>Abbott</td>
</tr>
<tr>
<td>19 Mar 2000</td>
<td>7</td>
<td>2.9</td>
<td>80</td>
<td>180</td>
<td>3820–6840</td>
<td>6</td>
<td>Abbott</td>
</tr>
<tr>
<td>20 Mar 2000</td>
<td>7</td>
<td>1.6</td>
<td>64</td>
<td>120</td>
<td>3820–6840</td>
<td>5</td>
<td>Abbott</td>
</tr>
<tr>
<td>3 Jan 2001</td>
<td>8</td>
<td>1.6</td>
<td>21</td>
<td>360</td>
<td>5810–8350</td>
<td>5</td>
<td>Rolfe, Abbott</td>
</tr>
<tr>
<td>4 Jan 2001</td>
<td>8</td>
<td>1.2</td>
<td>18</td>
<td>300</td>
<td>5810–8350</td>
<td>5</td>
<td>Rolfe, Abbott</td>
</tr>
<tr>
<td>6 Jan 2001</td>
<td>7</td>
<td>1.1</td>
<td>58</td>
<td>60</td>
<td>3820–6840</td>
<td>5</td>
<td>Rolfe, Abbott</td>
</tr>
<tr>
<td>7 Jan 2001</td>
<td>7</td>
<td>2.2</td>
<td>131</td>
<td>60</td>
<td>3820–6840</td>
<td>5</td>
<td>Rolfe, Abbott</td>
</tr>
</tbody>
</table>

Table 7.2: HJD at the start of each set of observations.

<table>
<thead>
<tr>
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<th>HJD</th>
</tr>
</thead>
<tbody>
<tr>
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<td>2451622.40</td>
</tr>
<tr>
<td>19 Mar 2000</td>
<td>2451623.42</td>
</tr>
<tr>
<td>20 Mar 2000</td>
<td>2451624.38</td>
</tr>
<tr>
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<td>6 Jan 2001</td>
<td>2451916.61</td>
</tr>
<tr>
<td>7 Jan 2001</td>
<td>2451917.59</td>
</tr>
</tbody>
</table>

of a flux standard, were taken for flux and slit loss calibration. Arc lamp exposures were taken at each telescope position. Various halogen lamp and sky flats were also taken.

7.3 Data reduction

The March 2000 observations were reduced by Dr. Tim Abbott, following the standard procedures described in Chapter 2 using IRAF. Corrections for slit losses could not be made since the comparison star was not placed in the slit, but flux standard observations were
used to correct the spectra for atmospheric extinction and the instrumental response. The March 18th spectra are only reliable between about 4000Å and 5200Å due to wavelength calibration problems.

For the January 2001 grism 7 observations a 98×2048-pixel strip of the CCD was used, minimizing the readout time while including both IY UMa and the comparison in each exposure. The blue grism suffers from none of the fringing problems of the red grism described in the previous chapter. Standard reduction steps were followed using IRAF packages. No illumination correction was applied since there was no significant spatial gradient in the sky exposures. Observations of the standard star reveal no appreciable telluric features in this range, so telluric correction was not required. A second wide slit (5") exposure of IY UMa and comparison was taken, immediately preceded by wide slit exposure of the nearby flux standard. These exposures enabled a flux calibrated slit loss corrected spectrum of IY UMa and (more importantly) its comparison to be obtained.

The flux calibrated comparison spectrum was binned into 40 wavelength bins, and divided by the similarly binned comparison spectrum in each uncalibrated frame. Fitting a 3rd order polynomial to this divided spectrum provided a smooth wavelength-dependent calibration spectrum for the frame. Each IY UMa spectrum was then multiplied by its calibration spectrum, thus producing flux calibrated and slit loss corrected spectra of IY UMa. The comparison star is very faint towards the blue end of the spectrum (roughly the first 650 pixels) causing the trace of its aperture to be lost in this range in many of the narrow slit object frames. Therefore, the above flux calibration can only be trusted for longer wavelengths than about 4800Å. Using the wavelength dependent sensitivity function obtained when calibrating the wide slit exposures with the wide slit flux standard, all the object frames were calibrated to produce a second set of flux calibrated spectra. This set of spectra is not corrected for slit losses, and so is calibrated only as well as the March 2000 spectra. However, by scaling each one of these spectra to match the corresponding slit loss corrected spectrum in the 100 pixels around 4800Å, and then replacing the slit loss corrected spectra with these scaled spectra below this range, a set of spectra was produced which is fully flux calibrated and slit loss corrected above ~4800Å, and degraded only by the wavelength dependent slit losses at shorter wavelengths.

The red end of the wide slit comparison spectrum was also used to flux and slit loss calibrate the blue end of the red spectra from 3rd and 4th January.
7.4 Average spectra

7.4.1 March 2000

Figure 7.1 shows the average nightly spectra for the March 2000 run. Plotted are the average spectrum during white dwarf eclipse (phases -0.03 to 0.03) and an average of phases 0.15 to 0.9, when neither hotspot nor white dwarf is eclipsed. These phase ranges were chosen after studying the lightcurves discussed in Section 7.6.1. The spectra have been gaussian smoothed with FWHM 4.5Å (3 pixels) which is less than the instrumental resolution and so does not remove any spectral information.

The spectra reveal a blue continuum. There are strong, double-peaked Balmer, He I and Fe II emission lines, clear signatures of Doppler shifted emission from an accretion disc, and commonly seen in dwarf nova spectra. Broad absorption wings are seen around the Balmer emission lines except Hα. Deep narrow absorption cores are seen in the emission lines, strongest for the higher order Balmer lines, He I and Fe II. These features are all seen in two other high inclination dwarf novae in quiescence - Z Cha (Marsh, Horne & Shipman 1987) and OY Car (Hessman et al. 1989). These systems both have very similar orbital parameters to IY UMa. Z Cha has $q=0.15$ and $P_{orb}=1.80$ hours (Wade & Horne 1988; Robinson et al. 1995) while OY Car has $q=0.10$ and $P_{orb}=1.51$ hours (Wood et al. 1986a). There is also a weak Na I doublet absorption at 5890-5896Å superimposed on the He I 5876Å emission.

The full width of the Balmer absorption wings is very large - about 20,000 km s$^{-1}$. This is far too large to be a Doppler shift from fast moving disc material, even in the inner disc. At a radius of $R_{wd} \approx 0.012a$ (the white dwarf radius, estimated from the 25-second white dwarf eclipse ingress/egress duration measured in P2000), the Keplerian disc velocity is still only $\sim 4300$ km s$^{-1}$. There must be some other broadening mechanism at work to produce such wide lines. Marsh, Horne & Shipman (1987) concluded in a study of Z Cha that Stark broadening is being seen. Stark broadening results from perturbations of energy levels in the emitters by the electric fields of nearby charged particles, and is thus only expected in the high density regions on or close to the white dwarf. The spectra in Figure 7.1 suggest that the broad absorption wings only disappear during the white dwarf eclipse. Figure 7.2 shows the average continuum-subtracted flux in the blue wing of Hβ over velocity range -5000 km s$^{-1}$ to -2500 km s$^{-1}$, combining the data from March 18th, 19th and 20th, folding and binning them in orbital phase. The average error in each phase bin is $0.15 \times 10^{-16}$ erg s$^{-1}$ Å$^{-1}$ cm$^{-2}$. The depth of absorption away from eclipse
7.4 Average spectra

Figure 7.1: Average spectra of IY UMa during quiescence in March 2000.
7.4 Average spectra

is around $0.6 \times 10^{-16} \text{erg s}^{-1} \text{Å}^{-1} \text{cm}^{-2}$ (about 12% of the continuum at $\text{H}\beta$), while during white dwarf eclipse (marked with the dashed vertical lines) it decreases rapidly to around 0.

![Figure 7.2: Phase folded and binned continuum-subtracted flux averaged over the blue absorption wing of $\text{H}\beta$.](image)

We now follow the arguments applied to the H$\beta$ blue wing lightcurves for Z Cha in quiescence (Marsh, Horne & Shipman 1987). Just outside white dwarf eclipse, in the phase bins corresponding to phases 0.94 to 0.96 and 0.04 to 0.06, the absorption is only about 25% to 30% weaker than away from eclipse, providing convincing evidence that the absorption is eclipsed only in phase range -0.06 to 0.06. The white dwarf mid-ingress is at phase -0.032, and lasts for 0.0039 in phase (25 seconds, P2000), so the H$\beta$ absorption is confined to within radius $(0.06 - 0.032)/0.0039 R_{\text{wd}} = 7.2 R_{\text{wd}} = 0.09a$. Stanishev et al. (2001) used the eclipse mapping technique (Horne 1985) to deduce a roughly flat temperature distribution for the quiescent disc in IY UMa. Assuming a flat temperature distribution and a disc radius of 0.25a (Stanishev et al. 2001), we find that the central 0.09a of the
7.4 Average spectra

disc can contribute only 13% of the disc flux. Looking at the orbital lightcurve of IY UMa shown in Figure 5.1, the white dwarf flux contributes about 70% of the continuum flux away from the orbital hump, so the disc can contribute at most 30% of the continuum flux. The inner 0.09a can therefore contribute at most about 4% of the continuum flux, but, as already noted, away from eclipse the depth of absorption in the blue wing of Hβ is about 12% of the continuum. Therefore the absorption in the disc can contribute only about 1/3 of the broad Balmer absorption, (c.f. 30% in Z Cha (Marsh, Horne & Shipman 1987)). The broad absorption must come predominantly from the white dwarf, as in Z Cha and also OY Car (Hessman et al. 1989).

