ROSAT and HST Observations
of Magnetic Cataclysmic Variables.

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This thesis describes principally ROSAT and HST observations of magnetic (or proposed magnetic) cataclysmic variables. Chapter 2 details a ROSAT observation of the polar EK Uma. The orbital light curve reveals a single bright phase. During this bright phase deep dips in the flux are seen, consistent with accretion stream occultation. The soft X-ray spectrum has an unusually high temperature of $50 < kT_{BB} < 62$ eV. No hard X-ray flux is detected.

Chapter 3 details ROSAT observations of proposed intermediate polars SW Uma and 1H0709-360. The previously reported X-ray/optical periodicities in SW Uma were not detected. Spectral analysis indicates a two-temperature model is appropriate. The weak signal from 1H0709-360 precludes a detailed spectral analysis. 1H0709-360 has dropped in flux by $\sim 2$ orders of magnitude since its detection. The intermediate polar classification of these two systems remains unconfirmed.

Chapter 4 describes ROSAT and contemporaneous optical and HST observations of the intermediate polar AE Aqu. During this observation X-ray flares were detected for the first time. The white dwarf spin modulated count rate increased only slightly with increased intensity. AE Aqu has an unusually soft spectrum which is only fit by a two temperature optically thin plasma emission model.

Chapters 5 and 6 detail ROSAT, HST, EUVE and complementary optical observations of the polar QS Tel. The soft X-ray/EUV data showed a bright-faint morphology. The EUVE observations, $\sim 1$ year later, revealed a change in morphology, indicating that both poles were active. A deep narrow dip is observed in both the ROSAT and EUVE observations. The HST (POS) observation had a mean spectrum of an underlying continuum with strong emission lines superposed. Orbital modulation was present in both. A narrow dip is observed in the continuum fold which is consistent with the dip observed by ROSAT and EUVE.
Declaration

I hereby declare that no part of this thesis has been previously submitted to this or any other University as part of the requirement for a higher degree. The work described herein was conducted by the undersigned except for contributions from colleagues as acknowledged in the text.

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Publications

Clayton K.L., Osborne J.P.
'The Soft X-ray Properties of the Polar EK UMa'

Rosen S.R., Clayton K.L., Osborne J.P.
'ROSAT Constraints on the Intermediate Polar Candidates V426 Oph, SW UMa and 1H0709-360'

'ROSAT Observations of AN UMa and MR Ser; the Status of the Soft X-ray Excess in AM Her Stars'

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'A ROSAT Observation of the Peculiar Magnetic Cataclysmic Variable- AE Aqr'

Clayton K.L., Osborne J.P.
'ROSAT Observations of AE Aqr made during the Whole Earth Telescope and World Astronomy Day Campaign'

'A ROSAT Observation of the Polar RE1938-461 (QS Tel)'

'Simultaneous X-ray, UV and Optical Observations of Flaring in AE Aqr'

Eracleous M., Horne K., Osborne J.P., Clayton K.L.
'HST Observations of AE Aqr during the 1993 WAD Campaign'

'Accretion Mode Changes in QS Tel (RE1938-461): EUVE, ROSAT and Optical Observations'

Clayton K.L., Rosen S.R., Osborne J.P.
'Recent Progress on QS Tel: The HST Results'
This thesis is dedicated, with love, to

Mum, Dad and Gill

Thank You
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Chapter 1

Introduction.

This thesis describes observations of five different cataclysmic variables, all of which are or are proposed to be magnetic systems (polars and intermediate polars). These observations are made principally in the soft X-ray band, EUV and UV with supporting optical photometry. This work includes analysis of ROSAT soft X-ray data of the most unusual cataclysmic variable yet observed - AE Aqr, to better determine the nature of the X-ray emission. The polars EK UMa and QS Tel are studied in the soft X-ray (ROSAT observation), EUVE and HST data are also reported in the case of QS Tel. Limits as to the nature of proposed intermediate polars SW UMa and 1H0709-360 are made.

This chapter describes the nature and properties of cataclysmic variables. Also included are details of the major emission and absorption mechanisms relevant to the observations made in the thesis.

1.1 Cataclysmic Variables

1.1.1 Overview

Cataclysmic variables (CVs) are semi-detached binary systems comprised of an accreting white dwarf star and a mass donating secondary, generally a late type star on, or near, the
main sequence. Orbital periods vary between ~ 80 min and a few days. The separation, \( a \), of the two stars, from Kepler's third law, is given by

\[
a = 3.5 \times 10^{10} M_1^{1/3} (1 + q)^{1/3} \frac{P}{24} \text{ cm}
\]  

(1.1)

where the masses are expressed in solar units and \( M_1 \) is the primary mass, \( q \) is the mass ratio \( M_2 / M_1 \) and \( P \) is in hours. Typical separations are such that the entire system could easily fit inside a star of the size of our sun.

In such closely bound systems it is possible for the outer layers of one star to be disrupted and drawn off by the gravitational pull of the companion. The potential felt by a test particle within such a system is approximated by Roche to be

\[
\Phi_R(r) = -\frac{G M_1 M_2}{|r - r_1|} - \frac{G M_2 M_1}{|r - r_2|} - \frac{1}{2} \omega \cdot r^2
\]  

(1.2)

where

\[
\omega = \left[ \frac{G (M_1 + M_2) M_2}{a^3} \right]^{1/2} \hat{u}
\]  

(1.3)

where \( \hat{u} \) is the unit vector normal to the orbital plane. This assumes that both the stars are point masses (this is a good approximation considering the central condensations found in stars) and are in circular orbits. Figure 1.1 shows some example Roche potentials for a system with a 3.3 hr orbital period and a mass ratio \( q = 0.62 \). The crosses indicate the centre of the two stars and the centre of mass is also indicated (CM). The potentials are plotted from the inside increasing outwards. Close to the stars the potential is essentially that of a single star, as you go further out however the potential becomes distorted until the 'figure of eight' or the Roche lobes are produced. The point where the two lobes meet is known as the inner Lagrangian point (labeled as \( L_1 \) in the figure). Material near this point can easily pass from one lobe to the other, as the potential at this point is less than the surrounding critical surface: an analogy of this is that the \( L_1 \) point is like a high
mountain pass between two valleys. Thus if the secondary fills its Roche lobe then mass transfer onto the companion can occur. This is commonly known as Roche lobe overflow.

Plavec and Kratochvil (1964) gave the distance from the $L_1$ point to the centre of the primary ($b_1$) as

$$\frac{b_1}{a} = 0.500 - 0.227 \log q$$

(1.4)

If this accreting material flows directly onto the primary then the ensuing luminosity is given by

$$L_{\text{acc}} = \frac{GM_1 \dot{M} M_0}{R_1}$$

(1.5)

or

$$L_{\text{acc}} = 1.3 \times 10^{39} \dot{M}_{16} M_1 (10^5 \text{ cm}/R_1) \text{ erg s}^{-1}$$

(1.6)

where $\dot{M}_{16}$ is the accretion rate in units of $10^{16} \text{ g s}^{-1}$. For typical accretion luminosities seen in CVs this can lead to accretion rates of $10^{16} \text{ g s}^{-1}$. This is much less than the limiting Eddington luminosity, where radiation pressure becomes sufficient to prevent the infall of further material, which is

$$L_{\text{edd}} \approx 1.3 \times 10^{39} M_1 \text{ erg s}^{-1}$$

(1.7)

The lowest temperature at which a body of radius $R_1$ can radiate at a luminosity $L_{\text{acc}}$ is the blackbody temperature ($T_{\text{bb}}$),

$$T_{\text{bb}} = \left( \frac{L_{\text{acc}}}{4 \pi R_1^2 \sigma} \right)^{1/4}$$

(1.8)
Figure 1.1: Roche lobe geometry. Roche equipotentials in the orbital plane are plotted for a system with $q = 0.62$ and $P_{\text{orb}} = 3.3$ hr. Crosses mark the centres of the two stars of mass $M_1$ and $M_2$. The centre of mass is indicated by CM and the solid equipotential which intersects at the inner Lagrangian point, $L_1$, defines the Roche lobes of the two stars.
where $\sigma$ is the Stefan-Boltzmann constant. The maximum temperature that can be achieved by the accreting gas is when the gravitational potential energy is converted entirely into thermal energy ($T_{\text{th}}$) or blackbody radiation,

$$T_{\text{th}} = \frac{GM_1M_2\sin^2 i}{3kR_1}$$ (1.9)

The temperature at which the accretion luminosity is radiated ($T_{\text{rad}}$) would be expected to lie between these two limits. Therefore for a $1M_\odot$ white dwarf accreting at a typical luminosity of $10^{33}$ erg s$^{-1}$ the expected range of temperatures from accretion would be

$$6 \text{ eV} < kT < 100 \text{ keV}$$

Hence the importance of studying CVs at X-ray wavelengths. These limits do not constrain the temperature of the accreting object very well, a more detailed examination of emission processes must be made to obtain better limits.

### 1.1.2 The white dwarf primary

The upper limit for the mass of a white dwarf was determined by Chandrasekhar (1931) to be $\sim 1.4M_\odot$. White dwarfs are one of the possible evolutionary endpoints after nuclear burning has been exhausted. Different initial masses of the progenitors will lead to a range of final compositions (Iben & Webbink 1989, Mazzitelli 1989, and references therein). Determining the mass of the white dwarf is rather difficult as, unlike the secondary, it does not fill its Roche lobe and hence has no volume constraint. Masses of both components can be calculated using Keplers third law if the radial velocity variations of both stars can be measured and the inclination of the system is known. The inclination is only well determined for eclipsing systems.

The mass-radius relationship is fairly well determined, and is dependent on the chemical composition of the white dwarf. An approximation of the analytical solution by Nauenberg (1972) is sufficient for the needs of this thesis and is given by
A feature of this relationship is that the radius of the white dwarf decreases with increasing mass. A 1 M⊙ white dwarf will have a radius of ~ 5 × 10^6 cm.

1.1.3 The secondary

The secondaries in CVs are mainly considered to be low mass late type main sequence stars, though there are exceptions, eg GK Per. Precise determination of their spectral classification is difficult as the bulk of the luminosity observed in CVs arises from the accretion process. Most determinations of the secondary type are made using red spectra where the accretion luminosity is much lower. In general for periods of < 0.25 days these stars have a red spectrum characteristic of M dwarfs, for longer periods K dwarfs have been observed (see Ritter & Kolb 1993, and references therein).

Eggleton (1983a) derived an expression for the secondary Roche lobe volume equivalent radius as

\[
\frac{R_L}{a} = \frac{0.49 q^{2/3}}{0.6 q^{2/3} + \ln(1 + q^{1/3})}
\]  

(1.11)

this is a better estimate of the Roche lobe size than others such as Paczyński (1971) as it is valid for all q and is accurate to 1%.

From the relations for the binary separation and \( R_L \), and requiring that the secondary does fill the Roche lobe, it can be shown that the density within the lobe is a function of the binary period (for q< 0.8). Therefore by using the empirical mass-radius relation for low mass main sequence stars of Patterson (1984),

\[
\frac{R}{R_\odot} = \left( \frac{M}{M_\odot} \right)^{0.88}
\]  

(1.12)
it is possible to determine the mass and radius of the secondary from just knowing the orbital period,

\[ M_2 = 0.071 \, P_{\text{orb}}^{0.22} \, M_\odot \]  
(1.13)

\[ R_2 = 0.098 \, P_{\text{orb}}^{0.074} \, R_\odot \]  
(1.14)

1.1.4 Classification of CVs

CVs can be subdivided into several different classes depending on several properties such as the magnetic state of the white dwarf or photometric behaviour. The main division lies between magnetic and non/weakly magnetic white dwarf CVs.

Non-magnetic systems

As material leaves the secondary via the L1 point, it possesses high angular momentum, this prevents the material accreting directly onto the white dwarf. Test particles within such a system follow trajectories that first pass closely to the white dwarf and then out again almost to the Roche lobe, before returning to intersect the path at some intermediate radius. A stream of gas would be expected to collide with itself and dissipate energy, eventually forming a circular orbit, using Kepler's law and equation 1.4 it can be shown that the circularization radius is,

\[ R_{\text{circ}} = a(1 + q)[0.500 - 0.227 \log q]^{1/4} \]  
(1.15)

For the values of q of interest in CVs (Rappaport Joss & Webbink (1982) showed that stable mass transfer can only occur for values of q< 2/3) it is found that \( R_{\text{circ}} \) is a factor of \( \sim 3 \) smaller than the Roche lobe radius, but is larger than the maximum radius of the
White dwarf (~ $10^9$ cm for a low mass white dwarf). Thus the stream forms a ring in a Keplerian orbit around the white dwarf.

However, this ring is unlikely to be stable. Turbulent motion and viscosity in the gas convert kinetic into thermal energy which is then radiated away. With the loss of this kinetic energy the material must fall deeper into the potential well, but to remove the angular momentum some material must move to larger radii. The overall effect of this is to smear out the ring and produce an accretion disc.

The binding energy of a gas element of mass $m$ in the Kepler orbit which just grazes the surface of the primary is $G M m / 2 R_*$. Since the gas elements start at large distances from the star with negligible binding energy, the total disc luminosity in a steady state must be

$$L_{\text{disc}} = \frac{G M \dot{M}}{2 R_*} = \frac{1}{2} L_{\text{acc}}$$

(1.16)

where $\dot{M}$ is the accretion rate and $L_{\text{acc}}$ is the accretion luminosity (1.5).

The balance of the accretion energy is expected to be deposited in the boundary layer (BL), i.e.,

$$L_{\text{BL}} = L_{\text{disc}} \left(1 - \left[\frac{\Omega_*}{\Omega_{R(WD)}}\right]^2\right)$$

(1.17)

where $\Omega_*$ is the white dwarf angular velocity, and $\Omega_{R(WD)}$ is the angular velocity of the material at the white dwarf radius. Thus for a slowly rotating white dwarf, the BL luminosity should be approximately equal to the disk luminosity. The temperature of an optically thick BL is expected to be of order $10^5 - 10^6$ K (Pringle 1977).

Where material first impinges on the disc a 'bright spot' is formed (see fig 1.2). This spot presents a varying aspect during the orbital cycle, thus producing orbital period modulation. This can be seen as in some cases as an 'S-wave'. The bright spot is hot compared to the outer edge of the disk. While sometimes prominent at optical wavelengths the bright spot is 'hot' only relative to the quiescent outer disk temperature, e.g. in OY Car.
Figure 1.2: Schematic of a dwarf nova, viewed from above. The gas stream flowing from the Roche lobe filling secondary, via the L₁ point, impinges on the accretion disc creating the 'bright spot'. Material flows through the disc until eventually, when it has lost enough angular momentum, it is accreted onto the white dwarf.

(Wood et al. 1989). During outbursts the disk becomes relatively brighter and the spot is not discernible.

Thus, in steady-disc theory, the disc makes a contribution to all parts of the spectrum between the near infra-red and the far ultra-violet, while the BL can contribute approximately an equal amount of radiation in the EUV/soft X-ray spectral region. Other sources of luminosity in the system are the companion star, which contributes principally in the infra-red, the optical bright spot and various non-steady state sources like winds.

Outbursts are a common feature observed in non-magnetic CVs and are caused by enhanced accretion onto the white dwarf. Two models exist to explain this increased mass transfer rate through the disc. First: an instability in the red dwarf causes the mass transfer through the L₁ point to increase, and subsequently through the disc and onto the white dwarf (Bath 1984, Bath, Clarke & Mantle 1986). Second: the mass loss rate from the secondary is constant and the mass transfer rate increase is caused by an instability within the accretion disc (Smak 1984; Frank, King & Raine 1992). Both models have had successes and failures in predicting the observed outbursts (Córdova 1993). In non-disc CVs (polars) high and low states have been observed showing that mass transfer from the
secondary can change, and it is likely that this can occur in dwarf novae. However simple mass transfer burst models predict that the bright spot luminosity should significantly increase during the outburst, which is not observed (Meyer-Hofmeister & Ritter 1993), though more detailed calculations by Dgani, Livio & Soker (1989) show that this effect is reduced as there is some degree of stream penetration into the disc. The disc instability models predicts that the bright spot luminosity, if anything, should decrease during the outburst. This model can also account for the different classes of non-magnetic CVs (Meyer-Hofmeister & Ritter 1993). However this model can not account for the rise in UV flux 0.5-1 days after the optical. It does also predict that there should be a gradual rise in the interoutburst level, which is not observed (Córdova 1993; Meyer-Hofmeister & Ritter 1993).

Non-magnetic systems can be sub-divided into many groups based principally on their photometric behaviour, the most common of which are the dwarf novae. Table 1.1 shows the range of systems observed.

U Geminorum was the first system to be classed as a dwarf nova in the latter half of last century as a result of its observed optical variability (Hind 1856 and Pogson 1883). The variability of dwarf novae was characterised by recurrent outbursts in which the source brightens by 2-5 magnitudes for a period typically of a few days before returning to its quiescent state. The outbursts recur on timescales ranging from tens-hundreds of days depending on the object. With binary periods of a few hours and one component thought
to be a late type dwarf the dwarf novae were classed as a subset of cataclysmic variables.

The brightest members of the class have $m_v \sim 8 - 9$ while in outburst. As a result of different outburst morphologies dwarf novae are sub-divided into three sub-groups (see table 1.1).

The U Gem variables are characterised by outbursts of similar total luminosity. There are however different types of outburst commonly exhibited by this class. Several different schemes of classification of outbursts exist. The simplest has only three categories, 'short', 'long' and 'anomalous'. The short and long outbursts tend to alternate. The anomalous outbursts are different because they occur infrequently and typically have a long duration, which is normally also followed by longer periods of quiescence.

The difference between short and long outbursts gets more pronounced as the orbital period gets shorter. For dwarf novae with periods of less than $\sim 2$ hours the long outbursts are typically a magnitude brighter than the normal outbursts and of longer duration. These are called super-outbursts, and have a total outburst energy which is an order of magnitude greater than the short type. The tendency for the alteration of the outburst disappears and super-outbursts occur less frequently. Systems exhibiting this type of behaviour are known as SU UMa variables. Another feature of these SU UMa systems (in contrast to U Gem systems) is that the optical emission during the super-outburst is smoothly modulated with a amplitude of $\sim 0.3$ mags at a period which is a few percent longer than the orbital period. This modulation is referred to as the 'superhump' behaviour. These superhumps have been shown to be not a function of inclination. Patterson et al. (1983) found that this behaviour is common amongst CVs (not just dwarf novae) of high mass accretion rate and short orbital period.

The final subgroup are the Z Cam variables. These exhibit a period of 'standstill' where the source brightness remains fixed at an intermediate intensity level for an extended period of time. Z Cam itself, for example, has an outburst recurrence time of 23 days but has been observed in standstill for up to about eighteen months. The standstills always start during the decline from outburst and end with a return to quiescence.

Spectra of dwarf novae during quiescence are characterised by strong, sometimes broad,
emission lines of H, He I and sometimes He II, with a flat or blue continuum. In outburst the continuum becomes bluer and the neutral hydrogen and helium lines tend to become broad absorption features.

Other non-magnetic systems are:

- Classical novae. These exhibit a single outburst with an increase of brightness of \( \sim 10 \text{ mag} \) in a few days, and decrease to quiescent level over a period of months to years. The spectrum changes during the outburst decline from outburst, passing through a 'post-novae' stage which is similar to dwarf novae in outburst, to an 'old-novae' state which is similar to dwarf novae in quiescence. The cause of the outburst is different from that seen in dwarf novae. As hydrogen-rich material accretes onto the white dwarf surface a layer of unburnt nuclear fuel builds up. The temperature and density at the base of this layer rise gradually until nuclear reactions can begin. The temperature and luminosity run away until the Eddington luminosity limit is reached, and thus the outburst is produced (Bath & Pringle 1985).

- Recurrent Novae. These show large amplitude outbursts similar to those seen in classical novae but these recur on a timescale of a few tens of years. If these outbursts are produced by the same mechanism as that seen in classical novae, this implies that either material must be being accreted onto the primary at a higher rate, or less material is required in these systems (Wade & Ward 1985).

- 'Nova-like' variables. These have similar spectral properties to novae, but have not been observed undergoing outburst. This class also includes the polars. The UX UMa class have a low level variability (\( \sim 0.5 \text{ mag} \)), and appear as dwarf novae in a permanent bright state. The VY Scl types have been termed 'anti-dwarf-novae', and as this suggests, they are usually in what appears to be dwarf nova outburst state, but show occasional dips into quiescence.
Magnetic CVs

The presence of a magnetic field with a magnetic moment $\mu > 10^{32} \text{G cm}^3$ (Lamb & Melia 1989) on the white dwarf is thought to disrupt the accretion process by channeling material onto the field lines and accreting material near to the polar regions. This results in pulsed emission at the spin period of the white dwarf. Several different accretion geometries can be seen depending on the strength of the magnetic field.

For the strongest fields ($\mu > 10^{33} - 10^{34} \text{ G cm}^3$) accretion discs are completely disrupted (see fig 1.3)(Lamb & Melia 1989) and the emission from the system is dominated by the funneled accretion flow. Because of the strong magnetic field the white dwarf is rotating synchronously with the orbit in these systems. They are known as polars or AM Herculis systems and are the subject of chapters 2, 4 & 5, I shall give more detail later in the sections following.

For lower magnetic fields the situation is more complex (for further details see section 1.1.8), and depends on the orbital period, the spin of the white dwarf and the magnetic field strength. The common view has the inner portions of the accretion disc disrupted by the magnetic field funnelling material onto the polar regions (see fig 1.4). The observed X-ray emission is a combination of that from the disc and the pulsed polar emission at the spin period of the white dwarf. These systems are known as intermediate polars (IPs) or DQ Herculis systems, (see Patterson 1994 for a review). Two suspected IPs are the subject of chapter 3 and the most unusual of these systems is examined in chapter 4.

It is also believed that fields of $10^4 \text{ G} (\mu \sim 10^{33} \text{ G cm}^3)$ may result in some tiny disruption of discs (Livio & Pringle 1992). This is thought to produce the UV delay seen during some dwarf novae outbursts (in the context of the disc instability model). Although the field is not strong enough to channel the accretion flow it is sufficient to evacuate the inner portions of the disc during quiescence. At the onset of an outburst there is a delay before the inner disc regions are filled and can emit in the UV. Warner (1983) suggested that a weak magnetic field may produce quasi periodic oscillations (QFO) which are observed in dwarf novae.
Figure 1.3: Schematic of a polar. Material leaving the Roche lobe filling secondary via the inner Lagrangian point follows a ballistic trajectory until it encounters the magnetic field of the white dwarf. In polars this is sufficient to channel the accretion onto the poles of the white dwarf. The white dwarf is in synchronism with the orbital period.

Figure 1.4: Schematic of an intermediate polar (IP). In this case the magnetic field is not strong enough to prevent the formation of a disc. The inner portions of the disc are disrupted by the magnetic field and accretion is channelled onto the white dwarf via the field lines. In this case the white dwarf is spinning faster than the orbital period.
1.1.5 Accretion onto magnetic white dwarfs

In this section I shall first describe the 'standard model' of 10 years ago, and then detail the more recent departures from this model.

The standard model

As already mentioned the magnetic field of the white dwarf disrupts the accretion flow, either preventing the formation of a disc or removing the inner portion. The disruption occurs where the ram pressure of the accreting gas, \( P_{\text{ram}} = \rho v^2 \) (where \( \rho \) is the density of the gas and \( v \) is the velocity), balances the pressure exerted by the magnetic field, \( P_{\text{mag}} = B^2/(8\pi) \). For spherical accretion this is known as the Alfvén radius, \( r_A \)

\[
    r_A = 5.1 \times 10^8 \frac{M_7^{-2/7} M_4^{-1/7}}{\mu_{30}^{4/7}} \text{ cm}
\]  

(1.18)

where \( \mu_{30} \) is the magnetic moment of the white dwarf in units of \( 10^{30} \text{ G cm}^3 \). In polars the stream must distort the lines resulting in a smaller radius, \( r_{\text{mag}} < 0.4r_A \) (Hameury, King & Lasota 1986). In the disc disruption case the magnetospheric radius should be \( R_{\text{mag}} \approx 0.5r_A \). The stream may follow its ballistic journey still further, due to screening effects of the surface currents on the material, thus excluding the magnetic field (Lamb 1989).

The schematic view of the accretion region onto the polar regions of a white dwarf is shown in fig 1.5. Infalling cold supersonic gas forms a stand-off shock at some height above the white dwarf surface, with a characteristic temperature, \( T_s \), given by

\[
    T_s = 3.7 \times 10^8 M_1 R_p^{-1} \text{ K}
\]  

(1.19)

The height of the shock is found by considering that the cooling post-shock flow must have time to cool before reaching the white dwarf surface, for Bremsstrahlung dominated cooling this is
Figure 1.5: Accretion column geometry at the surface of a magnetised white dwarf (Frank, King & Raine 1992). Cyclotron emission comes from the heated post shock region.

\[ D_{\text{red}} \approx 9 \times 10^8 M_{12}^{-1} f_{-2} M_{1}^{3/2} R_{\odot}^{1/2} \text{ cm} \]  

(1.20)

where \( f \) is the fractional surface of the white dwarf and \( f_{-2} \) is \( 10^2 f \). A more rigorous treatment of post-shock conditions finds that the actual shock height, \( D \), is approximately one third of \( D_{\text{red}} \) (Frank, King & Raine 1992).

The emission from this accretion region can be divided into three main components (Lamb & Masters 1979, Imamura and Durisen 1983) namely;

- Optically thin bremsstrahlung emission. This emission originates from the shock heated accretion material. This material is cooling from \( T_s \) and hence bremsstrahlung temperatures are of \( \sim \text{ few } \times 10 \text{ keV} \).
- Optical and near Infra-red cyclotron emission. As electrons spiral down the magnetic field lines cyclotron emission is produced (see emission processes section), this
emission has certain characteristics;

- the intensity is beamed perpendicular to the magnetic field
- radiation (in the optically thin regime) is strongly polarised. When viewing perpendicular to the magnetic field lines the polarisation is linear, and circular polarisation is observed when viewing along the field.

- Soft X-ray and EUV emission blackbody emission from the white dwarf surface of temperature of a few \( \times 10^8 \) K. This emission results from reprocessed hard X-ray and cyclotron emission by the white dwarf.

For the shock height predicted by the standard model the soft X-ray emission is not expected to exceed the sum of the other two components

\[
L_{\text{soft}} \sim L_{\text{hard}} + L_{\text{cyc}}
\]  

The soft X-ray emission is likely not to be just blackbody emission, but this emission modified by atmospheric absorption. Models of white dwarf atmospheres by Williams, King and Brooker (1987) in the EUV and soft X-ray band have shown that blackbody models overestimate temperatures by factors of 2-5, but the luminosity is accurate to a factor of 2.

1.1.6 Problems with the standard models

The soft X-ray excess

X-ray observations of polars have found that the soft X-ray luminosity exceeds the hard X-ray and cyclotron flux by factors of 10 or more (King and Watson 1987, Ramsay et al. 1994). Three methods of producing this excess have been put forward:

- 'Blobby' accretion. Kuipers & Pringle (1982) first proposed that inhomogeneities may exist within the accretion flow. Denser filaments (or 'blobs') would shock within the surface of the white dwarf, here the hard X-ray radiation is quickly thermalised
thus producing just soft X-ray emission. Less dense material would shock in the same way as the standard model above the white dwarf surface producing the predicted amount of hard and soft X-ray emission, see fig 1.6 (Frank, King & Lasota 1988). The combination of these results in an enhancement in the soft X-ray flux.

- The ram pressure of the accreting material may depress the white dwarf surface by a few scale heights (to a depth approaching that of the shock height) at the base of the accretion column (Beuermann 1989, Stockman 1989). At present only qualitative predictions have been made. The soft X-ray enhancement results from hard X-rays being intercepted and thermalised by the walls of the trench.

- For low accretion rates ($\dot{M} \leq 10^{-1} \, \text{g cm}^{-2} \, \text{s}^{-1}$) Kuipers & Pringle (1982) predicted that ions in the accretion stream would lose their kinetic energy via Coulomb collisions with atmospheric electrons. The accretion energy from this can be radiated away by cyclotron radiation quickly and thus no shock forms. Lamb & Masters (1979) first highlighted this effect as the non-hydrodynamic regime and Thompson & Cawthorne (1987) investigated this and termed it the 'bombardment' solution. Woelk & Beuermann (1992, 1993) presented a full frequency dependent LTE solution of this bombardment as a function of white dwarf mass $M$, magnetic field strength $B$ and the specific accretion rate $\dot{\nu}$. They found that for low values of $\dot{\nu}$ the bremsstrahlung component is suppressed.

For IPs there is a lack of hard X-rays in the emitted spectra (Patterson 1994). It is likely that this problem is analogous to the soft X-ray problem in the polars, and may well be a similar effect producing the under luminosity. The reprocessed flux would be expected in the soft X-ray and EUV, but because many IPs are at distances of $\sim 200$ pc away they tend to suffer from interstellar absorption. An observation of RE0751+144, which is in a location of low interstellar absorption, found a strong soft X-ray/EUV component and it is reasonable to assume that this is the missing flux (Mason et al. 1992).
Soft X-rays

Photosphere Diffuse blob shocks above the photosphere, producing hard X-rays

Dense blobs shock below the photosphere, producing soft X-rays

Figure 1.6: Inhomogeneous (or 'blobby') accretion. Blobs of varying density and length hit the surface of the white dwarf. Sufficiently dense blobs penetrate the photosphere and radiate almost all their energy as soft X-rays (Frank, King & Raine 1992).

Where might inhomogeneities be produced?

There are two areas which might be expected to produce inhomogeneities within the accretion flow:

- The L₁ point. As material leaves the secondary via the L₁ point perturbations on the dynamical timescale in the Roche lobe potential might be expected (King 1989), thus causing disruption of the accretion flow. These would exist for all CVs, but for non-magnetic systems and IPs containing a disc these would not be observed due to smearing effects of the disc.

- In the region where material threads onto the magnetic field lines inhomogeneities might be expected to form. These may be caused by interchange (eg Rayleigh-Taylor) instabilities, which dominate mass flow through the magnetosphere (Arons & Lea 1976a,b; Arons & Lea 1980; Elsner & Lamb 1977). Alternatively complex magneto hydrodynamic (MHD) instabilities can cause blobs to be formed, eg see Aly & Kuijpers (1990).
Structured emission regions

Observations of polars (e.g. DP Leo (Schmidt 1989), VV Pup & VS34 Cen (Wickramasinghe 1989, Cropper 1989, Ferrario & Wickramasinghe 1990)) have shown that the emission regions must be extended. Optical polarisation studies show that the emission is best fit when modelled as coming from an arc-like region. Figure 1.7 shows the way accretion regions may be structured in the case of 'blobby accretion', it shows that the fractional area can be divided into three different components (Frank, King & Raine 1992), namely:

\[ f_{\text{acc}} \sim 10^{-5} \quad \text{the bombarded area, where hard X-rays would be produced if the blob was sufficiently diffuse} \]

\[ f_{\text{acc}} \sim 10^{-4} - 10^{-5} \quad \text{the radiating area, region over which soft X-rays would be produced} \]

\[ f_{\text{facc}} \sim 10^{-3} - 10^{-4} \quad \text{the region over which blobs usually land} \]

where

\[ f_{\text{acc}} \ll f_{\text{eff}} < 1 \ll f_{\text{facc}} < 1 \]  \hspace{1cm} (1.22)

see fig 1.7.

For IPs determining the size of the emission region in itself is a difficult task. Two indirect clues are; 1) the X-Ray and optical pulse shapes are approximately sinusoidal, which suggests a fairly large structure (King & Shaviv 1984). 2) The lack of any observable soft X-ray component could also signify a large accretion area. An observation of eclipses in H0253-193 from ingress/egress timings indicated a projected diameter of \( \sim 1.5 \times 10^9 \) cm (Kamata et al. 1991), however this assumes that the secondary’s edge is sharp.

1.1.7 The condition for synchronism

Magnetospheric interaction between the magnetic field of the white dwarf and the intrinsic field of the secondary produces a synchronism torque. If this is sufficient to balance the
Figure 1.7: Schematic picture of three fractional areas $f_{\text{acc}}$, $f_{\text{eff}}$ and $f_{\text{zone}}$. The fractional areas over which the blobs accrete $f_{\text{acc}}$ are always much smaller than that over which soft X-rays emerge $f_{\text{eff}}$. The geometrical region over which blobs land is characterized by $f_{\text{zone}}$. In normal accretion states $f_{\text{eff}} \leq f_{\text{zone}}$ while in anomalous states $f_{\text{eff}} \ll f_{\text{zone}}$ (Frank, King & Raine 1992).
accretion torque than the system becomes synchronised. A rough estimate of the required
\( \mu \) can be found by setting \( R_{\text{mag}} = R_{\text{orb}} \) in 1.18 and applying Keplers third law,

\[
\mu_{32} = 110 \, M_2^{4/7} \left( 1 + q \right)^{7/12} M_1^{1/2} P_2^{7/6}
\]

(1.23)

where \( P_2 = P_{\text{orb}}/2 \) hr. Typically for polars (\( M_1 = 0.7 \), \( q = 0.25 \), \( M_17 = 0.1 \) and \( P_2 = 1 \))
this implies \( \mu_{32} > 30 \). It can be seen from 1.23 that for longer period systems synchro­
nism becomes harder. Recently a polar has been discovered with an 8hr orbital period
(RXJ051541+0104.6, Walter, Wolk & Adams 1994), this would require a reasonably large
magnetic field for synchronism to occur. The asynchronous IPs tend to have longer periods
than polars (see section 1.1.9 for a discussion of this).

1.1.8 Polars

The white dwarf in polars is synchronised with the orbital period. Some systems have
been found to have polarized emission at a slightly lower period than the orbital period
(e.g. V1500 Cyg: Schmidt, Stockman & Lamb 1988; BY Cam: Silber et al.1992) and
RXJ1940-1025 is suspected to be asynchronous in the other sense (Watson et al.1995).
Whilst this technically makes these systems IPs, they resemble polars so closely that they
are regarded as polars that have ’gone astray’ (possibly because of lack of phase lock in a
classical nova eruption), rather than another member of the IP group.

Currently there have been 31 polars identified (Ritter & Kolb 1993), ranging from periods
of ~ 8 – 1.32 hours and magnetic fields of 10-80 MG. The prototype system, AM Herculis
was identified as a CV containing a magnetic white dwarf in 1976 (Tapia 1976) from the
presence of strong circular and linear polarisation modulated at a period of 3.1 hr in optical
data. At the same time a variable X-ray source was found to be present within the AM
Her error box (Hearn, Richardson & Clark 1976).
X-ray variability

Variability in the soft X-ray region is produced as the emission region rotates and consequently changes its projected area into our line of sight. Eclipses by the secondary are also observed in some systems producing another form of modulation. Other dip like features are seen in the X-ray light curves of polars, these have been shown to be caused by eclipsing the polar region by the accretion stream itself (e.g. Watson et al. 1989). At periods shorter than $P_{\text{orb}}$ variability has been seen on two timescales, of a few tens of seconds to a few tens of minutes and of the order of a second, in the optical. X-ray observations have found these quasi-periodic variations (see Cropper 1990 and references therein) on the longer timescale for a numbers of systems (the faster signal has not been seen due to insufficient signal to noise of X-ray detectors).

Two pole accretion

Various observations of polars have observed changes in the light curve which can only be explained by a change from one to two pole accretion. Simultaneous hard and soft X-ray observations of AM Her (Heise et al. 1985) showed the two light curves to be modulated $\sim 165^\circ$ out of phase. The soft X-ray maximum occurred at linear polarisation phase of 0.2, which was different to previous observations where the soft and hard X-ray maximums were coincident at phase 0.62 (Hearn & Richardson 1977, Tuohy et al. 1981). In this case the hard X-ray emission was from the usual upper pole, but the soft X-rays were being emitted from a second pole. QQ Vul also showed a change in the soft X-ray light curve (Osborne et al. 1987, Beardmore et al. 1995) which again is interpreted as emission from two poles.

Polarisation studies of some systems have shown structured linear polarisation pulses which could not be explained by a single emission region. Modelling of such data in AM Her found there to be two extended emission regions (Wickramasinghe 1991).
1.1.9 Intermediate polars (IPs)

IPs have smaller magnetic fields and thus do not achieve synchronism. Magnetic fields have been detected for two IPs:

- From polarisation measurements BG CMi was found to have a magnetic field of \( \sim 4 \times 10^8 \) G (Penning, Schmidt & Liebert 1986; West, Berriman & Schmidt 1987).
- RE0751-44 again from polarisation studies, has been found to have a magnetic field of 8-18 MG (Piirola et al. 1993).