Marsh, Horne & Shipman (1987) fitted model white dwarf spectra to the absorption wings in Z Cha revealing that although Hα does not exhibit the absorption wings, there is still a broad Stark-broadened line profile under the emission line. Hessman et al. (1989) obtained similar fits for OY Car, although their observations didn’t include Hα. However, we can safely expect that in both IY UMa and OY Car the Hα emission does sit on a broad Balmer absorption from the white dwarf.

The hotspot spectrum

The very high inclination of IY UMa leads to a prominent orbital hump as the hotspot comes into view. From the continuum lightcurves, discussed in Section 7.6.1, the phase ranges corresponding to maximum and minimum visibility of the orbital hump are 0.75 to 0.90 and 0.5 to 0.6 respectively. Average spectra were calculated for each night in each of these ranges. The spectra for the 19th and 20th March were grouped together since they cover the same spectral range and clearly agree well out of eclipse (Figure 7.1). The spectrum around hump minimum was then subtracted from that around hump maximum to produce two hotspot spectra - one for the 18th March, and one for the 19th and 20th March. These are shown in Figure 7.3, smoothed as in Figure 7.1. This difference in the spectrum between orbital hump maximum and minimum should be almost entirely due to the change in visibility of the hotspot, assuming the disc brightness is not changing with phase. The line profiles from the disc should also change little, with the double peaks simply shifting by less than 1Å due to the orbital motion of the white dwarf. The spectra in Figure 7.3 therefore provide a good measure of the hotspot spectrum, although due to the complicated structure of the hotspot, they are likely to only show a fraction of the hotspot flux.
Figure 7.3: The spectrum of the orbital hump, obtained by subtracting the average spectrum at hump minimum from the average at hump maximum. The smooth curves are blackbody fits to the data.

The most prominent features are the strong emission lines in the Balmer series. The He I and Fe II lines are also seen. The lines show one strong peak, as expected from a strongly localized emission region. There are also weaker peaks and dips in some of the lines, a result of the change in projected velocity of the hotspot between the peak and minimum of the orbital hump and possibly other small changes in the line profile unrelated to the hotspot. This spectrum is in stark contrast to the hump spectrum of Z Cha, obtained using the same technique, where no line emission from the hotspot was seen (Marsh, Horne & Shipman 1987). OY Car and V2051 Oph also show no sign of line emission from the hotspot (Bailey & Ward 1981; Watts et al. 1986), although there are a variety of systems in which line emission from the hotspot is seen - e.g. WZ Sge (Krzeminski & Kraft 1964; Spruit & Rutten 1998), U Gem (Smak 1976) and V893 Sco (Matsumoto, Mennickent & Kato 2000).
The continuum is blue, as we would expect from the hot stream-disc impact region. To estimate the temperature of the stream-disc impact, the continuum was fitted with a simple blackbody spectrum, calculated using Equation 7.1.

\[
F(\lambda, T) = 10^{-7} \frac{A}{d^2} \frac{1}{\lambda^5} \frac{1}{e^{hc/\lambda kT} - 1} \text{erg s}^{-1} \text{Å}^{-1} \text{cm}^{-2}
\]  

(7.1)

\(A\) is the area of the emitting region, \(d\) is the distance to the source and \(T\) is its temperature. The fit was done by minimizing the reduced \(\chi^2\), \(\chi_r^2\) (defined in Equation 3.2), using the amoeba routine from Press et al. (1992) as implemented in IDL. There were only two free parameters - \(T\) and \(A\). The effects of \(A\) and \(d\) are both just to scale the spectrum, so \(d\) was fixed using the estimate in P2000 of 190±60 parsecs. Formal errors on the fitted \(A\) and \(T\) were estimated by using a grid of values of \(\chi_r^2\) to determine the largest deviation of each parameter from the best fit for which \(\chi_r^2 \leq \chi_{r,\text{min}}^2 + \Delta\chi_r^2\). A value \(\Delta\chi_r^2 = 4.61/(N - M)\) was set, where \(N\) is the number of points in the spectrum and \(M = 2\) is the number of degrees of freedom in the model. If the noise in the spectra is normally distributed then the errors determined this way will be 90% confidence limits. Since the hotspot spectra come from averages of a number of spectra, the noise should be roughly normally distributed.

Fitting a blackbody to the combined March 18th, 19th and 20th hotspot spectra yields \(T = 10659\pm76\) K and \(A = 0.00051\pm0.00001\) \(a^2\). Taking into account the error range on the distance and fitting all three nights of data gives \(T = 10660\pm78\) K and \(A = 0.0005\pm0.0003\) \(a^2\). We take this as our most reliable determination of the temperature and effective emitting area of the hotspot in IY UMa.

The effective area determined here is consistent with what we expect for the stream-disc impact. From the basic model and parameters used in Section 5.5.1, the difference between the projected area of the hotspot at the maximum of the hump and at the minimum will be of order \(\sim 0.001 - 0.01a^2\). This is consistent with the fitted area given that this does not take into account likely anisotropy in the emission from the hotspot due to the complicated behaviour of the stream-disc impact.

Stanishev et al. (2001) determined the hotspot temperature in IY UMa on 6th February and 11th March 2000 as 15900K and 13800K respectively, both with errors of about 500K. Although our value does not agree with either of these values within confidence limits, the two values are also in disagreement with one another, suggesting either that the hotspot changed temperature, which is perfectly plausible, or that Stanishev et al. (2001)
underestimated their uncertainty. They used the total flux at orbital hump maximum and minimum, and so had to estimate the emitting area from maximum entropy eclipse maps, and then account for foreshortening. It is possible that this lead to an underestimation of $A$ by a factor of $\sim 5$, enough to account for the difference between their $T$ and this spectroscopic measurement. The spectroscopic blackbody fit fully constrains both $T$ and $A$ for a given value of $d$ (for which Stanishev et al. (2001) also used the 190 pc value) and so is more likely to reflect the true hotspot temperature.

Several measurements of the hotspot temperature during quiescence have been made in other similar systems. Robinson et al. (1995) determined the hotspot temperature in Z Cha as $16700\pm600$ K, significantly higher than this IY UMa measurement, however they also made an earlier measurement (Wood et al. 1986b) of $11300\pm2000$ K, again suggesting significant variations in hotspot temperature over time or uncertainty in photometric determination of temperature. Marsh, Horne & Shipman (1987) produced an orbital hump spectrum using the technique we use here, and found that a blackbody did not describe the spectrum well, observing a significant drop below 4000 Å, probably due to Balmer absorption. The spectral coverage of our observations prevents us from seeing any such drop in IY UMa. Schoembs & Hartmann (1983) estimated the hot spot temperature in OY Car as $15000\pm5000$ K.

All these measurement reveal hotspots with temperatures between 11000 and 17000 K. Variations of this size could simply result from the variation in disc radius and shape at different epochs of the eruption cycle, considering the simple model for hotspot brightness presented in Chapter 5. However the two different Z Cha temperatures correspond to the same disc radius, while the highest Stanishev et al. (2001) temperature corresponds to the largest radius, the reverse of what is expected if a varying disc radius causes these temperature changes. Variations in hotspot temperature might also result from other variations in the geometry of the impact, for example stream-disc overflow. The current picture is rather muddled, and clearly more accurate spectroscopic measurements of the hotspot temperature and accretion flow geometry are needed to understand this behaviour.

7.4.2 January 2001

Figure 7.4 shows the average spectra from the January 2001 observing run. Those from January 3 and 4 show just the blue end of the red spectra which were discussed in the previous chapter, and are grouped and averaged as those from March 2000. Note that the
Figure 7.4: Average spectra of IY UMa during rise to outburst in January 2001.

spectra for the 6 and 7 January have been plotted with larger flux scales and averaged differently, due to the different behaviour of the system on these days. The spectra have been gaussian smoothed with FWHM 4.5Å.

On 3 and 4 January 2001, the spectra show exactly the same Hα behaviour seen in March 2000, with strong double-peaked disc emission and a deep core. The He I 5876Å emis-
sion is also present as in March 2000, but the core absorption is significantly stronger. The unresolved Na I doublet is again seen. The continuum flux level is consistent with that in March 2000. The system appears to be in much the same quiescent state as during the earlier observations.

The spectra from the 6 January present a very different picture. Immediately we see that the system is 7-8 times brighter. The double-peaked emission lines are still present, and deep core absorption is seen in all emission lines. This core absorption is so strong that it completely dominates the emission in He II (except He I at 5876 Å) and Fe II, closely resembling OY Car during a normal outburst in 1991 (Harlaftis & Marsh 1996) and some of the superoutburst spectra of Z Cha from 1984 and 1985 (Honey et al. 1988). The Balmer absorption wings are no longer seen, except possibly around Hγ. We also see the emergence of emission in the C III/N III Bowen blend at 4640 Å and He II at 4686 Å. This feature was seen in spectroscopic observations of the January 2000 superoutburst (Wu et al. 2001), and has been seen in other outbursting dwarf novae, e.g. Z Cha in superoutburst (Honey et al. 1988), OY Car in normal outburst (Harlaftis & Marsh 1996) and IP Peg (Morales-Rueda, Marsh & Billington 2000), and in nova-likes e.g. V348 Pup (see Chapter 4), and UX UMa (Rutten et al. 1993). The He II emission probably results from the reprocessing of EUV and X-ray emission from the boundary layer between the inner disc and white dwarf (Patterson & Raymond 1985), while the N III emission may result from conversion of He II Lyα transition photons via O III to N III photons (Deguchi 1985). The line profile is broad, with a width similar to that of the Balmer lines, implying that this emission is coming from the disc.

The spectra from 7 January are different again. The continuum has become about 30% weaker and the core absorptions in the Balmer emission lines have become much weaker, particularly in Hα and Hβ. The absorption in Fe II has disappeared completely, although it remains in He I. Notably, the He II 4686 Å line and the Bowen blend have become much stronger, with the He II now being comparable in strength to Hβ. The He II line at 5412 Å is now faintly visible above the continuum. He II 4686 Å shows clear double-peaked structure, indicating its source in the disc. This is clear in Figure 7.5, which shows the average out of eclipse spectrum in the range 4530–5000 Å, with the rest wavelengths of C III/N III, He II, Hβ marked with dashed lines.

Clearly, IY UMa went into outburst between the 4th and 6th January, and continued in outburst on the 7th. The accretion rate onto the white dwarf must have been several
Figure 7.5: Average out of eclipse spectrum from 7 January 2001, showing the double-peaked emission in He II, with the Hβ profile for comparison.
7.5 The systemic velocity

The radial velocity of the Hα line was measured for each of the quiescent spectra from March 2000. This was done by fitting a double gaussian profile, and using the mid-velocity of the two peaks in each spectrum as the radial velocity. The measurements are shown in Figure 7.6.

A sinusoid of the form \( V = \gamma - V_1 \sin 2\pi (\phi - \phi_0) \) was fitted to those velocities outside eclipse during quiescence in March 2000. Hα was used since this is both the strongest line, and also does not exhibit the Balmer absorption wings. It is still likely to be sitting on a broad Balmer absorption line as is the case in Z Cha (Marsh, Horne & Shipman 1987), but suffers less from this effect than the other Balmer lines.

We obtain systemic velocity \( \gamma = 13.6 \pm 1 \, \text{km} \, \text{s}^{-1} \), velocity amplitude \( V_1 = 105.4 \pm 1 \, \text{km} \, \text{s}^{-1} \) and \( \phi_0 = 0.087 \pm 0.002 \). This measured systemic velocity is consistent with that measured during the January 2000 superoutburst by Wu et al. (2001), who measure \( \gamma = -4 \pm 32 \, \text{km} \, \text{s}^{-1} \) from Hα. The formal errors given here underestimate the errors. Systematic errors arise, particularly depending on the range of spectra excluded during eclipse. This result comes from excluding phases -0.15 to 0.15, while excluding just -0.1 to 0.1 gives \( \gamma = 20 \pm 1 \, \text{km} \, \text{s}^{-1} \), \( V_1 = -108.4 \pm 1 \, \text{km} \, \text{s}^{-1} \) and \( \phi_0 = 0.098 \pm 0.001 \).

\( V_1 \) is very different from the 62 km s\(^{-1}\) predicted velocity amplitude of the white dwarf. Along with the large phase shift, \( \phi_0 \), relative to white dwarf mid-eclipse, we see that the emission does not follow the motion of the white dwarf. The phase shift is the same as that seen in the unusual dwarf nova WZ Sge (Skidmore et al. 2000) and also similar to those in OY Car (measurements summarized in Hessman et al. (1989)) and Z Cha (Marsh, Horne...
7.5 The systemic velocity

Figure 7.6: Hα velocity measurements. These are heliocentric velocities. Purple points during eclipse are ignored during the fit. Green points are rejected by the fitting routine because the deviate by more than 3σ from the fit. The red curve is the best fit sinusoid.

& Shipman 1987). Such phase shifts are seen in quiescent low mass X-ray transients e.g. V616 Mon (Haswell & Shafer 1990). It has become clear that the measurement of the white dwarf velocity from emission lines is very unreliable, a result of the accretion flow being far more complicated than a simple circular disc. Marsh, Horne & Shipman (1987) concluded that the emission lines in Z Cha may have both non-axisymmetric velocity and brightness distributions. This is not surprising - the stream-disc impact will clearly have some effect on the brightness and velocity distribution of the disc. Tidal effects may also come into play.

In theory, the peak separation of the double peaked disc emission lines is a good measure of velocity at the outer edge of the disc (Smak 1981). This is because the low velocity of the outer disc leads to it occupying a small velocity range when projected towards the observer, concentrating the outer disc emission into the sharp peaks. Figure 7.7 shows the peak-peak velocity separation. We see strong variation in the peak separation as a function of orbital phase. Between phases 0.4 and 0.6 the separation is fairly constant at around 1400–1450 km s\(^{-1}\). However between about 0.1 and 0.4, and again from phase
0.6 to 0, the velocity dips down towards a minimum of about 1100 km s\(^{-1}\). This is easily understood, however. The dips correspond to phases when the velocity of emission from the hotspot is crossing between the disc peaks, first from red to blue, then from blue to red. This S-wave from the hotspot is presented and discussed in detail in Section 7.7.1. To obtain the best possible measure of the peak separation in IY UMa, we take the average separation in the phase range 0.4 to 0.6. This is 1412±50 km s\(^{-1}\). Assuming a Keplerian velocity distribution in the disc, the disc radius corresponding to half of this velocity separation is 0.45±0.02\(a\). This radius seems rather too large: it coincides with the location of the 3:1 resonance, something which we only expect during a superoutburst.

In practice, measuring the disc radius from the separation of double-peaked disc emission usually gives larger radii than using other methods (Wade & Horne 1988; Marsh 1988). Stanishev et al. (2001) determined a radius of 0.25\(a\) in IY UMa only a week earlier while P2000 estimated a quiescent disc radius of 0.28±0.02\(a\) from the hot spot S-wave. Given that the effect of the hotspot is to reduce the estimated disc velocity as it crosses between the peaks, it is possible that it is also having the same effect between phases 0.4 and 0.6, except that in this range the effect is fairly uniform since the projected spot velocity is at its maximum excursion. Any reduction in the peak separation will lead to an increase in the implied disc radius. If the hotspot reduces the apparent peak separation by 160 km s\(^{-1}\), then the true disc radius would in fact be 0.36\(a\). This is a possible explanation for the large radius measured here.

### 7.6 Lightcurves

Lightcurves of the continuum and the H\(\alpha\) and H\(\beta\) emission lines for the March 2000 observations are shown in Figure 7.8, while Figure 7.10 shows these plus the He\(\text{II}/\text{Bowen}\) blend lightcurve for the January 2001 run.

The total flux (top panels of the plots) is simply the integrated flux in each spectrum. The ‘continuum’ flux is the flux integrated in the range 5950 Å–6450 Å, this range being chosen simply because it is covered by all spectra from both observing runs, except 18 March 2000. The H\(\alpha\) lightcurve is the integrated flux over H\(\alpha\) from the continuum subtracted lightcurve, where the continuum was fitted with a 2nd order polynomial. For H\(\beta\), before integrating over the continuum subtracted line profile, the average flux either side of line was subtracted to reduce the effect of the Balmer absorption wings. For the He\(\text{II}/\text{Bowen}\) blend, the continuum was fitted with a straight line either side of the profile.
to improve the continuum fit.

7.6.1 March 2000

The total flux lightcurves for March 2000 reveal the strong orbital hump and eclipse, seen more clearly in the high resolution photometry shown in Figures 5.1 and 5.10. The low time resolution of these observations blurs out most of the eclipse structure, such as the different ingress times of the white dwarf and hotspot eclipses. The March 18 continuum flux is much lower than for March 19 and 20 due to the smaller wavelength range of the spectra on that night. The most notable difference between these lightcurves and those studied in Chapter 5 is the apparent double orbital hump. On the 19th and 20th, where the structure appears to be far more stable than on the 18th, the hump rises from a minimum around phase 0.5–0.6, and reaches minimum again at about 0.2, just after eclipse. It then rises to a lower maximum value around phase 0.3 to 0.4, before falling again to minimum around phase 0.5–0.6. This can be easily understood. The strongest orbital hump is the traditional hump, seen as the stream-disc impact comes into view on the near side of the disc. At phase 0.3-0.4, the stream disc impact is again seen roughly length ways, but on the far side of the disc, so we only see the top of the hotspot, and consequently a weaker
hump. Such double humped orbital curves have also been seen in WZ Sge and AL Com (Robinson, Nather & Patterson 1978; Patterson et al. 1996), but with both humps having the same amplitude. Robinson, Nather & Patterson (1978) interpreted the modulation in WZ Sge as a result of obscuration of the hotspot by a thick disc rim at the hump minima, rather than due to a varying projected area of the hot spot. On 18th March, the structure is far less stable from one orbit to the next, possibly a result of unphotometric conditions and lack of a comparison star being placed on the slit. The behaviour is more like that seen in Figure 5.10, with one large orbital hump, although the last orbit is starting to look a little more like those of March 19th and 20th.

The continuum lightcurves in the second panel (covering only 5950Å–6450Å) look exactly the same as the total curves. The Hα curve is also very similar, although the relative intensity of the secondary to the primary orbital hump is higher than in the continuum. The Hα eclipse is much shallower than the continuum eclipse. This is because, as discussed in Section 7.4.1, the broad Balmer absorption lines will be obscured during white dwarf
eclipse, reducing the apparent eclipse depth in the continuum subtracted emission lines. In addition, the narrow absorption cores were also found to disappear during eclipse in Z Cha (Marsh, Horne & Shipman 1987), suggesting that these may also contribute to the shallow eclipse in Hα. The Hβ lightcurves for March 19th and 20th show the double-humped structure again, with a shallower eclipse than in Hα, reflecting the increased strength of the two absorption components relative to the emission line. The March 18th Hβ curve is again unstable, but to the extent to which we can draw conclusions from its structure, it also shows the double hump and weak eclipse seen in the following nights.