Some of the flux from the accretion region is intercepted by the secondary star and the bright spot on the disc (if the system is a disc system). Heating of these regions occurs at a slightly lower frequency with the heated target being illuminated one less time per orbit. This produces emission at the 'sideband' frequency, given by:

\[
\omega_{\text{repro}} = \omega_{\text{spin}} - \Omega_{\text{orb}}
\]  

Other sideband periods are also expected, see later in this section for more details.

Hard X-ray pulsations

Many IPs show strong hard X-ray pulses as the white dwarf rotates. King & Shaviv (1984) calculated the light curves expected from a tall, optically thin accretion column rotating in and out of view. These light curves were quasi-sinusoidal, with the detailed shape dependent on binary inclination and magnetic obliquity. The geometry alone dictates the visibility of the emission region, so there is no energy dependence. However Rosen, Mason & Cordova (1988) reported considerable energy dependence in the X-ray pulses of EX Hya, and developed an alternative 'accretion curtain' model from radial velocity variations, in which absorption effects in the column produce the waveform.

Examination of many systems pulse shapes at many energies (Norton & Watson 1989)
shows that photoelectric absorption is quite important, but so is geometry (since some stars show strong pulses even at very high energy). In general the pulse shapes need to be studied separately for each system.

Notches in the hard X-ray light curve of FO Aqr have been observed occurring at a constant spin phase (Norton et al. 1992) and have been shown to be energy independent. Thus this notch component is probably produced by the occultation of a small 'hot spot' on the surface of the white dwarf.

Orbital modulation and disc or no disc accretion

The majority of X-ray emitting IPs show some orbital modulation (Hellier, Garlick & Mason 1993), most showing a harder X-ray spectrum during the minimum of the orbital light curve. It is likely that the modulation is due to photoelectric absorption in a structure which orbits the white dwarf at the orbital period.

Dips occur ~ 0.2 cycles before inferior conjunction of the secondary; this is the phase where the bright spot at the edge of the disc is closest to the line of sight, hence the bright spot is a prime suspect for causing the modulation. However to explain the obscuration of non-eclipsing systems the gas must be at least 30° above the orbital plane. Also the minima are fairly broad, some even quasi-sinusoidal; a structure at the outer edge of the disc would have to be very large and extended azimuthally to produce such modulation.

It has been suggested that IPs may accrete directly from the secondary without the mediation of the disc (Hameury, King & Lasota 1986) or disk overflow of the stream could occur (Hellier 1995). The stream of material from the secondary to the magnetosphere may well be the cause of modulation as it could be locked to the binary period and rise well above the orbital plane, and could be extended in azimuth since it is likely to be sheared by the magnetosphere (Norton et al. 1992). This scenario would not be expected to produce stable X-ray dips, and indeed the observations indicate that the dips do tend to be quite unstable in depth and phase.

As already mentioned some form of non-disc accretion is occurring in IPs (Hameury, King
Lasota 1986, King & Lasota 1991). Many observations do show the presence of discs (Hellier 1991). Eclipses have been used to give the spatial extent of the eclipsed material and the 'S-wave' shape of the emission line profiles give the location of the accreting material; both types of observations imply that a disc is present. Hellier (1991) also argues in the case of FO Aqr the X-ray spin pulse with its sinusoidal profile requires an azimuthally uniform accretion flow, which is naturally provided in the case of disc accretion. Indeed in FO Aqr there is evidence that both accretion modes can occur simultaneously when the mass from the secondary overshoots the accretion disc and goes onto the magnetosphere directly (Hellier, Mason & Cropper 1990).

Wynn & King (1992) computed power spectra for IPs using various accretion geometries. They showed that it was difficult to rule out discless accretion on the basis of their power spectra. In particular, the lack of any strong beat-frequency ($\omega - \Omega$) modulation or dominant spin frequency modulation does not necessarily indicate that disc accretion is dominant. Within hard X-ray power spectra there were frequencies present which indicated the dominance of discless accretion mechanisms, notably the $2\omega - \Omega$ (orbital) component. Physical symmetries and absorption lead to difficulty in separating disc and discless accretion especially at lower energies, however harder X-ray power spectra ($\geq 4\text{keV}$) with no components but the spin and its harmonics would indicate that that system was accreting predominately via a disc. Therefore making power spectra, specifically at harder X-ray energies a useful diagnostic tool for determining the presence or not of a disc.

1.1.10 Evolution

Initially most CVs form from a main sequence binary system with a large separation ($P_{\text{orb}} \sim 1\text{ yr}$). The most massive star being the primary (white dwarf) 'parent'. The primary evolves quickly and during red giant expansion the primary's Roche lobe is filled and unstable mass transfer starts. The secondary is overwhelmed by this mass transfer and is thrown out of thermal equilibrium. This results in a 'Common Envelope (CE)' around the two stars (see Eggleton 1993b, Loore & Doom 1992, Ritter 1976 and Taam & Bodenheimer 1989).
During this CE phase, which is short-lived (~$10^3 - 10^4$ yrs), most of the systems angular momentum is lost due to mass loss from the system. This angular momentum loss is thought to be caused by deposition of energy into the envelope by frictional drag exerted on the secondary during its motion within the CE. This is sufficient to drive mass out of the binary, preferentially into the binary plane (Livio & Soker 1988, Taam & Bodenheimer 1989, Regos & Tout 1995). This angular momentum loss leads to a reduction in the binary separation and period. This phase concludes with the secondary rejoining the main sequence and the core of the primary being a white dwarf. Further angular momentum loss (either by gravitational radiation or magnetic braking) allows the secondary to come into contact with its Roche lobe and hence mass transfer restarts and a CV is born.

CVs follow a secular evolution path from longer to shorter periods via angular momentum loss (for a review of the evolution see Rappaport, Joss & Webbink 1982). For the system to remain a CV with stable mass transfer the secondary must remain in contact with its Roche lobe despite mass loss. There are two ways in which this could occur (King 1988);

1. The star can expand to fill its Roche lobe due to nuclear evolution of its core. For this to occur on a reasonable timescale $M_2 \geq 1M_\odot$ at the onset of mass transfer. GK Per is an example of such an evolved secondary system with a $P_{\text{orb}} \approx 48$ hr. However in general most CV secondaries are not this massive.

2. The Roche lobe shrinks onto the secondary. This requires there to be angular momentum loss in the system.

To produce this further angular momentum loss there are two mechanisms (King 1988), namely;

- Gravitational Radiation (Kraft, Matthews & Greenstein 1962). For systems with periods of $P_{\text{orb}} < 3$ hr predicted mass transfer rates are $\dot{M} < 2 \times 10^{-10} M_\odot$ yr$^{-1}$, which are consistent with observation (Patterson 1984). For periods of $P_{\text{orb}} > 3$ hr mass transfer rates are too low.

- Magnetic Braking. For systems with periods of $P_{\text{orb}} > 3$ hr magnetic braking due to the stellar wind from the secondary is hypothesized. Using the magnetic braking
law of Mestel & Spruit (1987) Hameury et al. (1988) found mass transfer rates in agreement with those observed for longer period systems.

The period distribution

Figure 1.8 shows the period distribution for CVs. This distribution shows three distinct features:

1. A short period cut off at $P_{\text{orb}} \sim 80$ min.
2. A long period cut off at $P_{\text{orb}} \sim 13$ hr.
3. A 'period gap' between 2-3 hrs.
The short period cut off is related to the extinction of Hydrogen burning in the core of the secondary. As degeneracy sets in the star is pushed out of thermal equilibrium and expands upon further mass loss. The binary expands and the orbital period increases. Rappaport, Joes & Webbink (1982) calculated that for a secondary with solar composition $P_{\text{min}} \sim 70$ min.

For stable mass transfer the secondary mass must be less than that of the primary. As the primary mass can not exceed the Chandrasekhar limit, the secondary mass also cannot exceed this limit of $\leq 1.4M_\odot$. The upper limit of $\sim 13$ hr for most CV periods (GK Per is an exception to this, as it has a more evolved secondary) is thus derived simply from the orbital period - mass relationship for CV secondaries (see prev).

The period gap is explained as the sudden reduction in the effectiveness of magnetic braking. At $P_{\text{orb}} \sim 3$ hr the mass of the secondary goes below $\sim 0.3 M_\odot$ and hence becomes fully convective. This is thought to either reduce the magnetic field of the secondary or reduce the stellar wind, either of which would reduce the effectiveness of magnetic braking (King 1988). The secondary then contracts below the Roche lobe and accretion ceases. Angular momentum loss still continues via gravitational radiation until eventually (at $P_{\text{orb}} \sim 2$ hr) the secondary regains contact with its Roche lobe and accretion restarts.

Recent ROSAT observations have found several new systems within the period gap (Buckley 1993, Schwope, Thomas & Beuermann 1993), leading to concerns that the period gap is much reduced or not present for polars. The evolutionary path described above cannot account for this, which leads to suggestions that in polars magnetic braking is not as efficient a mechanism for removing angular momentum as first thought (Wickramasinghe & Wu 1994; Li, Wu & Wickramasinghe 1994). However a comparison between the orbital period distributions of non-magnetic systems and polars below $P_{\text{orb}} = 5$ hr by Wheatley (1995) shows no significant differences between the two distributions. The differences in the orbital period distributions occurs above $P_{\text{orb}} = 5$ hr, where there are a lack of polars. This is well understood and is consequence of the requirement for synchronism. Thus the period gap problem in polars appears to be solved. One polar which appears in the gap, QS Tel with a $P_{\text{orb}} = 2.33$ hr, cannot explain its presence in the gap using secular evolution without using contrived arguments (Hameury, King & Lasota 1988). The most
likely reason for its presence in the gap is that it was born there (see e.g. Hameury et al. 1991).

Another feature of the period distribution is that most IPs lie above the period gap and polars are below it. Suggestions were made that this indicated that IPs evolve into polars (Chanmugam & Ray 1984, King, Frank & Ritter 1985). This assumes that the magnetic field strength of the IPs is basically similar to that of polars, however most IPs are unpolarised. It has been found to be difficult to hide polarised flux in, say, the accretion disc (Wickramasinghe, Wu & Ferrario 1991). The current view is that IPs have lower field strengths (Wickramasinghe, Wu & Ferrario 1991; Lamb & Melia 1988; King & Lasota 1991). This leads to two problems;

1. Where are the polar precursors?
2. If polars are not evolved IPs, where are the evolved IPs?

Patterson (1994) postulates some solutions to these problems. As the IP lifetime is probably shorter than polars then not many IPs are needed to make up the polar population. RE0751+144, with a B field of 8-18 MG is a good example of a polar ancestor. Finding a population of low field short P_{orb} stars which most IPs evolve into is harder. Patterson (1994) believes that these stars would be faint (as a low P_{orb} implies a low M) and as the disc is not normal it may not undergo dwarf novae outbursts, which is the usual way of detecting faint systems.

1.2 Emission and absorption processes

Here I shall only describe processes which are important for X-ray spectra in magnetic CVs.

1.2.1 Absorption

Absorption of radiation is characterised by
\[ I = I_0 e^{-\tau(E)} \]  
\[ \tau(E) = \sigma(E) N \]

where \( \tau \) is the optical depth, which is given by;

\[ \tau(E) = \sigma(E) N \]

where \( \sigma(E) \) is the absorption coefficient and \( N \) is the column density of the absorbing material. The absorption coefficient for a material is the sum of the absorption coefficients for all the mechanisms occurring and the ranges of species and relative densities present.

For this introduction I shall just deal with one absorption mechanism (as it is the most widely used during this work) any others that are used will be dealt with at that point.

**Photoelectric absorption**

Photoelectric (or bound-free) absorption occurs when incident radiation provides enough energy to an atomic electron for it to released from the atom. The cross section is highly complicated, for the purpose of the analysis in this thesis I used the effective photoelectric cross section (cross section per unit H) by Morrison and M‘Cammon (1983). They calculated the cross section of each species in a neutral cosmic abundance gas and then weighted these by the fractional abundance relative to H, thus determining a column density \( N_H \) for the medium.

1.2.2 Emission processes

**Blackbody emission**

Blackbody radiation is emitted by optically thick material \((\tau \gg 1)\) which is in thermal equilibrium. The emitted spectrum is characterised by the temperature of the material. The brightness of the blackbody is given by the Planck function;
Bremsstrahlung emission

As an electron interacts with Coulomb fields of ions, the particle is caused to accelerate. This gives rise to continuum emission known as Bremsstrahlung. This is the dominant cooling mechanism for optically thin plasma of temperatures of > 10^7 K. For a Maxwellian distribution of electron velocities, the spectral emission per unit volume is:

\[
\frac{dP_B(T)}{dV d\nu} = 6.8 \times 10^{28} T^{-1/2} e^{-E/kT} N_e N_i Z^2 g_B(T, E) \text{ erg cm}^{-3} \text{ s}^{-1} \text{ Hz}^{-1}
\]  

where \(N_e\) is the electron density, \(N_i\) is the ion density and \(g_B\) is the gaunt factor (Zombeck 1990).

The total bremsstrahlung emission is

\[
\frac{dP_B}{dV} = 2.4 \times 10^{-27} T^{1/2} N_e^2 \text{ erg cm}^{-3} \text{ s}^{-1}
\]  

(assuming \(g_B(T) \approx 1.2\))(Zombeck 1990).

Line emission

Raymond and Smith (1977) calculated the X-ray emission spectrum from a collisionally ionised plasma. The emitted spectrum is a strong function of the plasma temperature. At 10^6 K many lines are seen at soft X-ray energies, e.g C, N, O, Ne, Mg, Si, S and Fe. As the temperature increases (10^7 K) the line emission is dominated by emission from iron L between 0.7 and 1.0 keV. Iron K emission between 6.7-6.9 keV dominates at temperatures of 10^8 K. Their model includes using continua spectra including the effects of these emission lines.
For an atom which has had the K shell electron ejected (by the photoelectric effect) there are two ways in which this atom can de-excite:

1. Fluorescence: emission of an X-ray photon when the K shell is filled by an electron from another sub-shell, e.g. for argon $\text{Ar}(K^{-1}) \rightarrow \text{Ar}(L_{1}^{-1}) + h\nu$.

2. Auger decay: the K shell is filled by an L shell electron with the simultaneous emission of another L shell electron, e.g. $\text{Ar}(K^{-1}) \rightarrow \text{Ar}(L_{2}^{-2}) + e^{-}$.

The fluorescence yield $\omega_{k}$ (the probability that radiative transfer occurs post absorption) increases with increasing Z. As iron has the highest Z of species commonly found in astrophysics then one would expect this to be the most easily observed fluorescence line.

**Cyclotron emission**

Cyclotron emission is emitted by electrons rotating around magnetic field lines. The fundamental cyclotron frequency is given by:

$$\nu_{c} = \frac{eB}{2\pi c m_{e}} = 2.8 \times 10^{13} B_{\gamma} \text{ Hz} \quad (1.31)$$

where $B_{\gamma}$ is the magnetic field strength in units of $10^{7}$ G.

In the case of polars cyclotron emission is optically thick at the fundamental frequency and optically thin beyond some last harmonic (Chanmugam 1980, Chanmugam & Dulk 1981). Thus (for polars) as the electrons are heated to shock temperatures significant power occurs at high harmonics $m$ of $\nu_{c}$. 

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Chapter 2

The Soft X-ray Properties of EK UMa.

2.1 Introduction

EK UMa was discovered as the serendipitous Einstein X-ray source 1E 1048.5 + 5241 by Morris et al. (1987). Subsequent optical observations (namely spectroscopy and polarimetry) showed EK UMa to be a polar (see chapter 1) of faint optical magnitude ($m_v = 18.5 > 19.5$). The orbital period was determined to be $114.5 \pm 0.2$ min and the distance was estimated to be in the range 0.5 to 2 kpc. Morris et al. (1987), by examining the features of the optical light curve, also put some crude limits on the geometry of the system, specifically that the inclination and the magnetic colatitude are both $56^\circ \pm 19^\circ$. The short X-ray observation (exposure time of $\sim 1800$ s) only showed that the EK UMa was both soft and variable (see fig 2.1). Cyclotron humps in the optical were observed and modelled by Cropper, Mason & Mukai (1990), resulting in a magnetic field measurement of $47 \pm 2$ MG.

As the Einstein observation was short, a 14 ks ROSAT observation was performed to better characterize the X-ray features of EK UMa. This chapter describes this observation and contemporaneous optical and infra-red observations.
Figure 2.1: *Einstein* IPC X-ray light curve of EK UMa, covering ≈ 28% of the binary orbit, taken from Morris *et al.* (1987).
2.2 X-Ray Observation and Data Analysis

ROSAT (see appendix A) made a pointed observation of EK UMa on 12th May 1992. This observation consisted of six continuous time intervals, with a total exposure time of \( \approx 14 \text{ ksec} \). The observation log can be seen in table 2.1. EK UMa was observed with the PSPC (0.1-2.5 keV) 40 arcmin off-axis, thus limiting the effects of the detector structure induced by the satellite wobble (a more detailed explanation can be found in appendix A). The data was extracted within an accumulation radius of 4.8 arcmin using the Starlink package ASTERIX (Saxton 1992). The background was accumulated from a surrounding annulus. The average PSPC count rate, corrected for vignetting and obscuration from the wires was 1.4 cts s\(^{-1}\).

The EUV Wide Field Camera (WFC) (Pye, Watson & Pounds 1991) did not detect EK UMa during its quasi-simultaneous 15 000 s exposure. Upper limits (3\(\sigma\)) on the count rate for the sla (90-210 eV) and s2b (61-111 eV) filters were determined to be \(4.5 \times 10^{-3}\) and \(4.2 \times 10^{-3}\) cts s\(^{-1}\) respectively.

2.2.1 Timing Analysis

A light curve of the PSPC data on EK UMa covering all six observation sections was created, background subtracted and corrected for energy-dependent vignetting (using an energy of 0.2 keV), dead-time and PSF scattering. This light curve was then folded on the 114.5 minute orbital period, and can be seen in fig 2.2. Phasing of this plot is purely arbitrary as the ephemeris of Morris et al. (1987) is no longer valid. This folded light curve clearly shows the presence of a bright and faint phase, with the bright phase covering 0.63 of the orbital cycle. EK UMa was not detected during the faint phase to an upper limit of (3\(\sigma\)) of 0.01 cts s\(^{-1}\).

In addition to the bright and faint phases seen in fig 2.2, there is also a succession of dips in the light curve over the phase range 0.9-0.15. One of these dips, centered on phase 0.05, is broad (\(\Delta \phi = 0.04\)) and has zero flux during the dip minimum. Fig 2.3 shows
<table>
<thead>
<tr>
<th>Section</th>
<th>Start Date</th>
<th>Start Time (UT)</th>
<th>Duration (mins)</th>
<th>Instrument/Filter</th>
</tr>
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<td>XRT PSPC</td>
</tr>
<tr>
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<td>13 May 92</td>
<td>13 13</td>
<td>31.7</td>
<td>XRT PSPC</td>
</tr>
<tr>
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<td>16 28</td>
<td>30.2</td>
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</tr>
<tr>
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<td>21 23</td>
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</tr>
<tr>
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<td>23 01</td>
<td>37.4</td>
<td>XRT PSPC</td>
</tr>
<tr>
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<td>XRT PSPC</td>
</tr>
<tr>
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<td>21 29</td>
<td>40.5</td>
<td>WFC S1A</td>
<td></td>
</tr>
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</tr>
<tr>
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<td>21 27</td>
<td>52.0</td>
<td>WFC S2B</td>
<td></td>
</tr>
<tr>
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<td>23 01</td>
<td>39.9</td>
<td>WFC S2B</td>
<td></td>
</tr>
<tr>
<td>14 May 92</td>
<td>01 59</td>
<td>52.8</td>
<td>WFC S2B</td>
<td></td>
</tr>
<tr>
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<td>5</td>
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<td>10</td>
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<td>26 Feb 93</td>
<td>13 26</td>
<td>0.3</td>
<td>UKIRT/K</td>
<td></td>
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<tr>
<td>26 Feb 93</td>
<td>13 28</td>
<td>0.3</td>
<td>UKIRT/K</td>
<td></td>
</tr>
<tr>
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<td>13 29</td>
<td>0.3</td>
<td>UKIRT/K</td>
<td></td>
</tr>
<tr>
<td>26 Feb 93</td>
<td>13 30</td>
<td>0.3</td>
<td>UKIRT/K</td>
<td></td>
</tr>
<tr>
<td>26 Feb 93</td>
<td>13 31</td>
<td>0.3</td>
<td>UKIRT/K</td>
<td></td>
</tr>
</tbody>
</table>
Figure 2.2: The soft X-ray (0.1-2 keV) light curve of EK UMa, corrected for vignetting, shown folded on the 114 min orbital period and plotted twice in 0.5 min bins. The phase is arbitrary, and an epoch of HJD 2448755.388 was used.
the six separate observation slots and clearly shows that the broad dip was observed on
two different occasions. The light curve also shows other dips and flares, including a dip
at phase 0.1 which is seen on three occasions. The two prominent dips seen before phase
zero, each of width $\Delta \phi = 0.03$, both occurred in the same observation section and this
phase interval was not observed again, so it is impossible to say whether these dips are
permanent features of the light curve.

The flux transitions occur on short time-scales, of the order of $\approx 10 \text{ s}$. A Fourier analysis
was performed to search for minute-to-second quasi-periodic modulation. This was done
by using the raw (no background subtraction) light curve with a binning of 0.5 sec. Fourier
transforms (using the Leahy et al. 1983 normalisation, resulting in a mean noise power of
2) of the data sections were created (each section had to have the same length, 1000s).
The Fourier transforms were then rebinned by a factor of 4, resulting in a power spectrum
with $\sim 200$ bins. The five separate transform were then summed together to improve
signal to noise of the power spectrum. Fig 2.4 shows the result of this Fourier analysis.
The detection power in one bin (for a specific confidence, Van der Klis 1988) can be set
by using a scaled $\chi^2$ distribution and then divided by the number of trails (the number of
trails is just the product of the factor of rebinning and the number of transforms summed
together). No modulation on these time-scales was detected. The upper limit to the signal
power was thus set, $P_{ul}$,

$$P_{ul} = P_{max} - P_{exceed}$$

(2.1)

where $P_{max}$ is the maximum power in the frequency range of interest, and $P_{exceed}$ is the
power level usually exceeded by the noise (typically $\sim 2$). The rms variation of this upper
limit then can be determined by

$$\text{rms} = \sqrt{\frac{W P_{ul}}{N_{phot}}}$$

(2.2)

where $W$ is the rebin factor, and $N_{phot}$ is the number of photons, thus giving an upper
limit (90% confidence) on the rms variation of 6% in the frequency range $0.01$-$0.1 \text{ Hz}$. 

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Figure 2.3: The individual sections of the observation with 15 s binning. Panels (a) to (f) show the time sections 1 to 6, respectively (see Table 2.1). The orbital cycle numbers of the sections are 1, 9, 11, 13, 14 and 15 for (a)-(f) respectively.
Figure 2.4: The average FFT of EK UMa. The solid line indicates the mean noise power of 2 (set by the Leahy et al. (1983) normalisation). For this transform a rebinning of 4 was used and 5 transforms were summed together, thus resulting in a $P_{\text{detect}}$ of 4.3.
2.2.2 Spectral Analysis

The data were searched for spectral variation by the use of a hardness ratio. This ratio was defined as

\[ HR = \frac{\text{Hard} - \text{Soft}}{\text{Hard} + \text{Soft}} \]

with the PHA channel ranges for the hard and soft bands chosen such that the ratio for the averaged bright phase would be zero, i.e. the soft and hard bands were PHA channels 11-20 (0.1-0.2 keV) and 21-200 (0.2-2 keV) respectively. The hardness ratio showed no significant change in spite of large changes in count rate during the bright phase, as is illustrated for one section of data in fig 2.6. In an effort to explain this lack of change the hardness ratio was predicted. Assuming that any reduction in flux from maximum is due to increased cold absorption and adopting the best-fitting spectral model parameters (see later in this section) the resulting ratio as a function of \( N_H \) can be seen in fig 2.5.

This predicted ratio can also be seen overlaid on the actual ratio in fig 2.6. This and...
similar plots show that the PSPC is insensitive to hardness ratio changes caused by column increases of $8 \times 10^{18} < \Delta N_H < 3 \times 10^{20}$ atoms cm$^{-2}$, thus $N_H$ changes of this level can not be ruled out by a hardness ratio showing no variation. Given the lack of observed spectral variation all the data were summed together for the subsequent analysis.

Spectral fits were performed using the XSPEC package (version number 8.3, Shafer et al. 1991). The PHA data were binned into 26 bins of at least 10 counts in each, and each PHA channel was binned by at least a factor of 2. The response matrix used for this reduction was the version released in January 1993 (DRM 36). A systematic error of 1% was included to account for uncertainties in the response matrix. Initially the spectrum was modelled with an absorbed blackbody and details of this fit are in Table 2.2.

A good fit was achieved, with a reduced $\chi^2$ of 1.35, but on examining the $\chi^2$ residuals there appeared to be some evidence that there may also be a hard X-ray component at energies greater than 1 keV. To test this hypothesis the fit was repeated with the inclusion of an additional Bremsstrahlung component with a fixed temperature of 10 keV. The resulting reduction in $\chi^2$ (from 31 to 28) was shown, by means of an F-test, not to be significant even at the 90% confidence level. Thus the 95% upper limit on the hard X-ray flux (1-2 keV) emanating from EK UMa is $7.3 \times 10^{-14}$ erg cm$^{-2}$ s$^{-1}$.

The confidence contours of blackbody temperature and luminosity (assuming a distance of 0.4 kpc; see later for the derivation of the distance), together with contours of constant $N_H$ for the single-component fit, can be seen in figure 2.7. the best-fitting value is marked by a cross. The blackbody temperature is constrained to be in the range $50 < kT < 62$ eV.

<table>
<thead>
<tr>
<th>Model Parameter / Statistic</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>$N_H$ (atoms cm$^{-2}$)</td>
<td>$7.8^{+19}_{-7.6} \times 10^{18}$</td>
</tr>
<tr>
<td>Blackbody $kT$ (eV)</td>
<td>57$^{+1}_{-6}$</td>
</tr>
<tr>
<td>$\chi^2 / \nu$</td>
<td>31 / 23</td>
</tr>
</tbody>
</table>

Errors calculated for a $\Delta \chi^2$ of 4.61.
Figure 2.6: Section 6 of the soft X-ray data covering a possible absorption dip, with the corresponding hardness ratio beneath. The predicted hardness ratio is plotted in the lower panel as a continuous line, assuming that any reduction in the count rate is due to an increase in absorption.
Figure 2.7: The 68%, 90% and 99% confidence regions of blackbody luminosity versus kT. Included also are contours of constant $N_H$. The luminosity is calculated as $L = \pi f d^2$, where $f$ is the peak X-ray flux and $d=400$ pc.
The unabsorbed total flux of the soft X-ray blackbody component was determined to be in the range \((4.5 - 7.2) \times 10^{-12} \text{ erg cm}^{-2} \text{s}^{-1}\). The upper limit to the total hard X-ray flux was found to be \(5.8 \times 10^{-13} \text{ erg cm}^{-2} \text{s}^{-1}\), for an assumed temperature of 10 keV.

2.3 Optical and Infrared Observations and Determination of a Distance Limit

V-band measurements were made by P. Hakala on June 1st 1992, 20 days after the PSPC observation (see table 2.1), at the 2.56-m Nordic Optical Telescope (NOT). The observations used a front-illuminated liquid-nitrogen cooled Tektronix CCD with 512 x 512 0.2-arcsec pixels, with the CCD assembly located at the f/11 Cassegrain focus. Two measurements were made, with integration times of 5 and 10 min, giving \(V = 20.0 \pm 0.1\) and \(20.0 \pm 0.1\) at phases 0.44 and 0.49, respectively, with respect to the ROSAT observation (see fig 2.2). These observations were thus obtained during the faint X-ray phase.

On February 26th 1993, five-spot K-band measurements were made by P. Wheatley at the United Kingdom Infrared Telescope (UKIRT). These were 20s integrations taken over a 5 min period using the UKT 9 instrument, which is an InSb photovoltaic detector system. The details of these measurements can be seen in table 2.1. The measurements were taken at phase 0.64-0.68, i.e. during the X-ray faint phase. These gave an average value of \(K = 16.8 \pm 0.7\).

The distance to EK UMa was then calculated using the method of Bailey (1981). This makes use of the fact that the surface brightness of a late-type star in the K band \((S_K)\) is practically independent of surface temperature and gravity, so its apparent magnitude is a function only of surface area and of distance \((d)\) (extinction is a very small effect at this wavelength). The distance is found by using

\[
S_K = K + 5 - 5 \log d + 5 \log \left(\frac{R}{R_\odot}\right)
\]

where

\[
S_K = 2.56 + 0.508(V - K) \quad \text{(for } V - K < 3.5)\]
Of course, in EK UMa accretion processes such as cyclotron radiation can contribute to the K-band flux, so only a lower limit can be obtained for the distance. Taking the lower limit of the K-band magnitude, and making use of standard binary-star relationships (see Chapter 1), we find a lower limit to the distance of 390pc. ¹ A limit has also been derived making use of a more recent tight relationship between absolute K magnitude and mass for M dwarfs described by Henry & McCarthy (1990)

\[
\log M = -0.188M_K + 0.688
\]

where \( M \) is the mass of the M dwarf, determined by using the relation between secondary mass and period (see 1.13 in chapter 1) in solar masses, and \( M_K \) is the absolute K magnitude. From the mass of the secondary and measured apparent K magnitude we find \( d > 434 \) pc, using the standard distance modulus equation (\( M_K - m_K = 5 - 5 \log d \)). In what follows I have chosen a distance to EK UMa of 400pc.

2.4 Discussion

The folded X-ray light curve of EK UMa shows bright and faint phases which are consistent with a single emitting pole rotating in and out of view. This type of light curve is seen in a number of other polars such as VV Pup, ST LMi and UZ For (Mason 1985; Osborne et al. 1988). The dip seen at phase \( \phi = 0.05 \) was detected twice during this observation, and is also seen in the ROSAT all-sky survey data (Besermann and Thomas 1994). The hardness ratio did not show a significant change over the dip interval. However, modelling of the effect of increasing cold absorption on the count rate showed we cannot rule out the possibility that the dip is caused by such an increase. The dip is similar to those seen in other polar light curves, such as those of EF Eri and QQ Vul (Osborne et al. 1986; Watson et al. 1989), where they have been attributed to absorption of the flux from the X-ray emitting region by the accretion stream as it rises out of the orbital plane constrained by the magnetic field of the white dwarf. It is reasonable to assume the same for EK UMa. Modelling has enabled us to place limits on the accretion-stream column density of \( 8 \times 10^{18} < \Delta N_H < 3 \times 10^{20} \) atoms cm\(^{-2}\).

¹Using a revised calibration for \( S_K \) by Ramseyer (1994) results in a lower limit to the distance of \( \sim 550 \) pc.
The light curve shows several other dips in the range $\Delta \phi = 0.9 - 0.15$. The relatively broad dip seen at phase $\phi = 0.05$ was observed twice, and the dips at phases $\phi = 0.1$ and $0.13$ were observed three times each (i.e., every time these phases were observed). The earlier dips, at phases $\phi = 0.91$ and $0.95$, were only observed once. The total range over which these dips are seen is $\Delta \phi \approx 0.25$, that is, a $90^\circ$ range in azimuth. Although EK UMa appears to be an extreme example, multiple dips have been seen in other polars, for example in UZ For (Osborne et al. 1988), QQ Vul, AN UMa and V834 Cen (Mason 1985). If all the dips in EK UMa are due to absorption, then this may suggest that multiple discrete accretion streams are being lifted out of the orbital plane by the magnetic field of the white dwarf. These multiple streams might be due to a complex magnetic field geometry of the white dwarf, or due to density inhomogeneities in the accretion flow. However, high order multipolar field components tend to dominate close to the white dwarf surface, thus placing the absorption source near to the white dwarf, contrary to opinion that the dips result from occultation by distant parts ($\approx 10^{10}$ cm) of the accretion stream (King & Williams 1985). For the 'blobby' accretion model 'blobs' of material with different densities would penetrate to different depths of the magnetosphere and thus thread to different regions of the magnetic field (King 1995). This at first seems a more promising method of explaining the dips, however if the stream were composed of a continuous range of densities then a broad stream onto the white dwarf would be expected thus producing one broad dip, which is not seen. There are still many uncertainties concerning absorption dips, being able to determine their nature requires a detailed knowledge of the structure of the stream, the geometry of the system and details on the threading mechanism. Alternatively, the first two narrow dips seen in EK UMa might be due to genuine cessations of accretion. The repeated observation of the dips in the phase range $\phi = 0.0 - 1.3$ means, however, that at least these must be caused by fixed structures.

Three features of the light curve can be used to constrain the system inclination ($i$) and the magnetic colatitude ($\delta$) of the emission region. These are as follows:

1. The duration of the bright phase is given by

$$\Delta \phi \geq \frac{1}{\pi} \cos^{-1}(-\cot \delta \cot i)$$

(where the equality holds if the emission region is small and flat). The light curve
clearly shows $\Delta \phi = 0.63$, leading to a limit for $i$ and $\delta$.

2. The lack of an eclipse in a Roche-lobe-filling CV can provide an upper limit on the inclination of the system, by using the orbital period (114.5 min) and thus radius of the secondary, see Chapter 1 (in this the volume averaged radius). For EK UMa the upper limit on the inclination is $74^\circ$.

3. The presence of absorption dips suggests that the inclination must be greater than the magnetic colatitude.

The combination of these constraints requires the inclination to be in the range $i = 58^\circ - 74^\circ$ and the magnetic colatitude to be in the range $\delta = 36^\circ - 74^\circ$, as shown in fig 2.8. These results are consistent with the original estimates of Morris et al. (1987).

Spectral modelling has shown the blackbody temperature of EK UMa is $50 < kT < 62$ eV. This value is high compared to measurements of the temperature in other AM Her systems (e.g. VV Pup: 23-43 eV, EF Eri: 16-40eV, AN UMa:< 50eV, V834 Cen: 12-22 eV; Osborne 1988). Our confidence in these measurements, which were made by EXOSAT is established by the agreement of the ROSAT and EXOSAT temperatures of the AM Her system UZ For (i.e. ROSAT PSPC:19-25eV, Ramsay et al.1993; EXOSAT CMA: 13-25 eV, Osborne et al.1988).

The high temperature of EK UMa might be taken to suggest that the white dwarf in this system has a high mass. This follows from the consideration of the local Eddington luminosity,

$$L_{\text{Edd}} = 1.3 \times 10^{38} f \left( \frac{M}{M_\odot} \right) > L_{\text{bb}} \text{ \ erg s}^{-1}$$

where $f$ is the emitting fraction of the white dwarf surface and $M$ is the mass of the white dwarf, and the blackbody luminosity $L_{\text{bb}} = 4\pi f \sigma R^2 T^4$. The requirement that the specific flux be sub-Eddington leads to a maximum possible temperature which is a function of the white dwarf radius. Making use of an approximation to the the mass-radius relationship of Nauenberg (1972) (see 1.10 in chapter 1) and the temperature limit of $kT > 50$eV corresponds to a minimum white dwarf mass of $M > 0.55M_\odot$ which is typical of CV WD masses (see fig 2.9). Many polars have been found with periods around 114 mins, including EK UMa, which is known as the period spike. Hameury, King and Lasota (1988) showed,
Figure 2.8: The geometrical constraints on the location of the soft X-ray emission region in EK UMa. The hatched regions are disallowed.
by secular evolution models, that most of the period spike members would have white dwarf masses of 0.6-0.7 $M_{\odot}$. Similarly, Ritter and Kolb (1992) determined the white dwarf mass should be $> 0.7 M_{\odot}$, these mass ranges are consistent with the mass limit determined for EK UMa.

Williams, King and Brooker (1987), however, have pointed out that blackbody fits to the intrinsically complex soft X-ray spectra of AM Her will typically result in temperature over estimates. Nevertheless, they show that a high fitted temperature still implies a high effective temperature in the atmosphere (e.g. 43eV to 25eV and 25eV to 14 eV), and suggests that for EK UMa $kT_{\text{eff}} > 30$eV. In this case, although EK UMa would still be unusually hot compared to other AM Her systems, no useful limit could be put on the white dwarf mass.
The more detailed work of Kylafis and Lamb (1982) and Imamura and Durisen (1983) suggest that high blackbody temperatures are due to high white dwarf masses or high specific accretion rates. These models assume that hard X-rays from the shock are the primary radiation product of the accretion process. The increase in the blackbody temperature in these models is due to increased backscatter of this hard X-ray flux. From the measured X-ray luminosity, the accretion rate can be estimated (see equation 1.6).