7.6.2 January 2001

Figure 7.9 shows the lightcurves for 3rd and 4th January 2001 with a smaller flux scale than in Figure 7.10, making the details clearer than in the latter figure. The continuum lightcurves covering 5950Å–6450Å show just the usual single orbital hump, with no evidence of the second hump seen in the March 2000 lightcurves. We do not see the first eclipse on 3rd Jan simply because no spectrum was taken at mid-eclipse. The flux away from the hump is about $10^{-20} \times 10^{-14}\text{erg}\text{s}^{-1}\text{cm}^{-2}$ compared to about $10\times 10^{-14}\text{erg}\text{s}^{-1}\text{cm}^{-2}$ in March 2000. These values are consistent, particularly given that the 2000 observations are likely to have lost some flux due to slit losses. The peak of the orbital hump reaches at least $40\times 10^{-14}\text{erg}\text{s}^{-1}\text{cm}^{-2}$ on 3rd January, but is only $25\times 10^{-14}\text{erg}\text{s}^{-1}\text{cm}^{-2}$ on the
4th. These observations have been corrected for slit losses, and so this change in orbital hump amplitude is likely to be genuine. There are many possible explanations for such a change in the amplitude of the orbital hump. Chapter 5 showed how variations of the disc radius and stream-disc velocity orientation at the hotspot can lead to such changes in hotspot brightness. If this is the cause, then the disc radius at the impact point must have increased between the 3rd and 4th January. A change in the accretion rate from the stream, or a varying amount of stream overflowing the disc edge could also explain such a change.

The Hα curves are more unusual, with the second hump and eclipse clear on 3rd January, and a sharp drop in flux at phase 2.4. The eclipses are clear on the 4th, but the lightcurve is quite flat between the eclipse.

The low time resolution and limited number of orbits prevents any detailed analysis of these lightcurves, but they do suggest dramatic changes in the behaviour of the system over one day. This is a surprising result if it is linked to the following outburst: no advanced warning of an impending outburst has been seen before or is expected.

We now turn our attention to Figure 7.10, which includes the 3rd and 4th January lightcurves again, but now plotted with a larger flux scale, enabling the lightcurves from the 6th and 7th January to be plotted on the same scale for comparison.

We see that the total lightcurves in the top panel show the same behaviour as the 5950Å–6450Å continuum lightcurves in the second row, and so we discuss only the continuum lightcurves, for which we have coverage of all four nights. Between the 4th and 6th January, the continuum flux has risen by a factor of about 7, decreasing again by the beginning of the last night's observations (phase 54.3). During the last night, the flux continues to drop slowly by a further 20 percent by phase 56.6. The eclipses are all deep, with the eclipse at phase 42 being broader and shallower than those at phases 55 and 56 on the 7th March, implying a disc which shrinks between the 6th and 7th March. Measurement by eye yields a full eclipse width of 0.20±0.04 on the 6th January, and 0.14±0.02 on the 7th. These widths correspond to emission distributions of radius 0.37±0.10a and 0.21±0.05a on the 6th and 7th respectively. These radii are both much smaller than the 3:1 resonance radius (0.46a) which the disc must reach to trigger the tidal instability involved in a superoutburst, while P2000 measured a disc radius of 0.44±0.03a at the maximum of the January 2000 superoutburst. This further establishes that the outburst discussed here was a normal outburst. If this change in eclipse width was interpreted as evidence
Figure 7.10: January 2001 lightcurves
for a precessing eccentric disc, it would imply an unrealistically large eccentricity of $\sim 0.8$
for an elliptical disc.

I know of four sets of observations of eclipsing dwarf novae beginning early during
the rise to outburst. Vogt (1983a) presents the rise to a normal outburst of OY Car,
while Webb et al. (1999) presents an entire outburst of IP Peg. Ioannou et al. (1999)
present observations of another short period high-inclination dwarf nova, HT Cas, covering
the rise, peak and decline from a normal outburst in 1995. The OY Car observations
revealed very similar changes in the disc emission distribution to those in IY UMa. Vogt
(1983a) concluded that during the rise to outburst, the outer disc rapidly brightens (as
it enters the hot ionized state), with this hot bright region quickly propagating inwards
towards the white dwarf, with the outer radius remaining constant. As the outburst
reaches maximum brightness, the radius of the disc rapidly shrinks by about a third.
These results are beautifully illustrated in the Rutten et al. (1992) re-analysis of the
Vogt (1983a) observations. This re-analysis involves the production of maximum entropy
eclipse maps, providing spatial images of the emission distribution throughout the rise
and early maximum of the outburst, images which confirm the Vogt (1983a) conclusion
that the OY Car outburst was an outside-in outburst, beginning in the outer disc and
propagating inwards. Ioannou et al. (1999) find exactly the same outside-in behaviour in
HT Cas, additionally concluding that the disc becomes flared during outburst. Our low-
resolution lightcurves prevent us from using the eclipse mapping technique, but we note
that the shrinking emission region and increasing eclipse depth we see in IY UMa between
the outburst maximum and a day later is exactly what was seen in OY Car and HT Cas.
However, the disc shrinks during decline in most (if not all) dwarf novae, including the
two longer period dwarf novae where eclipses during rise to outburst suggest an outburst
beginning in the inner disc and propagating outwards (IP Peg (Webb et al. 1999) and
EX Dra (Baptista & Catalán 2001)).

The lightcurves of the emission lines are very similar to one another, although they are
noisier for H$\beta$ and the He ii/Bowen blend than for H$\alpha$. In contrast with the behaviour of
the continuum, the emission lines are still all rising on the 6th January, being nearly twice
as bright at the start of the observations on the 7th than during the 6th. These changes
are not simply a result of any variations in the core absorption. Lightcurves which exclude
the region between the peaks of the lines show the same behaviour. On the 7th the line
fluxes gradually decrease by about 10% to 20%, in the same way that the continuum does.
7.7 Doppler tomography

The eclipses in Hα suggest the same decrease in radius of the line emission region as for the continuum. The Hβ and HeⅡ/Bowen blend eclipses on the 6th are too messy to draw any such conclusions.

7.7 Doppler tomography

7.7.1 March 2000

Trailed spectra

We now begin to study the temporal variation of the emission line profiles in detail. Figure 7.11 shows trailed spectra of the March 2000 observations. These were produced by binning each line profile in both velocity space and in phase, with 40 bins each of width 100km s\(^{-1}\) in velocity, and 25 phase bins per orbit. The trailed spectra are shown for Hα, Hβ, Hγ and HeⅠ. For all but the 18th March, the HeⅠ 5876Å line is shown, which is the strongest of the HeⅠ lines. For the 18th March the HeⅠ 5016Å line is shown (the spectral range on this night did not cover 5876Å). The wavelength of the weak NaⅠ absorption doublet line at 5890–5896Å corresponds to a velocity range of about 700–1000 km s\(^{-1}\) in the trailed spectra, so care must be taken to be wary of this when studying the HeⅠ 5876Å images. The colour scale of each map goes from dark blue for the continuum level (taken as the average flux at velocities ±2000 km s\(^{-1}\)) to deep red for the maximum value. Levels below the continuum are shown in the same deep blue as the continuum level.

Apart from the 18th March HeⅠ 5016Å trail, which is too noisy to see any more than the clear double peaks from the disc, we see that the behaviour in each line is consistent for all 3 nights. We therefore consider the trailed spectra combining all three nights of data, shown in Figure 7.12. These trails are phase-folded, and binned into 30 bins per orbit.

In the Balmer lines we clearly see the double peaks of the disc emission, with faint evidence of emission from the disc also visible in the HeⅠ emission. The red peak (positive velocity) in the HeⅠ lines is weaker than the blue peak, and the region within a few hundred km s\(^{-1}\) of 1000 km s\(^{-1}\) has flux less than or equal to the continuum level. This is clearly the NaⅠ absorption. Between the disc peaks is the core and white dwarf Balmer absorption, although this is plotted at the continuum level in these trails. We see a very strong S-wave with phase-dependent brightness in all four lines, corresponding to a concentrated region of emission in velocity space. The strength of the disc emission is similar to that of the
Figure 7.11: Phase-binned, velocity-binned and continuum-subtracted trailed spectra. Highest fluxes are deep red, fluxes equal to and below continuum are deep blue.

S-wave emission in Hα, but the relative strength of disc to S-wave decreases for higher order Balmer lines, with the brightest parts of the S-wave dominating the disc emission in Hγ. The S-wave is also dominant in He I.