In the case of EK UMa this accretion rate is rather low at $\sim 5 \times 10^{14}$ g s$^{-1}$. Imamura and Durisen (1983) showed that for their standard accretion model increasing $M$ led to a corresponding increase in blackbody temperature, therefore for such a low $M$ EK UMa is unlikely to reach such high blackbody temperatures. Thus the model of Imamura and Durisen are unlikely to be directly applicable in this case.

The large soft X-ray excesses seen in many AM Her systems (e.g. King and Watson 1987) have been ascribed to the accretion of blobs which shock well below the white dwarf surface, with the hard X-ray flux from exposed shocks being much reduced and energy being radiated as soft X-rays from the photosphere (Kuijpers & Pringle 1982; Frank, King & Lasota 1988). The hard and soft total fluxes of $(4.5 - 7.2) \times 10^{-12}$ and $5.8 \times 10^{-13}$ erg cm$^{-2}$ s$^{-1}$, respectively, lead to a calculated soft/hard flux ratio of 7.7-12.2, indicating that EK UMa has a very substantial soft X-ray excess. Even if we assumed that the luminosity were a factor of 2 smaller by consideration of the over estimation of luminosity by blackbody models (Williams et al. 1987), this excess would still be present. Recent work by Ramsay et al. (1994) has shown that such soft X-ray excesses are present in most of the polar systems.
Chapter 3

ROSAT constraints on the intermediate polar candidates SW UMa and 1H0709-360.

3.1 Introduction

SW UMa and 1H0709-360 are both X-ray sources which have been claimed to be members of the intermediate polar (IP) class of magnetic CVs. As IPs contain an asynchronously rotating white dwarf it is expected that they would show X-ray modulation at the white dwarf spin period.

Shafter et al. (1986) were the first to propose the IP classification for SW UMa based on optical and X-ray observations in quiescence. They detected an 81.8 min radial velocity period and a 15.9 min optical photometric period, which they also observed in low energy EXOSAT (0.05-2 keV) data. The detection of superhumps by Robinson et al. (1987) showed SW UMa also to be an SU UMa-type system. An EXOSAT observation during superoutburst showed no X-ray modulation at either the 81 or 15.9 min periods (Szkody, Osborne & Hassall 1988). Additionally, the 15.9 min period has not been detected in subsequent optical observations, either when the system was in an anomalous faint state.
(Howell & Szkody 1988) or in a superoutburst state (Kato, Hirata & Mineshige 1992).

1H0709-360 was discovered by Tuohy et al. (1990) as an eclipsing CV with a binary period of 2.44 hr, within the 2-3 hr period gap. These authors suggested, on the basis of its strong HeII emission, the presence of a hard X-ray flux and the possible existence of a second periodicity close to the orbital period, that 1H0709-360 might be an IP close to synchronism.

This chapter described ROSAT observations made with the intention of verifying or discounting the IP nature of these two systems.

3.2 Observations

3.2.1 SW UMa

ROSAT observations of SW UMa were made during two occasions, one on April 11th 1992 and the second on May 3rd 1992. Both PSPC observations used the filter wheel at the open position, and they had a combined exposure time of 6 ks in 3 slots. WFC observations of 2580 and 1040s were recorded through the s1a and s2b (112-200 eV: 60-110 eV) filters, respectively. The observations were performed with the target on-axis.

3.2.2 1H0709-360

1H0709-360 was the target of a short ROSAT PSPC pointing on October 4th 1991, with a total exposure of 8ks during an interval of ~ 7 hrs. The target was observed on-axis with the filter wheel again in the open position.
3.3 Data Analysis and Results

3.3.1 SW UMa

Timing Analysis

The position of source in the PSPC is within 7 arcsec of the optical position. The PSPC lightcurve was extracted using an accumulation radius of 3.5 arc min and a binning of 30s. A background region offset by 12 arc min was used for subtraction as the annulus region had sources present within it. The light curve was also corrected for vignetting and obscuration by the wires. SW UMa was detected at a mean count rate of 0.35 cts s\(^{-1}\) in the PSPC. Although the light curve (fig 3.1) shows no obvious variability within either data section, the target was 27% brighter on April 11 than on May 3. SW UMa was not detected by the WFC in either the pointed observations or the survey. Upper limits (3\(\sigma\)) based on the WFC survey scans are 0.020 cts s\(^{-1}\) (2120s exposure) and 0.021 cts s\(^{-1}\) (1830s exposure) in the s1a and s2b filters respectively. An optical light curve provided by the AAVSO (J. Mattei 1993, private communication) is shown in fig 3.2 and indicates that SW UMa was observed during a quiescent state by ROSAT, beginning about 4d after the end of the previous outburst.

The PSPC data were examined using both Fourier and folding techniques. A Fourier analysis, restricted to periods between 30 and 2000s found no significant power in this range, with a 3\(\sigma\) upper limit on the rms variation of 21\% (for the technique see section 2.2.1 in chapter 2). The accuracy of the 15.9 min period (assuming an uncertainty of ±0.1 min) reported by Shafter et al. (1986) is insufficient, over the 19d separation of the 2 ROSAT observations, to generate phasing of a folded light curve of the combined data, therefore each observation was treated separately. A low order polynomial was subtracted from the data to remove any possible orbital/longer period modulation, the resulting residuals were folded on the 15.9 min period. The tighter measurement yields a 95\% upper limit of 20\% on the semi-amplitude of any 15.9 min modulation. It must be emphasized, however, that neither ROSAT observation spans more than three spin cycles (15.9m) in phase. Although neither ROSAT observation alone samples more than 60\% of the 81 m orbital period, we
Figure 3.1: The ROSAT PSPC light curves of SW UMa. The upper panel shows the data set taken on 1992 April 11, whilst the data set from 1992 May 3 occupies the lower panel.
Figure 3.2: The AAVSO lightcurve of SW UMa spanning an interval of 200 d around the time of the ROSAT observation. The open circles indicate upper limits. Arrows mark the epochs of the ROSAT observations.
tested for its presence by fitting a sinusoid to each section separately. No such modulation
was detected with a 95% upper limit of 58% on the fractional semi-amplitude about the
mean.

Spectral Analysis

A search for spectral features in the SW UMa PSPC data were conducted by computing
a time-resolved curve of the hardness ratio, using the pulse-height channels 11-30/31-200 (i.e. 0.1-0.3 keV/ 0.3-2 keV) in 60s bins. Tests of the data, with various binning
factors, against the hypothesis of constancy revealed no significant variation on time-
scales greater then 60s. As a result, spectral fits were made of the time-averaged spectrum,
using blackbody, bremsstrahlung, power-law and Raymond & Smith models. The results
are displayed in table 3.1. Simple models do not fit the data. Two-component models
provided a better match. An F test showed that the two component models provide a
better representation of the data, with an $F_{obs}$ of 4.7, larger than the 99% confidence, $F_X$
of 2.35. A two-component, Raymond & Smith model with temperatures of $0.37^{+0.07}_{-0.14}$ keV
and $2.8^{+1.5}_{-1.2}$ keV and a common absorbing column, $N_H = 0.4^{+0.4}_{-0.2} \times 10^{19}$ atoms cm$^{-2}$ yielded
the best fit with a $\chi^2$ of 35.4, although other two-component models, e.g. blackbody plus
bremsstrahlung, also provided acceptable fits ($\chi^2 = 37.3, \nu = 28$). Confidence limits on
the two temperatures and the absorbing column of the two-component, Raymond & Smith
model are also presented graphically in fig 3.3. The implied 2-6keV (chosen so a comparison
with the EXOSAT ME can be made) band (absorbed) flux is $1.1 \times 10^{-12}$ erg cm$^{-2}$ s$^{-1}$.
The EXOSAT ME reveals a flux of $3.2 \times 10^{-12}$ erg cm$^{-2}$ s$^{-1}$ for a 70eV blackbody
or $5 \times 10^{-12}$ erg cm$^{-2}$ s$^{-1}$ for a 0.21keV thermal bremsstrahlung (Shafter, Szkody &
Thorstensen 1986).

3.4 1H0709-360

X-ray emission from a point source whose centroid is located within 12arcsec of the re-
ported position of 1H0709-360, was detected at a mean count rate of $0.003 \pm 0.001$ cts s$^{-1}$
(4.1$\sigma$ detection) in the ROSAT PSPC field. This ROSAT observation represents the first
Table 3.1: Spectral parameters of SW UMa for the best fitting blackbody and Bremsstrahlung (separately and together), power-law, and Single and Two Temperature Raymond Smith Models. Uncertainties are based on two parameters of interest and are the 90% confidence values.

### Power law model

<table>
<thead>
<tr>
<th>$N_H \times 10^{19}$</th>
<th>$\alpha$</th>
<th>$\chi^2$, $\nu$</th>
<th>Flux (0.1-2.5 keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>(cm$^{-2}$)</td>
<td></td>
<td>(ergs s$^{-1}$ cm$^{-2}$)</td>
<td></td>
</tr>
<tr>
<td>9.1$^{+4.6}_{-3.8}$</td>
<td>1.78±0.2</td>
<td>56.7, 29</td>
<td>2.9×10$^{-12}$</td>
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</tbody>
</table>

### Bremsstrahlung model

<table>
<thead>
<tr>
<th>$N_H \times 10^{19}$</th>
<th>$kT$</th>
<th>$\chi^2$, $\nu$</th>
<th>Flux (0.1-2.5 keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>(cm$^{-2}$)</td>
<td>(keV)</td>
<td>(ergs s$^{-1}$ cm$^{-2}$)</td>
<td></td>
</tr>
<tr>
<td>5.7$^{+2.8}_{-1.4}$</td>
<td>1.71±0.6</td>
<td>44.9, 29</td>
<td>2.8×10$^{-12}$</td>
</tr>
</tbody>
</table>

### Raymond Smith model

<table>
<thead>
<tr>
<th>$N_H \times 10^{19}$</th>
<th>$kT$</th>
<th>$\chi^2$, $\nu$</th>
<th>Flux (0.1-2.5 keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>(cm$^{-2}$)</td>
<td>(keV)</td>
<td>(ergs s$^{-1}$ cm$^{-2}$)</td>
<td></td>
</tr>
<tr>
<td>0.64</td>
<td>2.69</td>
<td>66.85, 29</td>
<td>2.8×10$^{-12}$</td>
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</table>

### Blackbody and Bremsstrahlung model

<table>
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<tr>
<th>$N_H \times 10^{19}$</th>
<th>$kT_{bb}$</th>
<th>$kT_{br}$</th>
<th>$\chi^2$, $\nu$</th>
<th>Flux (0.1-2.5 keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>(cm$^{-2}$)</td>
<td>(keV)</td>
<td>(keV)</td>
<td>(ergs s$^{-1}$ cm$^{-2}$)</td>
<td></td>
</tr>
<tr>
<td>1.7$^{+3.4}_{-1.6}$</td>
<td>2.3±0.07</td>
<td>2 (frozen)</td>
<td>37.3, 28</td>
<td>2.7×10$^{-12}$</td>
</tr>
</tbody>
</table>

### Two Temperature Raymond Smith model

<table>
<thead>
<tr>
<th>$N_H \times 10^{19}$</th>
<th>$kT$</th>
<th>$kT$</th>
<th>$\chi^2$, $\nu$</th>
<th>Flux (0.1-2.5 keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>(cm$^{-2}$)</td>
<td>(keV)</td>
<td>(keV)</td>
<td>(ergs s$^{-1}$ cm$^{-2}$)</td>
<td></td>
</tr>
<tr>
<td>0.37$^{+2.3}_{-0.37}$</td>
<td>0.37$^{+6.0}_{-0.17}$</td>
<td>2.8$^{+5.9}_{-1.6}$</td>
<td>35.36, 27</td>
<td>2.7×10$^{-12}$</td>
</tr>
</tbody>
</table>
Figure 3.3: 68% (innermost), 95% and 99% (outermost) confidence contours for the parameters of the two-component, Raymond & Smith model applied to the ROSAT spectrum of SW UMa.
X-ray imaging detection of the X-ray counterpart to the optical CV. The source was not detected in the WFC in either the pointed or survey observations. The survey data provide 3$\sigma$ upper limits of 0.01 cts s$^{-1}$ in both the s1a and s2b filters, based on an exposure times of $\sim$2400 and $\sim$2050 s, respectively.

The PSPC light curve of 1H0709-360 was binned into 100s bins. Whilst no significant evidence of systematic variability in the light curve when folded on the 2.444h orbital period identified by Tuohy et al. (1990) is found, the poor statistical quality of the data permits a 100% modulation within the 95% confidence level.

The low PSPC count rate measured for 1H0709-360 precludes a detailed spectral investigation. Nevertheless, some constraints can be placed on the spectral parameters of the source by forming an average hardness ratio from the counts measured in the pulse height channels 41-200 (0.4-2 keV) and 8-40 (0.1-0.4 keV) and then comparing the resulting ratio, 5.79 $\pm$ 7.34, with values expected for a Raymond & Smith plasma model, evaluated over a grid of points in the kT-$N_H$ plane. Fig 3.4 shows the allowed regions of parameter space. While the temperature is essentially unconstrained, at the 95% level, the column density is less than $10^{21}$ atoms cm$^{-2}$, unless the temperature is below about 0.4 keV. For temperatures above 1 keV, the column density is less than $7 \times 10^{20}$ atoms cm$^{-2}$, with 95% confidence. Assuming the temperature to lie in the range 0.1-30 keV, an upper limit on the observed flux at the Earth can be placed on the data of $1 \times 10^{-13}$erg s$^{-1}$ cm$^{-2}$ in the 0.1-2.5 keV band and, by extrapolation, a corresponding 2-10 keV flux of $2.8 \times 10^{-13}$erg s$^{-1}$ cm$^{-2}$.

3.5 Discussion and Conclusions

3.5.1 SW UMa

The ROSAT PSPC observations of SW UMa show no convincing evidence of the 15.9 min modulation that was witnessed by the EXOSAT ME (1-8 keV) data of the source taken during quiescence (Shafter et al. 1986). The 3$\sigma$ upper limit of 21% on the modulation semi-amplitude, is however consistent with the 28 $\pm$ 17% semi-amplitude observed in the EXOSAT data. As such, this data does not exclude the presence of the suggested 15.9 m...
Figure 3.4: Contours of the measured hardness ratio (dashed line) and its 68% (leftmost solid line) and 95% (rightmost solid line) confidence limits for 1H0709-360 are mapped on the grid of such values in the kT-N_H plane for a Raymond & Smith emission model.
Based on an F-test, the spectral profile measured from the PSPC data indicates that a two-temperature representation is preferred with about 90% confidence over one component models. It is interesting to note that the temperatures (0.37 and 2.8 keV) measured for SW UMa in these data are comparable to those (0.7 and 9 keV) derived for the IP, EX Hya (Singh & Swank 1993). A recent ASCA study of EX Hya showed, from the analysis of the intensity ratios of He-like and H-like Kα emission lines of Mg to Fe, that the temperature distribution could not be parameterised by a two emission component model, but rather a distribution of temperatures (Ishida, Mukai, Osborne 1994). Harberl & Motch (1995) found three new IPs (RXJ0028+59, RXJ0153+74, RXJ1712-24) which are characterized by hard X-ray spectra consistent with intrinsically absorbed thermal bremsstrahlung as was seen from IPs known before ROSAT, but are still much harder than SW UMa. They also found three other new IPs (RXJ0558+53, RXJ1914+24) with soft X-ray components similar to those seen in polars, with blackbody temperatures of 40-60 eV. Although these new IPs have soft spectra, they still do not resemble the spectrum observed from SW UMa.

Two component models that have been applied to non-magnetic dwarf novae such as SS Cyg indicate a much lower temperature, blackbody source of ~ 20 eV which varies in luminosity depending on the state, being much stronger during outburst (e.g. Jones & Watson 1992, Cordova et al. 1980). ASCA PV data of SS Cyg during an anomalous outburst was best modelled by a two temperature Raymond and Smith model (kT= 0.8 and 3.5 keV) plus a bremsstrahlung of ~ 18 keV (Nousek et al.1994). Van Teeseling & Verbunt (1994) have recently studied ten different non-magnetic CVs with ROSAT and arrived at a range of two temperature models which are broadly consistent with the temperatures observed in SW UMa (see table 3.2), thus the observed spectrum can give us no clues as to the class of object SW UMa belongs to.
Table 3.2: Temperatures determined from 2T Mewe models for a selection of CVs taken from van Teeseling A. & Verbunt F. (1994).

<table>
<thead>
<tr>
<th>Object</th>
<th>Subclass*</th>
<th>$T_1$(keV)</th>
<th>$T_2$(keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>TY PSA</td>
<td>SU</td>
<td>0.72</td>
<td>12</td>
</tr>
<tr>
<td>IX Vel (1991)</td>
<td>UX</td>
<td>0.48</td>
<td>2.2</td>
</tr>
<tr>
<td>IX Vel (1993)</td>
<td></td>
<td>0.68</td>
<td>&gt; 7</td>
</tr>
<tr>
<td>V3885 Sgr</td>
<td>UX</td>
<td>0.55</td>
<td>3.4</td>
</tr>
<tr>
<td>GP Com</td>
<td>DD</td>
<td>0.14</td>
<td>2.0</td>
</tr>
<tr>
<td>VW Hyi</td>
<td>SU</td>
<td>0.79</td>
<td>4.2</td>
</tr>
</tbody>
</table>

*Subclasses from Ritter (1990), SU= SU UMa type dwarf nova, UX= UX UMa type nova-like and DD = double degenerate.

3.5.2 1H0709-360

This ROSAT PSPC observation was the first imaging X-ray detection of the X-ray counterpart to the optical CV that was associated with the HEAO-1 X-ray source, 1H0709-360. The source is, however, only detected weakly by ROSAT, and thus I am unable to confirm or reject the presence of any orbital modulation. The spectrum is also poorly constrained, but the intervening absorbing column density is likely to be less than $10^{21}$ atoms cm$^{-2}$.