The eclipse of the disc begins in the blue and ends in the red, as first the approaching
7.7 Doppler tomography

side of the disc and then the receding side is occulted by the donor star. In all four lines, the disc emission in the red wing around phase 0.3 is weaker than at other phases (except eclipse). It looks like this weak region of red wing emission possibly stretches back as far as the eclipse, although the S-wave around these phases makes it difficult to be completely certain. A similar effect was seen in WZ Sge by Spruit & Rutten (1998), where an S-wave of absorption was seen delayed in phase relative to the hotspot S-wave. There is also a brightening in the blue wing of the disc emission around phase 0.3. This is seen in all four lines, with the He I trails showing clearly that this is a brightening in the disc, and not merely the stream and disc emission coinciding in velocity space. This corresponds to the far side of the disc from the donor having a bright region. The phasing and amplitude of the S-wave is clearly the same for all lines, with maximum velocity around phase 0.1 and minimum velocity around phase 0.6. The variation in brightness of the S-wave is clearest in the He I line, where we see that is is brightest around phase 0.8, also bright around phase 0.2 and faintest between phases 0.3 and 0.6: the lightcurve of the S-wave follows that of the orbital hump. The S-wave is eclipsed, with eclipse beginning about phase 0. and ending around 0.1, telling us that its source is on the same side of the line of centres as the receding half of the disc. This phasing, late eclipse and lightcurve of the S-wave all tell us that this is the emission from the hotspot.

![Figure 7.12](image_url)

Figure 7.12: Phase-folded, velocity-binned and continuum-subtracted trailed spectra combining all 3 nights of data from March 2000.
Doppler maps

Figure 7.13 shows Fourier-filtered back-projections produced for each of the trailed spectra shown in Figure 7.7.1. Phases -0.1 to 0.1 (the eclipse) were left out, since eclipses violate assumption of Doppler tomography that all regions in the map have a constant visibility throughout the orbit. It should be noted that the hotspot also violates this assumption, and so Doppler maps will show an average of the hotspot emission visible throughout the orbit. The geometry plotted in white is explained in Chapter 2.2.2.

The Hα map shows a ring of disc emission with Keplerian velocities corresponding to locations within the tidal radius. The stream-disc impact should be seen between the two arcs corresponding to the stream trajectory and its Keplerian velocity in the top left of the map, as seen in maps of U Gem (Marsh et al. 1990) and WZ Sge (Spruit & Rutten 1998). This is seen clearly in the March 20th map, and there is also emission at the right velocity in the March 19th map, although the thickness and brightness of the disc ring make it less

![Fourier-filtered back-projections](image)

Figure 7.13: Fourier-filtered back-projections. White markings are described in Chapter 2.2.2.
Figure 7.14: Spatial geometry interpreted from Doppler map. Stream from donor star impacts disc at hotspot (white region with thick outline). Disc emission (grey) is obscured in this region, and at different regions close to the hotspot depending on orbital phase.

distinct in that image. In both Hα maps the disc emission just to the left of the donor star is very much weaker than elsewhere. This region of weaker disc emission corresponds to the weak disc emission in the red peak of the trails just after eclipse. Assuming a Keplerian velocity field, this corresponds to weaker emission from the region of the disc around the stream-disc impact (Figure 7.14). We do not expect a Keplerian velocity field where the stream and disc merge, and any underlying Keplerian emission could be obscured by an optically thick region around the impact. It is therefore no surprise that we see weaker disc emission in this velocity region.

The Hβ and Hγ maps again show the disc emission and hotspot, but in these maps the hotspot is much stronger relative to the disc than in Hα. The disc emission is brighter towards the bottom in each map, corresponding to the bright region in the blue wing of the trail around phase 0.3. Despite being seen in all trails, this feature in the Doppler map is only seen clearly in the Hβ maps.

The He I 5876Å maps show no sign of the disc emission at all. They show just the very strong bright spot. There is a dark ring in the March 19th and 20th maps where the disc should be. This ring is due to the Na I doublet, which appears as absorption at constant
velocity, and when projected into velocity space has canceled out the weak disc emission.

The behaviour of the Balmer maps is the same as that seen in WZ Sge (Skidmore et al. 2000) and also very similar to that in HE 1047 (Skidmore et al. 2002). The strong disc component in both lines, and the increase in the relative strength of the stream-disc impact to the disc in Hβ is seen in both of these systems, while the weaker region of disc emission in Hα between the hotspot and donor is seen in WZ Sge but not in HE 1047. Spruit & Rutten (1998) show an Hα Doppler map of WZ Sge which clearly shows the disc and stream-disc impact, but no clearly missing disc emission, despite the strong absorption S-wave seen in the WZ Sge spectra.

**Modeling the emission**

The code used to model the hotspot lightcurves in Chapter 5 models emission line profiles as well as lightcurves, and was used to produce a simple 3D model for the main features of the Doppler maps presented here. The geometry for the hotspot is exactly as described in Section 5.5.1. We used a disc with height $H_{\text{disc}}$ (independent of $r$), with a Keplerian velocity distribution and surface brightness varying as $S \propto r^{-2}$ (consistent with the typical power laws found by Smak (1981)). Using a more realistic disc whose thickness increases linearly with $r$ had very little effect on the results (but see the comment near the end of this section). The velocity of the hotspot flow at the initial impact point is taken to be the ballistic stream velocity at that point, $\tilde{V}_{\text{stream,0}}$. At increasing hotspot azimuths, $\theta$ (see Figure 5.9), the hotspot flow velocity at a point, $\tilde{V}_{\text{spot}}$, approaches the Keplerian disc velocity at that point, $\tilde{V}_{\text{disc}}$, as

$$\tilde{V}_{\text{spot}} = \tilde{V}_{\text{disc}} + (\tilde{V}_{\text{stream,0}} - \tilde{V}_{\text{disc}})(1 - \theta/\Delta \theta_{\text{spot}}).$$

This prescription means that the flow velocity of material in the hotspot region has reached the disc velocity at the end of the impact region. See Section 5.5.1 for an explanation of the parameters not described here. The local emission line profiles are treated as gaussians. The local disc line width is taken as the thermal velocity, $V_{\text{th}}$, assumed equal to the local sound speed, $c_s$, calculated using Equation 1.6. The hotspot is also assumed to be thermally broadened, using the hotspot temperature 10660K measured in Section 7.4.1. We use an outer disc radius of $r_{\text{disc}} = 0.36a$, and inner radius of 0.05$a$. The hotspot height is $h_{\text{spot}} = 0.013a$ and the disc thickness $2H_{\text{disc}} = 0.018a$ as in Section 5.5.1. We do not account for the emission line anisotropy due to the disc shear and turbulence (described in Horne & Marsh (1986) and Horne (1995)). This is a simplification of the
likely behaviour of the stream-disc impact. Detailed numerical simulation of the stream-
disc impact (Armitage & Livio 1998) shows that the velocity field varies greatly with height
above the disc, with the stream being stopped rapidly at the disc edge close to the disc
mid-plane, and with the stream almost uninterrupted by the disc a few scale heights above
the mid-plane. Making direct predictions of observed line profiles from such simulations
will be valuable in the future, but our simple model is sufficiently detailed to explain
the basic structure of these low resolution observations. To reproduce the weakened disc
emission around phase 0.3, we extended the spot radius to 0.05a, with the spot path
centred on a circle of radius 0.9r_{disc} (the dotted circle in Figure 7.14). The code can check
for occultation of the disc by both the donor and the hotspot. Any parts of the hotspot
behind or underneath the disc surface (i.e. within r_{disc} and below H_{disc}) are assumed to
be completely obscured. No attempt was made to model the absorption effects seen in the
observations, but these do not affect the basic structure we are reproducing.

The model trailed spectra are shown in Figure 7.15. The model trails were convolved
with a gaussian of FWHM 230 km s^{-1} to simulate the instrumental resolution. Simulated
line profiles were produced with two different settings. Trail (a) was produced allowing
the hotspot to occult the disc emission, so that any emission from the disc whose line of
sight passes through the stream-disc impact region or donor is not seen. The second trail,
(b), considers only eclipse by the donor, so we are simply seeing emission from the entire
disc and the hotspot added together. (c) shows the difference between the two models,
trail (b) - trail (a). Underneath each trail is a Fourier filtered back-projection, produced
as for the real observations by omitting the eclipsing phases. The relative brightness of
the disc and hotspot were chosen to match these model Doppler maps with the observed
H\alpha maps.

Unsurprisingly, both models reproduce the basic morphology of the observed trails:
double-peaked disc emission, centred on the white dwarf velocity, a deep eclipse of the
disc, beginning in the blue peak and ending in the red peak, and an S-wave from the
hotspot with maximum velocity around mid-eclipse and minimum velocity around phase
0.6. The region between the double peaks is not as dark as in the observations, but
this is mainly because the absorption effects were not modeled. The peaks of the disc
emission are sharper than observed, probably a result of not modeling the shear effects,
which lead to broader peaks and also to a deeper more V-shaped dip between the peaks
(Horne 1995). The apparently double-humped nature of the hotspot light curves is also
not seen. Such a variation would require the height of the stream-disc impact above the
disc to be similar to the spot radius, $0.05a$. The only difference between trails (a) and
(b) is that we see some weakening of the disc peak at about phase 0.2, just as seen in the
observations. There is a similar weaker region of the blue peak about phase 0.75. Indeed,
close inspection of trail (a) reveals an S-wave of absorption with the same amplitude as the
hotspot wave, but delayed by about 0.1 in phase. This was seen clearly in observations of
WZ Sge (Spruit & Rutten 1998). Trail (c), which shows (b)-(a), make this feature clear. It
corresponds to the region of disc emission under the stream-disc impact, and was modeled
in WZ Sge by Spruit & Rutten (1998) by removing disc emission from the stream-disc

![Figure 7.15: Model trailed spectra and Doppler maps.](image)
impact and alternatively by assuming absorption of hotspot emission by intervening disc material along the line of sight.

The model Doppler maps provide another, easier to interpret, way of studying the models. Doppler maps effectively produce the orbit-averaged velocity distribution, giving a better agreement between the model and observed maps than between model and observed trails, because the variations in hotspot flux are averaged out. The models match the Hα Doppler maps well. The model disc appears brighter around maximum and minimum $V_y$ than around $V_y = 0$, but this results directly from leaving out the eclipsing phases. The hotspot appears very much like that in the observations. The significant difference between the models is the region between the hotspot and the donor. As expected, this faint region of disc emission only appears when the stream-disc impact is allowed to obscure the disc: it is seen in map (a) but not map (b). The third map is just map (b) - map (a), identifying clearly in velocity space the region being obscured by the stream-disc impact. With the values of $h_{\text{spot}} - H_{\text{disc}}$ and $r_{\text{spot}}$ used in the model, most of the obscured emission is that from the regions of the disc under the hotspot, rather than the regions around the hotspot which are obscured at different phases. For higher values of $h_{\text{spot}}$, larger regions of the disc around the hotspot will be obscured. However, for this disc radius with the hotspot at $0.9r_{\text{disc}}$, a hotspot with $h_{\text{spot}} > 0.018a$ would eclipse the white dwarf. Increasing $h_{\text{spot}}$ to $0.018a$ has no significant effect on the model trails or maps. If the disc is not flat, for example increasing in thickness linearly with distance from the white dwarf, then the effect of the hotspot obscuring nearby disc regions would be increased, possibly enabling a lower value of $r_{\text{spot}}$ to be used while still reproducing the hotspot disc obscuration effects. This was tested by using a flared disc whose height, $h$, increases linearly with distance, $r$, from the white dwarf: $h = H_{\text{disc}}r/r_{\text{disc}}$. The strength of the obscuration$^1$ was increased by about 20 percent without any notable change in the model trails or maps.

The result of this modeling is to confirm the identification of the disc, hotspot and obscuration of the disc by the hotspot using a simple 3-dimensional model for the accretion flow and emission characteristics. With higher resolution higher S/N observations, and more computing power, this 3D modeling technique could be used to develop a more detailed model of the stream-disc impact, employing $\chi^2$ minimization techniques to determine best fit values of parameters such as disc radius and hotspot location.

$^1$The strength of the obscuration was quantified as maximum pixel flux in map (c).
7.7 Doppler tomography

7.7.2 January 2001

Finally, we study the detailed velocity structure and accretion flow behaviour during the dramatic change from quiescence to normal outburst in January 2001.

Trailed spectra

Trailed spectra for Hα, Hβ, Hγ and HeI 5876Å are shown in Figure 7.16. The 3rd and 4th January trails have 25 phase bins per orbit, while those from the 6th and 7th have 35 bins per orbit, reflecting the higher time resolution of these observations. All have 40 velocity bins in the range -2000 to 2000 km s⁻¹.

The Hα trail for 3rd January shows the two disc peaks, with the blue-to-red eclipse. The S-wave from the hotspot is not visible, however in the first orbit, the bright regions where the S-wave crosses the disc velocity and/or is brightest are seen around phases 0.3–0.5 (blue peak) and 0.6–0.8 (blue moving to red), along with evidence of the S-wave moving red-blue around phase 0.25. The deeper core absorption in these spectra compared to the March 2000 observations, along with the much lower temporal resolution of the observations, explains why the S-wave cannot be seen crossing zero velocity in this trail. This trail is therefore consistent with those from March 2000. The low S/N in He I prevents any structure being seen in that trail, apart from the core absorption and possibly a bright region around the blue peak between about phase 0.6 and 0.75, probably due to the S-wave.

The January 4th Hα trail again shows clearly the disc peaks and the blue-red eclipse. As seen in the lightcurve, there is no significant change in brightness, apart from eclipse. Evidence of the bright spot S-wave can again be seen in the red peak from phase 0–0.25 and in the blue peak around phase 0.6 to 0.9. The deep core absorption again prevents us seeing the S-wave crossing low velocities. The He I is too noisy to see anything but the core absorption and hints of the brighter region before eclipse.

For 6th January, Figure 7.16 shows trails for Hα, Hβ, Hγ and HeI. All show the double-peaked disc emission. The structure looks the same in all four trails, except that the red peak in the He I is weaker than the blue peak, a result of the NaI absorption. The eclipse, seen most clearly in Hα and Hβ (the brightest lines) is clearly a typical blue-red eclipse, as expected for a prograde disc. Apart from the rise in brightness throughout the night, there is no other easily identifiable behaviour. The very deep absorption cores in the lines again hide any structure between the disc peaks.

Only on the 7th January, when the absorption cores have weakened substantially, are
we able to see more complicated structure which is stable throughout the night. Again, the behaviour is very much the same for all four lines, apart from the relative strengths of the various components. Double-peaked disc emission with a blue-red eclipse is seen, with the region between the peaks being shallowest for Hα, and getting deeper for the higher order Balmer lines and for He I, as seen in the average spectra (Figure 7.4).

In the Balmer lines there is a clear S-wave component, crossing from maximum redshift at about phase 0.25, to minimum redshift around phase 0.75. This S-wave disappears
between about phase -0.2 and 0.2, and has a velocity semi-amplitude in the range about 300–600 km s\(^{-1}\). These characteristics suggest that this S-wave originates from the inner face of the donor star, whose velocity is \(\approx 442\) km s\(^{-1}\) (P2000 orbital parameters), and whose phasing would be the same as this S-wave. In addition, the inner face of the donor is invisible in the phase range about -0.2 to 0.2 (the exact region depending on the area of the inner donor face which is emitting). This feature has also been seen repeatedly in IP Peg during outburst e.g. in Morales-Rueda, Marsh & Billington (2000). There is one other feature in the trails, suggestive of another S-wave. In the blue disc wing around phase 0.4, centred on a velocity of about -700 km s\(^{-1}\), we see a bright region which covers a phase range of less than 0.1, and a velocity range of more than 500 km s\(^{-1}\). This is clearly seen in all three orbits in H\(\alpha\) and H\(\beta\), and can also be seen more faintly in H\(\gamma\) and He\(I\). In the red wing at phase 0.9 we see a similar feature at velocity of about 700 km s\(^{-1}\). This looks like a distinct feature, unrelated to S-wave thought to come from the donor, and is very likely to come from the same source as the feature around -700 km s\(^{-1}\), with both sharp features being close to opposite extrema of the same S-wave. Doppler mapping is needed to help further disentangle this detailed structure. These features look very much like those modeled in Harlaftis & Marsh (1996), where they modeled the emission during outburst in OY Car, including a component from the donor star, and with the higher velocity S-wave coming from the accretion stream.

Figure 7.17 shows the trailed spectrum of He\(\Pi\) and the Bowen blend region binned into
50 bins of 100 km s\(^{-1}\), with zero velocity corresponding to the wavelength of the He\(\text{II}\) line. It clearly shows double-peaked disc emission in He\(\text{II}\), but the eclipses do not start in the blue and move to the red. This is likely just a result of the Bowen blend centred around velocity \(-3000\) km s\(^{-1}\) distorting the line profiles. There is no other identifiable structure in the eclipses, probably a result of the blending of the lines.

**Doppler maps for 3rd and 4th January**

Figure 7.18 shows Fourier filtered back projections of the 3rd and 4th January 2001 observations. As for the March 2000 Doppler maps, orbital phases -0.1 to 0.1 were omitted to prevent eclipse spectra from distorting the maps.

The 5 and 6-minute exposure times of the spectra form 3rd and 4th January respectively will lead to azimuthal blurring in the maps of about 20°. Combined with the limited number of spectra contributing to the maps (16 and 13 respectively), we should not draw any conclusion about the fine detail in these maps.

![Figure 7.18: Fourier-filtered back-projections for 3rd and 4th January 2001.](image)

The H\(\alpha\) map for 3rd January shows a clear ring of disc emission. The disc emission shows two bright regions on opposite sides of the ring. The orientation of the bright regions suggests that they come from the brighter spectra around phase 0.7. The assumption
of Doppler tomography that the brightness of all points in velocity space is constant throughout the orbit is broken by these observations, leading to this apparent azimuthal variation in the brightness of the disc ring. By dividing each line profile by the integrated flux under it (from the lightcurve in Figure 7.10), a new normalized Doppler map was produced for which the assumption of a non-varying velocity space emission distribution is better matched. This is shown in Figure 7.19. The ring in this map does not show such strong azimuthal variation in brightness, confirming that the varying ring brightness in the unnormalized map is indeed an artifact resulting from the varying brightness of the line. We can therefore only conclude from this map that we see the disc emission. There is a blob consistent with the location of the bright spot, but with so few spectra, it cannot be confirmed as such. The January 4th Ha map also shows a clear ring of disc emission, this time with a fairly uniform brightness (there are no strong variations in the line flux out of eclipse). There is a brighter region consistent with the hotspot, but again, due to the limited number of spectra, this cannot be reliably identified as such. The HeI maps for the 3rd and 4th, which come from the very noisy spectra shown in the previous section, both show a bright region at the expected hotspot location, with no other bright region on the map for the 4th. There is one other bright spot in the map for the 3rd, but it is fainter than the region at the hotspot, and is not coincident with any expected source of emission. The detection of emission from the hotspot velocity on both nights suggests that this is genuinely the stream-disc impact emission. Therefore, from these very limited observations, coming from the blue end of near IR spectra which were optimized to look for the donor, we have still detected the main features of the quiescent accretion flow, as seen in more detail in the March 2000 observations.