Within the constraints imposed by the ROSAT data, the predicted flux in the 2-10 keV band of $2.8 \times 10^{-13}$ erg cm$^{-2}$ s$^{-1}$ indicated that the point source in the ROSAT image is a factor $\geq 70$ fainter than the source of emission detected by HEAO-1 (Tuohy et al. 1990). We find no X-ray objects visible within the ROSAT field that lie within both the HEAO-1 MC and LASS error boxes (see Tuohy et al. 1990), that could explain this large flux discrepancy. There are three possible explanations; the target, 1H0709-360, may be highly variable, perhaps having been in outburst during the HEAO-1 observation. In the context of IPs, whilst four of the group are known to exhibit outburst behaviour (EX Hya - e.g. Hellier et al. 1989; V1223 Sgr - van Amerongen & van Paradijs 1989; TV Col -Szkody & Mateo 1984; GK Per - e.g. King, Ricketts & Warwick 1979), only in the last case have the X-ray observations been secured in both quiescence and outburst states (Watson, King &
Osborne 1985, Ishida 1992). The implied flux change in 1H0709-360 would be significantly larger than anything witnessed in GK Per during its outbursts, which was a factor of ~10 brighter than quiescence. An alternative scenario is that the HEAO-1 flux is dominated by one or more sources that are not intrinsically luminous in the 0.1-2.5 keV band and/or are heavily absorbed, explaining the lack of a prominent candidate in the ROSAT field. The third possibility is that the spectrum of 1H0709-360 is, in fact, composed of two components, one hard and one soft, with the ROSAT data being dominated by the soft component. If the hard component contributes only very weakly in the 0.1-2.5 keV band, it may be undetectable, leading to a substantial underestimate of the 2-10 keV flux based on an extrapolation of the soft-component spectrum. However, a hard bremsstrahlung or power-law component would have to be heavily absorbed in comparison to the soft component if it were not to dominate the observed 0.1-2.5 keV band flux. For example, a hard bremsstrahlung source with a temperature ~ 30 keV that can account for the HEAO-1 flux but only a fraction to the 0.4-2.0 keV band PSPC count rate requires $N_H \geq 5 \times 10^{22}$ atoms cm$^{-2}$. Higher column densities are required if the temperature is lowered. Other IPs that have been observed at X-ray wavelengths have shown hard X-ray spectra consistent with absorbed Bremsstrahlung emission (e.g. Haberl & Motch 1995).

Summary

These ROSAT observations of SW UMa and 1H0709-360 do not provide conclusive evidence in favour of the proposed IP classification. Continuous, extended X-ray observations of each target will be required to classify these systems unambiguously.
Chapter 4

A ROSAT Observation of the Peculiar Magnetic CV - AE Aqr.

4.1 Introduction

AE Aqr was first found to be a binary system by Zinner (1938). Further spectroscopic measurements by Joy (1954), Crawford & Kraft (1956) and Chincarini & Walker (1981) showed it to be a close binary with a K5V secondary and an orbital period of 9.88 hr. Detection of a 33s period in the optical and X-ray resulted in the IP classification (Patterson 1979, Patterson et al. 1980, see also Eracleous, Halpern & Patterson 1991). This modulation originates from the white dwarf (Welsh, Horne & Gomer 1993), and is interpreted as due to the changing visibility of the magnetically confined accretion regions as the white dwarf rotates. AE Aqr has the shortest spin period of its class. It has also recently been discovered to be rapidly spinning down (τ ~ 2 × 10^7 yrs, de Jager et al. 1994).

AE Aqr is unique in having frequent bright optical flares. These flares can reach as much as three times that of the quiescent level and last for ~ 10 - 60 mins or more. Bruch (1991) found that these flares can be classified into three separate types: 1) pure or predominately continuum flares, 2) pure hydrogen emission line flares and 3) Mixed continuum and lines flares. Recently pulsed TeV gamma ray emission has been observed
with a burst luminosity of $\sim 10^{34}$ erg s$^{-1}$, which is of the same order as the spindown luminosity of $\sim 6 \times 10^{34}$ erg s$^{-1}$ (Meintjes et al. 1994). These gamma ray bursts have been found to be correlated with the onset of optical flares. Strong and variable radio emission was first discovered by Bookbinder & Lamb (1987), and subsequent multiband observations at radio wavelengths (Bastian, Dulk & Chanmugam 1988) suggest that the emission is a result of superposition flarelike particle accretion events. Millimeter observations also show emission consistent with the superposition of flarelike events (Abada-Simon et al. 1993).

A short EXOSAT observation placed an upper limit to the temperature of 1.8keV on any optically thin X-ray emission from AE Aqr (Osborne 1990). Similarly an Einstein IPC measurement gave an average optically thin emission temperature of 1keV (Eracleous, Halpern & Patterson 1991). These X-ray temperatures are much lower than those measured for other DQ Her type systems, which have temperatures $\sim 10 - 30$keV (Ishida 1991). Up to this time no simultaneous optical data have been recorded, so it is unknown whether optical flares are correlated with flares in the X-ray or whether flares occur at all in the X-ray.

Many of the properties of AE Aqr show it to be a highly unusual CV. Simultaneous multiwaveband measurements were needed to start solving some of the mysteries of AE Aqr such as the origin of the flares. In an aim to to this a World Astronomy Day and Whole Earth Telescope (Nather et al. 1990) runs were organised for October 1993. This chapter describes the ROSAT observation that took place during this campaign and some of the simultaneous optical and HST UV photometry.

4.2 Observations

4.2.1 ROSAT PSPC observations

The data were taken using the ROSAT PSPC: 0.1-2.5 keV, (Trümper 1984) in 3 observations over a 3 day period during October 1993 (see Table 4.1). The observations were performed on-axis and had a total exposure time of 15 ksec. AE Aqr was detected at an average count rate of 1.3 cts s$^{-1}$. 
Table 4.1: ROSAT PSPC Observation Log for AE Aqr.

<table>
<thead>
<tr>
<th>Section</th>
<th>Start Date</th>
<th>Start Time (UT)</th>
<th>Duration</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td>(h m s)</td>
<td>(m s)</td>
</tr>
<tr>
<td>i</td>
<td>20 Oct 93</td>
<td>01 27 11</td>
<td>09 36</td>
</tr>
<tr>
<td>ii</td>
<td>20 Oct 93</td>
<td>04 40 02</td>
<td>23 43</td>
</tr>
<tr>
<td>iii</td>
<td>20 Oct 93</td>
<td>06 17 19</td>
<td>26 28</td>
</tr>
<tr>
<td>iv</td>
<td>20 Oct 93</td>
<td>19 09 34</td>
<td>25 07</td>
</tr>
<tr>
<td>v</td>
<td>20 Oct 93</td>
<td>20 47 36</td>
<td>22 40</td>
</tr>
<tr>
<td>vi</td>
<td>21 Oct 93</td>
<td>01 19 19</td>
<td>11 28</td>
</tr>
<tr>
<td>vii</td>
<td>21 Oct 93</td>
<td>02 56 04</td>
<td>21 59</td>
</tr>
<tr>
<td>viii</td>
<td>21 Oct 93</td>
<td>04 33 35</td>
<td>24 13</td>
</tr>
<tr>
<td>ix</td>
<td>21 Oct 93</td>
<td>06 10 35</td>
<td>26 41</td>
</tr>
<tr>
<td>x</td>
<td>21 Oct 93</td>
<td>19 03 04</td>
<td>25 31</td>
</tr>
<tr>
<td>xi</td>
<td>22 Oct 93</td>
<td>01 12 34</td>
<td>12 14</td>
</tr>
<tr>
<td>xii</td>
<td>22 Oct 93</td>
<td>02 49 50</td>
<td>21 57</td>
</tr>
</tbody>
</table>
The 3 data sets, consisting of 12 continuous intervals in total, were extracted using the Starlink ASTERIX (Saxton 1992) analysis package. An accumulation radius of 2.7 arcmin with a background annulus around, but separated from, the source was used for the extraction of data. In this way no events were lost through electronic ghost imaging (see Appendix A).

As the observation was performed with AE Aqr on-axis, care has to be taken in the interpretation of light curves. Fine wire structures are present on the instrument which can occult a celestial X-ray source. To avoid permanent occultation by a wire, the spacecraft does not point steadily at the source, but is 'wobbled' by ±3' with a period of 400sec, thus producing a source flux modulation at this period and its harmonics (see appendix A for a more detailed description).

AE Aqr was simultaneously observed with the ROSAT WFC (Simms et al. 1990) using the S2b filter (61-111 eV). A total exposure of 17.8 ksec was obtained during the three observations. The data were corrected for exposure, vignetting and secular efficiency degradation. AE Aqr was not detected, giving a 3σ upper limit of $4.5 \times 10^{-3}$ cts s$^{-1}$.

4.2.2 Optical and HST UV data

Simultaneous optical photometry at 4 s resolution were taken, by the WET team, in blue light using the 1.9m telescope at the SAAO on the 20 and 21 October and at the McDonald telescope on the 22nd (O'Donoghue et al. 1995). The observations were made using unfiltered light and therefore had varying bandpasses at the different telescopes. Both telescopes monitored a nearby A star which was used for calibrating for this effect and atmospheric extinction. The data sets taken at the SAAO and the McDonald telescope were simultaneous with PSPC sections v, x and xii respectively and had overlap times of ~ 670, 1500 and 1300 secs.

AE Aqr was observed by the HST's Faint Object Spectrograph on October 21st 1993. The observation lasted for 11.4 hours and was divided into 8 slots (Eracleous et al. 1995). During this time ~ 30 mins of data simultaneous with the ROSAT observation were recorded. I am only going to deal with this simultaneous data. This HST data consists of
four slots coincident with the ROSAT sections (vi), (vii), (viii) and (ix) (see fig 4.1). Two light curves are recorded for the HST data. One is the mean UV flux taken by averaging the observed flux from several emission line-free regions between 1300 and 2100 Å. The other is the zeroth order light curve from a passband of width 1900Å centered on 3400Å.

4.3 Data Analysis and Results

4.3.1 ROSAT Timing Analysis

Light curves, extracted initially using the full energy band, were background subtracted and corrected as usual for vignetting, deadtime and a constant term for the reduction in counts due to the wires (see fig.4.1). These light curves show that AE Aqr spent most of the time at a count rate of ~ 1 cts s⁻¹, with occasional excursions to higher count rates. There is one well delineated flare, in section iii, starting at ~ 6.3 hrs UT on October 20th, which lasted for ~ 20 minutes and peaked at 3.5 cts s⁻¹. Smaller flares may be present in sections ii, x and xi.

The data have been folded at the orbital period (Porb=9.88hr), see fig 4.2. Phase coverage was poor, with no phase in the orbit being covered more than once. The apparent orbital modulation is dominated by the flares.

The data were extracted into 1 second bins in order to search for periods around 33 seconds and their harmonics. After heliocentric correction of each data section a fast Fourier transform (FFT) was performed on the entire data set. The resulting FFT can be seen in fig. 4.3. These clearly show the presence of a signal at 33.0781 ±0.0021s, consistent with the optical period of 33.0767 s (de Jager et al. 1994). Fig 4.3 also shows a small peak at the 16.5 s optical period. Note, that folding at this period (see later) shows this to be insignificant. Orbital sideband periods are absent, as is shown in fig 4.3. A further FFT was performed on the window function to determine the effect of aliasing. This showed that the window function produces power in the range 0.001-0.015 Hz, out of the region of interest. This procedure was also followed for soft (0.1-0.45 keV) and hard (0.45-2 keV) bands and the resulting FFTs can be seen in fig. 4.4 and 4.5. Again the 33.0767s period
Figure 4.1: The ROSAT soft X-ray light curve of AE Aqr plotted in 1 minute bins. The spacecraft wobble causes periodic partial modulation at 100s, resulting in non-Poisson bin to bin variability.
Figure 4.2: The ROSAT soft X-ray light curve of AE Aqr folded on the orbital period ($P_{orb}=9.88$ hr). Phase zero corresponds to inferior spectroscopic conjunction of the K5V star. Any secular orbital modulation appears to be swamped by large flares.
Table 4.2: Fits to Data Folded at various Periods

<table>
<thead>
<tr>
<th>Period (s)</th>
<th>Const. $\chi^2, \text{dof}$</th>
<th>Const. &amp; Sine $\chi^2, \text{dof}$</th>
<th>Const. $^1$</th>
<th>Amplitude $^1$ (errors are for $\Delta \chi^2 = 2.3$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>33.0767</td>
<td>187, 31</td>
<td>23.7, 29</td>
<td>1.21</td>
<td>$0.17 \pm 0.02$</td>
</tr>
<tr>
<td>16.5380</td>
<td>8.43, 15</td>
<td>3.81, 13</td>
<td>1.21</td>
<td>$&lt; 0.05$</td>
</tr>
<tr>
<td>33.0153</td>
<td>42.9, 31</td>
<td>20.6, 29</td>
<td>1.21</td>
<td>$&lt; 0.08$</td>
</tr>
<tr>
<td>33.0460</td>
<td>29.3, 31</td>
<td>24.5, 29</td>
<td>1.21</td>
<td>$&lt; 0.05$</td>
</tr>
<tr>
<td>33.1070</td>
<td>75.8, 31</td>
<td>33.3, 29</td>
<td>1.21</td>
<td>$&lt; 0.11$</td>
</tr>
<tr>
<td>33.1380</td>
<td>27.4, 31</td>
<td>14.4, 29</td>
<td>1.21</td>
<td>$&lt; 0.07$</td>
</tr>
</tbody>
</table>

1 Units are cts s$^{-1}$.

was detected to a similar accuracy and the 16s period was absent.

The data were folded at various interesting periods and fitted with a constant plus a sinusoid to quantify the level of modulation. The results are presented in Table 4.2 (the amplitude being the semi-amplitude of the sinusoid). The only period strongly detected is that at 33.0767 s with modulation at the 16.5 s period being $< 29\%$ of the 33.0767 level. Modulation at the other periods is either not detected, or is detected at a level consistent with random variability in the source. The quiescent data is seen folded on the quadratic ephemeris of de Jager et al. (1994) in figure 4.6. It clearly shows that the modulation is single peaked, unlike the double peaked modulations reported for the optical and UV (Robinson et al. 1991, Eracleous et al. 1994).

Power spectra of individual data sections showed differing levels of 33s power. This prompted an investigation of the amplitude of this modulation as a function of count rate. The data were grouped into 400s sections so as to remove effects due to the wires, with the times of the events in each section corrected for the motions of the Earth around the sun and the motion of the white dwarf in the binary system. This latter variation was taken from the recent optical determination of the delay of the spin pulses (which may be as much as $\pm 2$ sec) by Eracleous et al. (1994) and the orbital ephemeris of Welsh, Horne & Gomer (1993). (The low statistical quality of the data do not require this correction). Each 400s section of data was then folded using the spin period of 33.0767 s. We have fit-
Figure 4.3: Upper panel: The power spectrum of the entire ROSAT PSPC data set for frequencies 0 to 0.1 Hz. This clearly shows the presence of periods around 33 s and the lack of any period at 16.5 s. Aliasing of the observation slots causes the power at the lower end of the power spectrum. The feature seen at 0.02 Hz is due to the wire structure. The lower panel: A high resolution power spectrum around the spin period. The dominant period is coincident with the optical period at 33.0767 s. Orbital sideband periods are also shown.
Figure 4.4: Hard band (0.45-2 keV) power spectra shown at the same resolutions as the previous figure. Again the most dominant period is the 33.0767 period, with neither the 16.5s or orbital sideband periods being detected. The modulation produced by the wires appears to be greatly reduced at this energy, as the count rate is lower in this band.
Figure 4.5: Soft Band (0.1-0.45 keV) power spectra shown at the same resolutions as the previous two figures. The only 2 periods detected are the 33s spin modulation of the white dwarf and the 0.02 Hz power spike of the wire structure modulation.
Pulse Modulation Profile during Quiescence

Figure 4.6: The spin modulated profile of AE Aqr during quiescence, folded on the quadratic ephemeris of de Jager et al. (1994) (the error in the ephemeris is $\phi \approx 0.04$). It can be clearly seen that the profile is single peaked and broad.

This analysis was repeated for two ROSAT PSPC observations of AE Aqr from the public archive (Nov 92 : 20 Ksec, Apr 91 : 3 Ksec). The results from these data sets are also shown in fig 4.7. A non-parametric correlation test (Kendall's $\tau$) shows the data to be positively correlated to a confidence of 98.9%. Taking out the flare (brightest) point reduces the correlation, but it is still present at 98.2% confidence.

Energy resolved light curves at 1s resolution using the hard and soft bands of 0.45-2keV
Figure 4.7: The absolute amplitude of the 33sec modulation as a function of the unmodulated count rate derived from fits to 400s data groups. The error bars correspond to ±68% confidence for 2 parameters, and are shown for the October 1993 data set only. Data from the archive is shown as circles (Nov92) and squares (Apr 91). Errors for these data points are comparable.
and 0.1-0.45keV respectively have been extracted. Using this data times series of the hardness ratio (hard/soft) were created. These were again folded on the quadratic ephemeris of the white dwarf spin of de Jager et al. (1994) after correction for the time delays described above. These folded light curves were fitted with a constant plus a sine wave. Modulation of the hardness ratio was shown by an F test ($F(99.9\%)=2.94$, $F_{obs} = 11.4$), to be present during quiescence (sections iv-ix in fig 4.1), with maximum hardness at the time of minimum intensity (see fig. 4.8). This hardness ratio variation could not be detected during the flare ($F_{obs} = 2.0$, section iii in fig 1), although it could still be present at just below the quiescent modulation fraction. Folded hardness ratios are presented in fig. 4.8. Table 4.3 shows the results of the sine fits.
4.3.2 ROSAT Spectral Analysis

A spectrum was extracted from those sections of data in which no flares could be seen (sections iv-ix in figure 4.1). This was binned into 32 pha bins having at least 10 counts per bin. A systematic error of 2% was added to each spectral bin to account for uncertainties in the PSPC gain. The data was then compared to various spectral models, all including a single absorption component (Morrison and McCammon 1983), using the XSPEC Spectral Analysis Package (Shafer et al. 1991). The best fit values are given in table 4.4. The only model which gave an acceptable fit to the data was a two-temperature optically thin plasma emission model (Raymond & Smith 1977) having temperatures of 0.21 and 0.99 keV. Full details of the best fit parameters are given in table 4.5, the spectral fit is illustrated in fig. 4.9.

The spectrum of AE Aqr during the flare (section iii in fig. 4.1) was also extracted, and was also fit with the two temperature Raymond & Smith model. The results of the fits are shown in table 4.5. A contour plot (see fig 4.10) shows that the lower, and possibly the higher, temperature component does rise during the flare.

Separate spectral fits to the bright and faint phases of the 33 s modulation during quiescence did not reveal any significant differences (see table 4.6).

For a distance to AE Aqr of 100pc (Welsh, Horne & Oke 1993) and for a PSPC count rate

1 used the PSPC response matrix appropriate to the data of the observation (i.e. DRM 36), results obtained with the other response matrix (DRM 06) were not significantly different.
AE Aqr fitted with a 2T Raymond & Smith Model

Figure 4.9: The ROSAT PSPC spectrum of AE Aqr during quiescence. The crosses show the data, the solid line shows the best fit model folded through the response of the instrument. The contributions of the 0.2 and 1 keV optically thin emission components are shown dotted.

Table 4.4: Spectral fits for various models during quiescence

<table>
<thead>
<tr>
<th>Model</th>
<th>$\chi^2, \text{dof}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Bremsstrahlung</td>
<td>224, 29</td>
</tr>
<tr>
<td>Power Law</td>
<td>377, 29</td>
</tr>
<tr>
<td>Raymond &amp; Smith</td>
<td>536, 29</td>
</tr>
<tr>
<td>Soft Blackbody plus Bremsstrahlung</td>
<td>49.9, 27</td>
</tr>
<tr>
<td>Power law temperature distribution (R &amp; S)</td>
<td>49.3, 27</td>
</tr>
<tr>
<td>2 T Raymond &amp; Smith</td>
<td>28.1, 27</td>
</tr>
</tbody>
</table>
Table 4.5: Spectral Fit during Quiescence and Flare for a two temperature Raymond & Smith Model.

<table>
<thead>
<tr>
<th>Model Parameter / Statistic</th>
<th>Quiescent Value</th>
<th>Flare Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>$N_H$ ($10^{19}$ atoms cm$^{-2}$)</td>
<td>$3.70_{-2.06}^{+2.41}$</td>
<td>$1.77_{-0.77}^{+4.41}$</td>
</tr>
<tr>
<td>$kT$ (keV)</td>
<td>$0.20_{-0.03}^{+0.05}$</td>
<td>$0.31_{-0.10}^{+0.12}$</td>
</tr>
<tr>
<td>($\times 10^6$K)</td>
<td>$2.32_{-0.35}^{+0.38}$</td>
<td>$3.60_{-1.39}^{+1.16}$</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2\text{keV})$($\times 10^{-12}\text{ergs cm}^{-2}\text{s}^{-1}$)</td>
<td>$2.49_{-0.15}^{+0.21}$</td>
<td>$5.88_{-0.67}^{+0.08}$</td>
</tr>
<tr>
<td>$kT$ (keV)</td>
<td>$1.05_{-0.09}^{+0.06}$</td>
<td>$1.36_{-0.41}^{+2.28}$</td>
</tr>
<tr>
<td>($\times 10^6$K)</td>
<td>$12.3_{-1.9}^{+0.9}$</td>
<td>$15.8_{-4.6}^{+2.6}$</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2\text{keV})$($\times 10^{-12}\text{ergs cm}^{-2}\text{s}^{-1}$)</td>
<td>$3.56_{-0.20}^{+0.09}$</td>
<td>$5.99_{-0.16}^{+1.17}$</td>
</tr>
<tr>
<td>$\chi^2, dof$</td>
<td>28.1, 28</td>
<td>37.1, 28</td>
</tr>
</tbody>
</table>

Table 4.6: Spectral Fits for the Bright and Faint Phases of the 33 s Modulation.

<table>
<thead>
<tr>
<th>Model Parameter / Statistic</th>
<th>Bright Phase Value</th>
<th>Faint Phase Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>$N_H$ ($10^{19}$ atoms cm$^{-2}$)</td>
<td>$3.19_{-3.19}^{+2.32}$</td>
<td>$3.43_{-3.43}^{+3.45}$</td>
</tr>
<tr>
<td>$kT$ (keV)</td>
<td>$0.20_{-0.03}^{+0.04}$</td>
<td>$0.21_{-0.06}^{+0.10}$</td>
</tr>
<tr>
<td>($\times 10^6$K)</td>
<td>$2.32_{-0.35}^{+0.38}$</td>
<td>$2.44_{-0.4}^{+1.16}$</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2\text{keV})$($\times 10^{-12}\text{ergs cm}^{-2}\text{s}^{-1}$)</td>
<td>$2.76_{-0.23}^{+0.07}$</td>
<td>$2.42_{-0.39}^{+0.3}$</td>
</tr>
<tr>
<td>$kT$ (keV)</td>
<td>$1.06_{-0.13}^{+1.47}$</td>
<td>$1.02_{-0.19}^{+0.42}$</td>
</tr>
<tr>
<td>($\times 10^6$K)</td>
<td>$12.3_{-1.5}^{+1.1}$</td>
<td>$11.8_{-2.2}^{+4.7}$</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2\text{keV})$($\times 10^{-12}\text{ergs cm}^{-2}\text{s}^{-1}$)</td>
<td>$3.90_{-0.17}^{+0.21}$</td>
<td>$2.68_{-0.43}^{+0.19}$</td>
</tr>
<tr>
<td>$\chi^2, dof$</td>
<td>27.2, 25</td>
<td>20.1, 25</td>
</tr>
</tbody>
</table>
Figure 4.10: Temperature confidence contours (68%, 90% and 99.7%) derived from the quiescent and flare (dotted) spectra.
of 1.0 cts s\(^{-1}\) (the minimum observed), the total X-ray luminosity of the two temperature plasma is \(L = 4\pi f d^2 = 1.5 \times 10^{39}\text{ ergs cm}^{-2}\text{s}^{-1}\).

### 4.3.3 Optical- X-ray comparison

The simultaneous X-ray and WET data

The three optical observations are examined separately as there is still some uncertainty in their absolute flux calibration. The first observation starts half way through slot v (see fig 4.1) of the X-ray data. Fig 4.11, plotted on an arbitrary x scale, clearly shows the presence of 2 large flares, which exhibit some flickering. The simultaneous X-ray data coincides with the rise of the first flare and a very slight increase in the X-ray is observed. The second observation (see fig 4.12) also exhibits flares and these are correlated with X-ray flaring activity. The final observation has a quiescent phase followed by a large flare (see fig 4.13). Unfortunately the simultaneous X-ray data only covers the quiescent phase, little or no modulation is seen either in the optical or the X-ray at this time.

The three data sets were then each folded over the spin period of the white dwarf using the same quadratic ephemeris as before. Fig 4.14 clearly shows that the folded light curves each have a different morphology. The first data optical fold shows a single bright and faint phase, with minimum occurring at \(\phi \sim 0.5\). The second fold has a minimum at the same phase as the first observation, though the shape of the peak is different. The final fold shows two peaks, the 16.5s modulation being present, the main minima again at \(\phi \sim 0.5\) and the second minima at \(\phi \sim 0.9\). These folds include flaring activity. The differing forms of modulation may be related to this flaring. The level of modulation in all the cases is very low ranging from 0.1-0.35%.

The simultaneous X-ray and HST data

The four UV and zero-order HST data sections were each compared with the X-ray light curves. Fig 4.15 shows the result of this comparison. It is also clear that the X-ray flux in the main is correlated with the UV flux, the best example of this can be seen in the
Figure 4.11: Upper panel: Optical light curve taken on 20th October '93. The two lines indicate the position of the simultaneous ROSAT data. Lower panel: Comparison of the simultaneous optical and X-ray data.
Figure 4.12: Upper panel: Optical light curve observed on the 21st October. The vertical lines indicate the start and end of the ROSAT observation. Lower panel: Comparison of the optical and X-ray data clearly showing a correlation between the two.
Figure 4.13: Upper panel: The optical light curve for the 22nd October '93. Again the vertical lines indicating the start and end of the ROSAT observation. Lower panel: Optical and X-ray comparison light curves, little activity is observed in either the X-ray or the optical.
Figure 4.14: 33s folds for each of the three optical observations using the ephemeris of de Jager et al. (1994). The minimum occurs at the same phase for each observation, but the modulation appears different in each case.
bottom left panel in fig 4.15. It is unfortunate that no major flares occurred during these observations.

The two HST bands were then folded on the spin ephemeris of de Jager (1994) and the resulting folds were then compared to the X-ray fold. Fig 4.16 clearly shows the double peaked nature of the UV modulation, the minimum of the X-ray pulse coincides with the smaller of the UV peaks and the main pulse in the UV is broadly consistent with the X-ray pulse.

Eracleous et al. (1995) showed that the spin modulation in the UV is very high, reaching an amplitude of up 50% of the mean quiescent level. This modulation amplitude was also shown (like the X-ray) not to vary with mean intensity (see fig 4.17).

4.4 Discussion

In this observation X-ray flaring has been observed for the first time, one flare being particularly prominent with a three fold increase in flux and lasting for approximately 20 mins. I have shown that the X-ray period, $P_x = 33.0781 \pm 0.0021$ s, is the same as the optical period, which must be the spin period of the white dwarf. No power in excess of local noise is seen at $P_x/2 = 16.5$ s or at any of the orbital sideband periods. The amplitude of the 33.08 s modulation increases with intensity, but only very weakly, the modulation fraction decreases with increasing intensity. Folding the quiescent hardness ratio at $P_x$ shows that the spectrum is harder during the light curve minimum, in common with other IP systems. The spectrum is best fit by a two temperature optically thin plasma emission model with temperatures 0.2 and 1.0 keV. Slight temperature differences are measured between quiescence and the flare. The spectral analysis of the spin cycle bright and faint phases support the hardness ratio change as the $N_H$ changes however the increase on its own is not enough to reduce the bright phase flux to the faint phase flux. It has also been shown that on the whole the X-ray and optical activity are correlated.

Clearly this ROSAT observation has presented with some problems to be addressed. These are:
Figure 4.15: The simultaneous HST and PSPC light curves. The zeroth-order represents a band center of 3400Å and a width of 1900Å and the UV represents the UV continuum (Emission line free regions from 1300-2100Å). The top left and right panels are coincident with slots (vi) and (vii), and the bottom left and right panels are simultaneous with slots (viii) and (ix) in the ROSAT light curve respectively. A binning of 60 s is used in each plot.
Figure 4.16: Spin folded light curves. Upper panel: the zero-order fold i.e. center wavelength of 1300Å and a width of 1900Å. Middle Panel: UV continuum fold. Lower Panel: the soft X-ray fold. All these use the ephemeris of de Jager et al. (1994).
Figure 4.17: Upper panel: The pulsed fraction of the major and minor peaks of the 33 sec UV modulation. Lower Panel: The mean level UV light curve. This figure is taken from Eracleous et al. 1995).

- The spectrum is quite different to that seen from other IP systems. The temperatures are low, whereas one would expect a harder spectrum with a X-ray Bremsstrahlung emission component of temperatures of 10keV and upwards (Ishida 1991). The spectrum does, in fact, resemble those observed from coronally active systems such as RS CVns.
- The flaring behaviour, being both frequent and bright, does not fit the pattern of IPs.
- The spin modulated count rate increases only very weakly with increasing average intensity. Intuitively, one might expect, if accretion were the cause of any X-ray emission, the amplitude of modulation would have a strong relationship with average intensity.

4.4.1 Coronal Emission from Secondary Star?

The brightest coronal emission systems are the binary RS CVns. Hall (1976) defined RS CVns to be binaries with a hotter component (spectral type F- to G- and luminosity class
V or IV) and cooler component (usually a subgiant or giant K-type star) which is the more massive component. These systems exhibit strong CaII H+K, strong coronal X-ray, radio and flare emission.

Surveys of groups of stars such as the Pleiades (Stauffer et al. 1994) and IC 2602 (a young open cluster, Randich et al. 1995) have studied their coronal X-ray emission. Stellar luminosity in these systems increases with increasing rotation until a certain level where the emission seems to 'saturate'; this tends to occur at log \(L_x/L_{bol} = -3.0\), see fig 4.18 for an example of this taken from the young open cluster IC 2602 (Randich et al. 1995). This saturation has been ascribed to the filling of all the available surface area by coronal activity. RS CVn can reach higher luminosities, in the brightest systems coronal loops are occurring between the two stars, hence increasing the emission volume. Exceptions to this saturation level have been observed (e.g. HD 22403 (Dempsey et al. 1993b) and the star R95B in Randich et al. (1995)) but these systems have doubts surrounding their classification.

Addressing the question of the spectrum, does the similarity to that of coronally active systems indicate that we may be observing X-ray emission originating at the secondary rather that accretion onto the white dwarf? Typical RS CVn spectra are best fit by two temperature Raymond and Smith models of temperatures 0.1 \(- 3 \times 10^6\) and 10 \(- 40 \times 10^6\) K (Dempsey et al. 1993a), which is consistent with the AE Aqr spectrum. Luminosities of these coronally active systems are typically lower than the total X-ray luminosity of AE Aqr \((1.5 \times 10^{31} \text{ergs cm}^{-2} \text{s}^{-1})\), the brightest dwarf RS CVn having a luminosity of \(1.0 \times 10^{32} \text{ergs cm}^{-2} \text{s}^{-1}\) (BD+25380, Dempsey et al. 1993b). Although AE Aqr is brighter than these systems it is not inconceivably bright, especially considering that the presence of a spin modulation does mean that some portion of the X-ray flux must be emanating from the white dwarf and not the secondary. Therefore I cannot rule out the possibility that some of the X-ray flux is arising from coronal activity from the secondary. The modulation fraction in the UV is higher than in the X-ray, this might be expected if coronal emission were weaker in the UV; hence a higher modulation as the proportion of accretion to coronal emission is greater.

The existence of flaring activity also may indicate that the secondary is active with the
Figure 4.18: Log of the X-ray luminosity to bolometric luminosity ratio vs. $B - V$ colour for objects in the young open cluster IC 2602 (taken from Randich et al. 1995). Only objects with known $B - V$ have been plotted. AE Aqr is also indicated using an X-ray luminosity allowing for 28% of the flux being due to accretion and assuming a standard $B - V$ for a K5 V secondary.
large flare (section iii fig. 4.1) having a profile with a steep rise and slow decline, similar to that seen in stellar flares. Ishida (1991) showed many IP light curves absent of flaring. Temperature variations during stellar flares are observed for some systems, where the lower temperature stays constant and the hotter component rises (Haisch et al. 1991). This is not observed in AE Aqr as the low temperature is shown to rise and the high temperature may also rise. An increase of spin modulation during the largest flare argues against a stellar origin. Observed UV flares have been shown not to be coronal flares because the emission line radial velocity curve does not track the motion of the companion, the velocities implied by the widths of the lines are greater than those observed for stellar flares and the continuum and emission lines decline in step, unlike stellar flares where the continuum declines much faster than the lines (Eracleous et al. 1993). In contrast O'Donoghue (1995) have shown that the optical spin modulation is a function of mean intensity and hence increases during the flares. This suggests that the optical flares originate on or near the white dwarf. IUE measurements taken during the WAD/WET campaign (de Martino, Wamsteker & Bromage 1995) confuse matters. N V, Si IV and HeII emission lines are diagnostic of transition-subcoronal regions while O I and Si II are representative of chromospheric regions. The enhancements for these lines during the flares are found to consistent with those observed for RS CVn systems. However, the far UV line luminosity in the flares of AE Aqr are greater by factors of 2-7 then the most active systems V711 Tau and UX Ari (Linsky et al. 1989).

4.4.2 AE Aqr as an IP

Imamura and Durisen (1983) modelled the accretion onto magnetic white dwarfs in CVs and found that the resulting spectrum should consist of a hard X-ray Bremsstrahlung component of ~ 30 keV from the post-shock region and a soft component of reprocessed bremsstrahlung from the white dwarf. Kylafis and Lamb (1982), when modelling non/weakly magnetic systems also found there to be a hard X-ray component. These hard X-rays are produced when the accreting matter shocks, resulting in a gas of very high temperatures which cool via bremsstrahlung emission. They found that the observed temperature depended on both mass and accretion rate, which was less than the emission region temperature. Observations by Ginga have shown IPs to have hard X-ray
spectra which are highly complex and often highly absorbed (Ishida 1991). Unlike both predictions and observations of IPs AE Aqr is a source of only relatively low X-ray temperature. ASCA satellite observations have found a range of temperatures in FO Aqr and EX Hya (Ishida, Mukai & Osborne 1994 and Mukai, Ishida & Osborne 1994). This range of temperatures does not go as low as those observed in AE Aqr but this may be due to instrumental effects. More recently some IPs have been shown to have additional very soft spectral components as well as the more usual hard component (~ 50eV and 20keV in RE0751+14; Duck et al. 1994) (Haberl et al. 1994), but AE Aqr does not show temperatures as low as these.

From the measured X-ray luminosity a mass accretion rate can be determined. Assuming that the accretion process is 100% efficient leads us to a mass accretion rate of $8 \times 10^{13} \text{g s}^{-1}$. This is very low, most CVs have accretion rates of the order of $10^{16} \text{g s}^{-1}$. From this mass accretion rate it is possible to determine the corresponding shock height (Frank, King & Raine 1992)

$$D \approx \frac{1}{3} \times 9 \times 10^{8} M_{\odot} f_{-2} M_{\odot}^{2} R_{\odot}^{\frac{5}{3}} \text{cm}$$

where $f_{-2} = \text{fraction of the white dwarf surface area on which accretion occurs}$ units of $10^{-2}$, $R_{\odot}$ is the radius of the primary in units of $10^{8} \text{cm}$, $M_{\odot}$ is the mass of the primary in solar masses and $M_{\odot}$ is the mass accretion rate in units of $10^{15} \text{g s}^{-1}$. This leads us to $D \approx 2.0 \times 10^{10} \text{cm}$ using $f_{-2} = 1$, $M_{\odot}=0.91M_{\odot}$ (Reinsch et al. 1994) and $R_{\odot} = 0.6 \text{cm}$ (using the mass-radius relationship for WD of Nauenberg (1971)). The resulting shock temperature can be found by using the relation (Imamura and Durisen 1983)

$$kT_{\text{sh}} = 62 M_{\odot} \left( \frac{R_{\text{shock}}}{5 \times 10^{8} \text{cm}} \right)^{-1} \text{keV}$$

where $R_{\text{shock}}$ is the shock height from the center of the white dwarf. In this case the accretion rate results in a shock temperature of $kT_{\text{sh}} = 1.21 \text{keV}$. This value is of the same order of magnitude as is observed.