Figure 7.19: Normalized Ha back-projection for 3rd January 2001.
Doppler maps for 6th January: Spiral shocks?

The Doppler maps from 6th January, shown in Figure 7.20, present a very different picture. The Hα map shows a ring of disc emission which has two bright regions on opposite sides of the disc around maximum and minimum $V_y$. The same structure is seen in Hβ and Hγ. The He I map is consistent, although it also shows a bright blob of emission in the top half of the lower left quadrant of the disc. This extra emission region is identifiable but much weaker in the Balmer maps. The two-armed asymmetry is reminiscent of the tidally induced spiral shocks seen in the outbursts of several dwarf novae, most notably IP Peg (Steeghs, Harlaftis & Horne 1997).

Spiral shocks are of great interest because they provide a possible mechanism for the effective transport of angular momentum out through an accretion disc, essential to allow mass to flow in through the disc at rates high enough to explain inferred accretion rates in CVs (see Section 1.4). The observational picture of spiral waves in accretion discs is still very limited (Steeghs 2001), with most observations of the phenomenon being in IP Peg. Other systems which show evidence for spiral waves during outburst are EX Dra, SS Cyg and U Gem, dwarf novae with orbital periods 5.0, 6.6 and 4.2 hours respectively, and V347 Pup, a nova-like with period 5.6 hours (Joergens, Spruit & Rutten 2000; Martinez-Pais et al. 1994a; Martinez-Pais et al. 1994b; Groot 2001; Still, Buckley & Garlick 1998). No detection of spiral waves has yet been made in an SU UMa system, as noted in the comprehensive observational review of spiral waves in accretion discs in Steeghs (2001). It is not clear whether this is a result of selection effects - spiral waves are expected only during outburst when the disc radius is large enough for the tidal effect of the donor to be important, and outbursts in the shorter period SU UMa systems are usually short -
or because spiral waves simply do not occur in the discs of short period systems. Truss et al. (2000) found spiral waves in SPH superoutburst simulations of the short period system Z Cha. This possible detection of spiral waves in the disc of IY UMa would be the first observation of spiral waves in a short period system, and as such confirmation of this result is important.

Figure 7.21 shows spiral waves in Doppler maps of IP Peg 5 and 6 days after the start of an outburst in August 1994 (from Morales-Rueda, Marsh & Billington (2000)). It shows clearly the two-armed spiral wave and a faint blob of emission from the irradiated donor. Of all the Doppler maps of spiral waves in CVs, the orientation of the waves in this image is closest to that of the asymmetry in our IY UMa maps, although even in these maps, the spirals appear rotated clockwise from the features seen in IY UMa. In other maps, both in IP Peg and other systems, the bright spiral arms are rotated clockwise by up to about 45 degrees. It seems unlikely that the behaviour of the spiral waves will not change, particularly with mass ratio, since the tidal forces inducing the spiral structure and the geometry of the system are sensitive to $q$. Armitage & Murray (1998) carried out SPH simulations of the disc in IP Peg, producing spiral waves which convincingly reproduce the structures seen in the IP Peg maps, and predicted that in low mass ratio systems ($q < 1/4$), asymmetries would be excited which move in the corotating frame. This complicates matters, making it seem unjustified to assume that any two-armed structure in low mass ratio systems will have the same geometry as in systems with higher $q$. 
The orientation of the possible spiral waves in IY UMa suggests that they could simply be an artifact of excluding the spectra during eclipse. Double-peaked spectra around phase zero contribute flux exactly where the disc is faintest in our maps. To test this, the line profile modeling code in the last section was used to produce trailed spectra of a circular accretion disc using the parameters described in the last section, and a radius of 0.45a to roughly match the velocity radius seen in these observed maps for the 6th January. The hotspot was not included, and the line profiles were produced for the same orbital phases as the observations. Figure 7.22 shows the Doppler map produced from these line profiles. We do indeed see bright regions around the top of the disc, consistent with those seen in the observations, but covering a wider azimuthal range. The second panel in Figure 7.22 shows the non-axisymmetric component of the model map, produced by subtracting the symmetric component from the image. The symmetric component of each image was calculated by replacing each pixel by the median average value of the map at the same distance from the white dwarf velocity as that pixel, effectively averaging the Doppler map azimuthally about the white dwarf velocity. Figure 7.23 shows the non-axisymmetric components of the observed Doppler maps. We see clearly that the azimuthal extent of the bright regions of disc emission is larger in the model than in the observations, suggesting that the omission of eclipse spectra may not be the cause of the bright disc regions.

The maximum brightness of the radially averaged flux in the Doppler maps was mea-
7.7 Doppler tomography

sured in two azimuth ranges for each map. Defining $\theta$ as the angle measured clockwise about the white dwarf velocity from the positive $V_y$ axis, the azimuth ranges were from -18 to 18° and 162 to 198° covering the bright regions of the disc, with 54 to 126° and 234 to 306° for the faint regions around $V_y = 0$. Using these averages, the ratio of the flux in the bright region to that in the weak region was found to be 1.9, 2.8, 2.7 and 1.8 in the Hα, Hβ, Hγ and He I maps respectively, while in the model it was found to be 1.8. So while this ratio is nearly the same in Hα, He I and the model, the ratio in Hβ and Hγ is much larger, telling us that it is unlikely that the apparent non-axisymmetry of the Doppler maps results from the missing eclipse spectra. We also note that the eclipse spectra were omitted in producing all the Doppler maps shown in this chapter, yet none of the other maps shows a disc with this brightness variation, possibly due to variations in the brightness of the disc out of eclipse.

![Figure 7.22: Doppler map and non-axisymmetric component of eclipsed circular disc model.](image)

![Figure 7.23: Non-axisymmetric component of the January 6 Doppler maps.](image)
Normalized Doppler maps were produced by dividing each line profile by the flux underneath, as for the 3rd January data. The normalized Hα map showed the same structure as in Figure 7.20, but the normalized Hβ and Hγ maps became a mess because the poor quality of the lightcurves for these weaker lines introduced significant artifacts.

Figure 7.24 shows a phase-folded trailed spectrum of Hα (left panel) for the 6th January and a trail reconstructed from the corresponding Doppler map (right panel). The eclipses are not shown, as they were excluded when producing the Doppler map. The white strip around phase 0.5 is coverage missing from the observations. The main difference between the observed and reconstructed trails is around this phase, where the reconstruction predicts significant weakening in the double peaks and the filling in of the region between them. This is not seen in the observations. We therefore cannot be certain that the structure in the Doppler maps is genuine, noting that at this time of rapid change in the accretion flow, the effectiveness of Doppler tomography is compromised.

![Phase-folded observed and reconstructed trailed spectra of Hα for 6th January.](image)

We may only conclude that this two-armed asymmetry in the Doppler maps may be evidence of spiral shocks at the peak of outburst, and higher quality observations are essential to confirm this behaviour.

The non-axisymmetric components of the Doppler maps, shown in Figure 7.23, reveal emission from around the donor Roche lobe which was not clear in the maps without the symmetric component removed. We discuss this further in the next section.
In Figure 7.25 we see the Doppler maps for 7th January. Again, we see the behaviour has changed drastically in one day. In all four maps we see strong emission concentrated around the velocity of the donor star. This emission region corresponds to the S-wave observed in the trails which is eclipsed between phases -0.2 and 0.2, suggesting that it comes from the inner face of the donor star. We also detected emission from the donor star velocity in the non-axisymmetric components of the Doppler maps from January 6 (Figure 7.23). Given the resolution of observations, the position in the Doppler maps of this source is consistent with emission from the inner face of the donor. Such emission from the donor star during outburst has been seen before, e.g. in OY Car (Harlaftis & Marsh 1996) and IP Peg in many different observations (Steeghs (2001) and Figure 7.21 from Morales-Rueda, Marsh & Billington (2000)). In each case the donor emission has been ascribed to irradiation of the inner face of the donor by high energy radiation from the boundary layer. Evidence for the presence of EUV emission and X-ray emission from the boundary layer was provided by the detection of the He II 4686Å emission line, believed to be produced through reprocessing of EUV and X-ray emission. He II 4686Å is seen in our spectra on both the 6th and 7th January, and the double peaks (Figures 7.5 and 7.17) tell us that it is coming from the disc. Marsh & Horne (1990) concluded that reprocessing of radiation from the boundary layer is the most likely cause of the Balmer emission from the donor in IP Peg, and Harlaftis & Marsh (1996) argued that with a flux of high energy radiation sufficient to produce significant reprocessing in the disc in OY Car, irradiation of the donor star is also possible. This argument applies in IY UMa, particularly when we
note that the strength of He\textsc{ii} relative to the Balmer series on 6th January was similar to that seen in OY Car, while on the 7th January it was much greater, explaining the stronger emission from the donor on the 7th than on the 6th.

The other notable feature seen in these Doppler maps is an arc of emission on the left hand side (negative $V_x$) at velocities corresponding to the outer part of the disc. We do not see the full circle corresponding to the whole disc. The arcs are brightest around $V_y = 0$. In H\textalpha\ and H\beta, the arc stretches through about 270° (measured anti-clockwise from the donor velocity), while in H\gamma and He\textsc{i} it only reaches about 135°, although the He\textsc{i} map shows an isolated bright blob at about 225°. At a first glance, this structure appears similar to that seen in outburst in OY Car by Harlaftis & Marsh (1996), which followed the accretion stream, and was thus attributed to emission from the stream. However, in IY UMa this is not the case. The stream trajectory should follow the arc marked with open circles. As it merges with the disc the stream-disc impact should be seen between the stream trajectory arc and the arc showing the Keplerian velocity along the stream trajectory (marked with filled circles). The arc of emission does not follow the stream trajectory in any of the Doppler maps, except where it first leaves the L1 point. It very quickly follows the velocity of the outer disc. Noting that no stream-disc impact hotspot is seen, we see that this emission does not arise from dissipation of energy where the stream hits the disc edge. Such emission, seen in quiescence, is small compared to this line emission in outburst. The fact that this emission is seen predominantly at the disc edge in the same quadrant of the disc as the stream-disc impact suggests a possible interpretation. We have already seen that irradiation by high energy emission from the boundary layer (BL) is likely to be the source of Balmer emission on the inner face of the donor and He\textsc{ii} emission in the disc. This arc of emission could also result from reprocessing of this BL emission in the disc. The extent of reprocessing would be greater for regions of the disc with a greater vertical extent, simply because such regions would intercept more of the BL radiation. Despite that lack of emission from dissipation of kinetic energy at the stream-disc impact, some vertically extended structure is likely to be present where the stream impacts the disc, providing a region where we might see reprocessing of the EUV and X-rays from the BL. This region would be in the top left of the Doppler maps. If this vertical structure, raised at this point initially by the stream disc impact, extended further around the disc, then it could explain the brightest region of the emission line arcs seen here. Simulations of discs and the stream-disc impact do show vertical structure extended
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around the disc edge e.g. Armitage & Livio (1996) and Hirose, Osaki & Mineshige (1991). There is also much observational evidence for vertical structure extending around the disc rim. Billington et al. (1996) suggested variable vertical structure in the outer edge of the disc in OY Car during superoutburst as an explanation for dips in its UV flux which were coincident with superhumps. We already expect some reprocessing of the light from the BL throughout the disc, which could explain why the the ring extends most, if not all, of the way around the disc in the Hα and Hβ maps. Studies of accretion discs in outburst suggest the presence of flared discs whose vertical structure is variable and probably greatest at the height of outburst. Such evidence has been seen in HT Cas (Ioannou et al. 1999) and in OY Car (Naylor et al. 1987). Combined with additional vertical structure, near to and perhaps triggered by the stream-disc impact, reprocessing of EUV and X-ray emission from the BL could explain the structure seen in these maps.

The delay in the rise of the line emission compared to the rise of the continuum might also be understood if the line emission in outburst is powered by irradiation from the BL (as suggested by Marsh & Horne (1990)) and the outburst is of the outside-in type seen in OY Car (Vogt 1983a) and HT Cas (Ioannou et al. 1999). The continuum flux begins to rise as soon as the outer disc enters the high state, while the emission lines do not become significantly strongly until the heating wave has reached the inner disc, because the emission line flux follows the accretion rate through the BL.

7.8 Conclusions

In this Chapter we have looked at a variety of properties of IY UMa in quiescence and in outburst. In quiescence we have seen some behaviour which is typical of other similar high inclination extreme mass ratio SU UMa systems, particularly OY Car and Z Cha. In quiescence, we found the most significant difference between IY UMa and these systems to be associated with the stream-disc impact, notably the lack of emission lines from the hotspot in OY Car and Z Cha contrasted with a pronounced S-wave in the emission line profiles of IY UMa originating form the hotspot. We have seen evidence of the hotspot disrupting and obscuring the nearby accretion disc, producing a simple model for this behaviour which with higher quality observations is likely to lead to a better understanding of the stream-disc impact. The importance of the stream-disc impact in IY UMa, and the importance of IY UMa in understanding the stream-disc impact, is a direct result of the extremely high inclination of this system. We are seeing the disc almost edge on.
7.8 Conclusions

We were fortunate enough to observe IY UMa just before the rise to an outburst, and for two days after the outburst peak. Spectroscopy just before the rise to outburst looks much like that in quiescence a year earlier. Around the peak of the outburst, we detect possible evidence of two-armed asymmetry in the accretion disc, something which might be attributed to supersonic spiral shock waves induced by the tidal influence of the donor. Such shock waves have been seen before, predominantly observed in IP Peg, and maybe of importance in understanding the fundamental question of angular momentum transport in accretion discs. If interpretation of these structures in the IY UMa Doppler maps as spiral waves is correct, this would be the first detection in an SU UMa type CV. Confirmation of this result is important, but the unpredictability of outburst times makes planning new observations troublesome.

Only a day later, while still in outburst, the behaviour of IY UMa was different again, with strong emission from the donor star and non-axisymmetric emission in the outer disc. The donor emission is likely to result from reprocessing of high energy emission from the boundary layer, while vertical structure in the disc, possibly affected by the stream-disc impact, may also be reprocessing BL radiation to explain the outer disc emission.

Higher resolution observations of IY UMa promise to yield more valuable information about accretion flow, particularly about the shocked region where the accretion stream impacts the disc edge.
Chapter 8

Conclusions

The original project for this thesis was to use the January 1995 simultaneous photometry and spectroscopy of V348 Pup, combining the spatial information provided by the eclipses with the velocity information in the line profiles, to map in detail the kinematics of the precessing non-axisymmetric disc suspected to be causing the persistent superhumps seen in the system. The discovery that the line emission does not come from the disc made this impossible, but while compromising the original project to some extent, a detailed study of two radically different accretion phenomena was facilitated.

Combining the 1995 photometry of V348 Pup with the two other datasets from 1991 and 1993 provided a set of observations with complete coverage of the precession cycle. Modeling and mapping of the spatial intensity distribution in Chapter 3 reveals the presence of a progradely precessing non-axisymmetric accretion disc, interacting with the orbit of the donor to produce the superhump modulation. The eclipse profiles and orbital lightcurves suggest that the mechanism by which the superhump emission is produced is the varying dissipation of energy at the stream-disc impact as the converging flows of the outer disc and the accretion stream are affected by the changing orientation of the disc. ‘Common’ superhumps seen during the superoutbursts of SU UMa type dwarf novae result from the varying viscous dissipation in the accretion disc as it is tidally distorted by the gravitational influence of the donor star, and we might have expected persistent superhumps in the nova-like V348 Pup, which is effectively in permanent outburst, to arise from the same mechanism. However, since SU UMa systems in superoutburst never reach a steady state, it is also possible that persistent superhumps represent the behaviour which would
be seen in SU UMas if they had time to reach a steady state during superoutburst. It would then be unsurprising that persistent superhumps are more like the late superhumps (seen when a system has spent a long time in superoutburst) than they are like common superhumps. Further examination of this result is important, and a similar study of the nova-like UU Aqr is planned, using observations in hand. Simulated eclipse profiles from particle simulations of superhumping nova-likes should reveal whether it is possible that the complicated varying disc shapes and viscous dissipation patterns in tidally distorted discs could conspire to look like a superhump arising from the hotspot.

The role of the stream-disc impact in the superhump phenomenon has been suggested before, but it has only found favour as the likely source of 'late' superhumps, seen in SU UMa systems as they approach the quiescent state after a superoutburst. In Chapter 5 the deep eclipse and excellent view of the stream-disc impact in IY UMa made it possible to develop a simple model for the three-dimensional structure of, and energy dissipation at, the stream-disc impact. Combined with the eccentric disc shape determined from changing eclipse profiles, this model is able to convincingly reproduce the lightcurves of IY UMa as it exhibits late superhumps, confirming the hotspot as the source of late superhumps in this system.

Low resolution time-resolved spectroscopy of IY UMa in Chapter 7 reveals further signatures of the stream-disc impact, similar to the behaviour seen in several other CVs. These observations demonstrate how a higher resolution spectroscopic study would provide an excellent opportunity to test the theoretical models of this shocked region of converging accretion flows. Further spectroscopy reveals the dramatic changes in accretion behaviour as IY UMa rises from quiescence to a normal outburst. We see the system at the peak of the outburst possibly showing structures reminiscent of the spiral waves seen in the accretion discs of several longer period systems (Steeghs 2001), but never before seen in an SU UMa system, a result which could have implications for the transport of angular momentum in accretion discs, fundamental to accretion astrophysics. It should be noted, however, that all the reliable observations of spiral waves in accretions disc observed so far have been in emission lines, and therefore only tell us about the structure of the disc chromosphere. How this translates to the behaviour deeper in the disc is unclear, but could be of importance when considering the angular momentum budget of the disc. A day later, the system shows evidence of reprocessing on the donor star and in vertically extended structure in the outer disc of high energy radiation from the boundary layer where the accretion disc
meets the white dwarf. These observations will provide a valuable comparison for the rapidly-improving models of dwarf nova outbursts. Spectroscopic detection of the donor star in IY UMa (Chapter 6) promises the possibility of tighter direct constraints on the system parameters, making IY UMa with its deep eclipses likely to become one of the most valuable laboratories for studying accretion flow.

The spectroscopy of V348 Pup (Chapter 4) reveals it is another member of the SW Sex class of nova-like CVs, whose anomalous emission line behaviour has yet to be fully understood, with two popular models competing to provide an explanation. Simulation of the accretion flow using the disc-anchored magnetic propeller of Horne (1999) confirms that this model can explain most of the observed features of the SW Sex systems, including one unique to V348 Pup, where the transient blue-shifted tramline absorption briefly changes its velocity, width and depth before disappearing.

The work in this thesis has exploited both modeling and indirect imaging techniques to find simple, self-consistent, physically-based explanations for detailed phenomena observed in two cataclysmic variables.
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