The co-rotation radius from Kepler's law for AE Aqr is $1.4 \times 10^{9}$. Expressing 1.18 in terms of accretion luminosity rather than $\dot{M}$ gives the Alfvén radius as

$$R_{\text{Alfvén}} = 5.5 \times 10^{8} M_{\odot}^{\frac{5}{2}} R_{\odot}^{\frac{5}{2}} \mu_{50}^{\frac{1}{2}} \text{cm}$$

where $\mu = B_{*} R_{\odot}^{2}$. This results in $R_{\text{Alfvén}} = 2.7 \times 10^{10} \text{cm}$ (using $B=6 \times 10^{4} \text{G}$ inferred
from spin behaviour (Bookbinder & Lamb 1987\(^2\)). Thus threading onto the field lines cannot occur because the Alfvén radius is larger than the co-rotation radius as material is moving too quickly to be threaded.

Eracleous et al. (1994) determined the HST best fit parameters for the temperature of the unilluminated parts of the white dwarf surface, the maximum temperature at the tip of the polar cap, the height of the illuminating X-ray source, \( H \equiv D/R \), magnetic colatitude and the limb darkening coefficient. From these they determined that the X-ray luminosity was equal to,

\[
L_X = 3 \times 10^{37} \eta^{-1} (H R_g)^3 \text{ erg s}^{-1}
\]

(where \( \eta \) is the efficiency and \( H \) was determined, from entropy mapping, to be in the range 2-4. For the accretion rate observed this requires an efficiency of 115%. The determination of the shock height requires even greater efficiencies. This means that some of the luminosity observed in AE Aqr is probably not produced in a shock, thus suggesting that perhaps blobby accretion is occurring where energy is deposited within the white dwarf and diffuses back to the surface to produce broad hotspots which are observed in the UV.

The 'bombardment' model for low accretion rates onto polars (see section 1.1.6 for a description) may also explain the low X-ray temperatures that are observed. This model is a function of the white dwarf mass, magnetic field and the specific accretion rate (Beuermann 1992, 1993). For AE Aqr's parameters the temperatures expected from this model are still much higher than observed due to the low magnetic field in IPs. Thus this model cannot be used to explain the observed spectrum.

Spin modulations in IPs were considered by King and Shaviv (1984), they assumed that the accretion region had negligible height and occurred over a small area near the pole. They found that the shape of the occultation modulation depended on the fractional area \( f \), the magnetic co-latitude and the line of sight inclination. In general modulations could be explained if the fractional surface area of hard X-ray emission was \( f \sim 0.1 \). Rosen, Mason and Cordova (1988) arrived at an alternative suggestion for the production of spin modulation in IPs. Their model considered that the accretion flow formed a curtain of

\(^2\)There is much uncertainty about the magnitude of the B field in AE Aqr, this should be borne in mind during the following discussion.
material down to the magnetic pole. The resulting X-ray emission from this accretion curtain would have a minimum caused by the absorption and/or scattering by the matter in the accretion column, and maximum when the pole was facing away from the observer. Norton and Watson (1989) presented X-ray spin modulations from several IPs. They successfully modelled the data with a combination of photoelectric absorption and self occultation, inclusion of partial covering to the models helped the fits. This resulted in a picture of structured accretion flow and emission region, with absorption and occultation effects arising in phase and at the same pole of the white dwarf. They also found, in contrast to King and Shaviv (1984), that the emission region occupies at least one quarter of the white dwarf surface.

33 second spin modulation in the light curves and the quiescent hardness ratios, have the same character as seen in other IPs in that AE Aqr is hardest during minimum flux. Though it has to be noted that this is can not due entirely to an increase in $N_H$ as no significant change has been detected in the absorption. EXOSAT observations of IPs have shown higher modulations than that observed in AE Aqr (Hellier, Mason & Garlick 1993). As $N_H$ is thought to be the cause of this modulation observations at lower X-ray energies (eg the PSPC) might be expected to produce a higher amplitude of modulation. The amplitude of this modulated count rate is observed to vary, increasing quite substantially during the flare as would be expected from an increase in accretion rate. However other changes in count rate do not yield such increases in modulation as might be expected from accretion driven production of X-rays though there is a weak correlation present. The amplitude of modulation has been shown to be independent of intensity in recent HST observations both before and during our observations (Eracleous et al.1994 and Eracleous et al.1995). The amplitude of the EUV modulation was found to be $\approx 40\%$, much greater than the X-ray modulation.

Recent modelling by Wynn, King & Horne (1995) supports the low accretion rate measurement. They modelled the accretion as a set of diamagnetic blobs following a ballistic trajectory which are modified by the magnetic force density

$$ f_{\text{mag}} = -k v_r, $$

where $v_r$ is the relative velocity between the blob and the field, and $k \sim 1/t_{\text{drag}}$ is the
Figure 4.19: A view of the AE Aqr system from above showing the simulated accretion flow. The Roche lobes of the primary and secondary are shown, with the orbital motion being in the clockwise direction. The arrows indicate inertial frame velocity vectors. (Figure courtesy of G. Wynn).

drag coefficient. They fixed the orbital and spin periods, and the mass of the white dwarf and the mass ratio, $q$. Fig 4.19 shows the result of a typical simulation, it can be seen that most of the mass is being ejected from the system via a gas stream in the plane of the binary. This ejection in mass results from the rapid rotation of the white dwarf, which causes a large relative velocity between the blobs and the magnetic field. At closest approach the blobs gain orbital energy and angular momentum thus spinning down the white dwarf and then cruising out of the system. Doppler maps of AE Aqr have allowed the value of $k$ and other parameters to be determined. The predicted magnetic moment, $\mu$ is consistent with the previous estimate $\mu \sim 2 \times 10^{52}$ G cm$^3$ (Warner & Wickramasinghe 1991), the $L_1$ overflow rate is in the range $6 \times 10^{13} - 3 \times 10^{14}$ g s$^{-1}$ and the magnetic dissipation rate $10^{33} < L_{\text{mag}} < 10^{34}$ erg s$^{-1}$ which is consistent with the spin down power of the white dwarf of $6 \times 10^{52} L_{\odot}$ erg s$^{-1}$ (de Jager et al. 1994). In this scenario only $\sim 1\%$ of the material from the secondary is actually accreted.
4.4.3 Conclusions

In conclusion AE Aqr has shown itself to be a unique CV. The ROSAT observation has raised more questions than it has answered. I have determined its spectrum, the low temperature being consistent with its low accretion rate. The modulation does not behave in the way we would expect is it does not increase with increasing count rate, as would be expected if accretion were the source of the system's luminosity.

I cannot rule out the possibility that the secondary is contributing some of the X-ray flux. This may account for the lower modulation level than that seen in other IPs (Hellier, Mason & Garlick 1993). Also coronal emission reduces in the UV and so might explain the higher modulation level seen in the UV as relatively more of the emission is from accretion rather than a coronal source, though this does not explain the lack of variation in the modulation amplitude in the UV modulation.

Though I can just explain the X-ray emission from AE Aqr being due to coronal emission from the secondary, with a smaller accretion component this is unlikely to be the case. The UV flares have been shown not to be due to coronal activity. The X-ray flare is unlikely to be coronal as the temperature profile between quiescence and flare does not vary in the same way as coronal stellar flares. Also this active secondary behaviour, e.g. flares and spectrum, has not been observed for other CVs (Eracleous, Halpern & Patterson 1991).

The model of Wynn, King & Horne (1995) in which the majority of the material that leaves the secondary is flung out of the system entirely, though in its early stages of development, currently looks very promising, explaining many of the observed features such as the doppler maps, spindown power and the lack of much accretion.

The combined data sets show that AE Aqr spends approximately 50% of the time flaring. HST observations of flares reveals that these do not originate at the white dwarf due to the lack of change in the modulated flux and therefore AE Aqr spends most of the time not accreting or accreting very weakly. Measurements of line velocities show that the flares cannot be occurring on the secondary star, the accretion stream is thus the only remaining possible location for the flares. Similar conclusions were arrived at by Horne & Eracleous

100
This clearly lends support to the model of Wynn, King & Horne (1995).
Chapter 5

Soft X-ray and EUV Properties of the Polar QS Tel.

5.1 Introduction

QS Tel was discovered as one of the brightest sources in the ROSAT Wide Field Camera (WFC) all sky survey conducted between June 1990 and January 1991 (Buckley et al. 1993). It was independently discovered in the PSPC survey, and is reported in Beuermann & Thomas (1993). The WFC observation and subsequent optical photometry and spectroscopy showed it to be a polar with an orbital period of 2.33 hr, placing it well inside the cataclysmic variable 'period gap'. Initial results indicate that the system exhibits a soft X-ray excess of at least a factor $\sim 2$ over the hard bremsstrahlung component. This was supported by a GINGA observation taken in October 1991 which did not detect QS Tel, thus placing a limit on any hard X-ray flux of $\sim 2 \times 10^{-12}$ erg cm$^{-2}$ s$^{-1}$ in the 2-10 keV band. An orbital period fold revealed the presence of bright and faint phases caused by the accreting pole rotating in and out of view during the orbital cycle.

Optical spectropolarimetry during a low state revealed a magnetic field of 57 MG for the accreting pole (Ferrario et al. 1994). However a later bright phase optical observation (Schwope et al. 1995) found cyclotron emission lines of two magnetic poles of field strengths
~ 47MG and 70-80 MG, (the later being the largest field measured for polar systems so far) indicating the presence of a non-dipolar field. Schwopa et al. (1995) also took low-state spectra which showed the secondary to be an M dwarf, from which they inferred a distance to QS Tel of 150-190 pc.

A performance verification phase EUVE observation in July 1992 (Warren et al. 1993) found QS Tel in a deep low state where the EUVE light curve was dominated by two flare events. No orbital variation was found and thus they concluded that QS Tel had gone into a state of little/no accretion.

This chapter will describe pointed ROSAT PSPC and WFC observations taken in October 1992 and also simultaneous optical photometry. Comparison with EUVE data taken in Aug and Oct 1993 will also be made.

5.2 Observations

ROSAT (see appendix A) made a pointed observation of QS Tel in October 1992. QS Tel was observed in the PSPC 40 arcmin off-axis so as to limit the effects of the space craft wobble. The observation consisted of 19 separate slots with a total exposure of 22.6 ksec. The WFC also observed QS Tel using principally the s2b filter (61-111 eV) in 16 slots for 24.5 ksec. V and I band photometry were also made during this time. An observation log is detailed in Table 5.1.

The PSPC data were extracted using a source circle of 5 arcmin and a surrounding background annulus using the ASTERIX package Saxton (1991). The mean count rate in the PSPC was 1.53 ± 0.03 cts s⁻¹. The WFC data were extracted using a source circle of radius 3 arcmin and a background region nearby by of the same radius. The subtracted data were corrected for exposure, PSF, vignetting and the efficiency drop (0.187%). The mean count rate in the WFC (s2b filter) was 0.113 ± 0.011 cts s⁻¹.

V and I band photometry observations were made by D. Buckley using the 1m telescope at the South African Astronomical Observatory (SAAO) on 13 and 17 October, i.e. the first
observation was simultaneous with ROSAT and the second was 1.5 days after the end of the ROSAT observation (see table 5.1 for more detail). The mean V and I band magnitudes for the two datasets were \( V = 16.4 \pm 1.7, 16.2 \pm 1.3 \) and \( I = 15.6 \pm 1.7, 15.3 \pm 1.3 \) thus placing QS Tel in a fainter state than the observations made in August and October 1991 (Buckley et al. 1993) with drops in the V and I band magnitudes of \( \Delta V = 0.6 \) and \( \Delta I = 0.7 \).

QS Tel was the subject of two EUVE observations in Aug and Oct 1993, with exposure times of 29ks and 69ks respectively (Rosen et al. 1995). The EUVE satellite spectrometer samples an energy range 70-760\(\text{Å} \) which is subdivided by 3 gratings, the SW (70-190\(\text{Å} \)), MW (140-380\(\text{Å} \)) and the LW (280-760\(\text{Å} \)). QS Tel was only significantly detected in the SW filter at count rates of \( 0.197 \pm 0.004 \) and \( 0.224 \pm 0.002 \) cts s\(^{-1} \) for the two observations. The source is also observed simultaneously by the deep survey instrument (DS) which provides photometric information over the range 70\(\text{Å}-360\(\text{Å} \), the raw count rates were \( 2.020 \pm 0.004 \) and \( 1.497 \pm 0.005 \) cts s\(^{-1} \) for the August and October observations respectively. Applying corrections to account for Primsching/dead-time (\( \sim 1.09 \)) and the dead time spot effects (\( \sim 2.3 \)) yields absolute DS instrument count rates of \( 4.26 \pm 0.66 \) and \( 3.92 \pm 0.71 \) cts s\(^{-1} \) respectively.
### Table 5.1: Observation Log for QS TEL

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* Indicates a slot which is actually two slots separated by less than 10 seconds.
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* indicates a slot which is actually two slots separated by less than 10 seconds.

5.3 Results

5.3.1 Timing Analysis

PSPC

The light curve was extracted using 30 sec time bins and then was folded into 128 phase bins over the 2.33hr orbital period using the linear ephemeris

\[ HJD = 244 8894.5568(15) + 0.09718707(16) E \]

(Schwope et al. 1995) (see fig 5.1)¹, this ephemeris will be used in all of the following folds. The folded light curve is highly variable with both bright and faint phases. Fig 5.1 also shows the survey folded light curve again revealing the bright and faint phases, though here the bright phase count rate is much greater than the pointed PSPC observation (ie bright phase count rate ratio of 5:1 for the survey to October 1992 observations). The faint phase flux has not changed between the two observations (count rate ~ 1cts s⁻¹). Each

²Schwope et al. (1995) found marginal evidence of a quadratic term in the ephemeris, however for this work I shall not be using it as there are still some uncertainties in this
phase is covered at most twice, though in some cases only once. Examining individual time section folds reveals that the light curve is highly variable on a cycle to cycle basis (see fig 5.2).

The bright phase also shows a narrow dip centered at phase 1.1 with a width of $\Delta \phi = 0.025$. This dip was observed in ingress at the end of slot 3 and in egress at the beginning of slot 17, these slots were separated by 22 binary cycles. The minimum of the dip is not observed. Although the dip is not observed completely, it is unlikely to be an artifact of the instrument for two reasons, firstly the ingress and egress are observed in different binary cycles. Also examination of the background flux also revealed no event large enough to cause the dip. Housekeeping records of the PSPC also showed no peculiar event at either of these to phases. No other slot end showed a large change in count rate.

Light curves for three energy bands were then extracted, Band 1: 0.1-0.2 keV, Band 2: 0.2-0.5 keV and Band 3: 0.5-2 keV. The first two divide the soft energy component and the latter covers the lower energies of the hard component. The hardness ratio (Band 2/Band 1) of the soft component is noisy with the possibility of some real spectral variability over $\phi = 0.55 - 0.65$ (see fig 5.3 for the two soft band light curves and the resulting hardness ratio). The hard component also shows a slight variation over the orbital cycle, increasing during the bright phase (see fig 5.4). The poor signal to noise of the data also means it is impossible to say whether the narrow dip is present in the hard component.

WFC

The WFC light curve was extracted and folded using 20 phase bins. This again showed the presence of the same bright and faint phases (see fig 5.5), with a similar reduction in bright phase flux as compared to the ROSAT WFC survey. A dip is observed at the same phase as the PSPC dip and was almost completely sampled with a data gap of only 7 secs during the dip minimum. Extracting the data at a higher time resolution and folding into 50 phase bins (a bin length of 168s) reveals the dip in greater detail and show it to be broader than the PSPC dip with a phase width of 0.08, though the egress occurs at approximately the same phase ($\phi = 1.12$, see fig 5.5).
Figure 5.1: Upper Panel: The orbital light curve folded into 128 bins using the ephemeris of Schwope et al. (1995) with phase zero corresponding to the blue/red crossing time of the optical narrow emission line. The bright and faint phases are clearly flickering and a deep dip is observed at phase 1.1. (N.B. all future folds will be folded on the same ephemeris.)

Lower Panel: This shows the orbital light curve of the ROSAT PSPC data, taken from Beuermann & Thomas (1993). A large difference between the bright phase count rates between the two observations is clearly evident.
Figure 5.2: This figure shows individual data slots folded over the orbital period. The dotted light curve is the mean orbital fold. A large degree of cycle to cycle variability is evident.
QS Tel: Band1, Band2 and Hardness Ratio Folds

Figure 5.3: Upper Panel: The band 1 (0.1-0.2 keV) orbital period fold. Middle Panel: The band 2 orbital period fold (0.2-0.5 keV). Lower Panel: The band 2/band 1 ratio. These show that there is some hardness ratio variation over the orbital cycle (particularly during \( \phi = 0.55 - 0.65 \)).
Figure 5.4: The band 3 (0.5-2 keV) orbital period folded in 50 bins. The mean count rate is $\sim 0.04$ cts s$^{-1}$ with a slight increase during the bright phase.
Figure 5.5: Upper panel: The mean orbital light curve folded into 20 phase bins. Bright and faint phases are observed with a deep dip at phase 1.1, coincident with the dip observed by the PSPC. Lower Panel: This shows one data section containing the dip plotted at a higher resolution.
A two bin hardness ratio of the PSPC data to the WFC data reveals no significant differences, with the mean bright and faint phase ratios being $13.1 \pm 1.5$ and $15.8 \pm 2.6$ respectively, thus for future analysis, when comparing the WFC and PSPC data the phase averaged spectrum will be used.

**Optical data**

The simultaneous V and I band photometry coincided with the PSPC slots 3 and 4, see fig 5.6. As the optical data covers ~ 1.5 orbital cycles the folded light curve has not been binned up see fig 5.7. This fold shows some orbital modulation which has a similar morphology to that observed in August 1991 (Buckley et al. 1993), the V band bright phase being roughly consistent with that seen by ROSAT between phases 0.85 and 1.4. There is a broad dip at phase 1.05-1.2 and depths ~ 0.1 and ~ 0.2 mags for the V and I bands respectively, which again is coincident with the ROSAT dip. A fainter phase is present between phases 0.5 and 0.85, being ~ 0.3 mags fainter in the V band and ~ 0.2 in the I band.

Even with this limited orbital coverage with only phases 1.0-1.4 being covered twice it is obvious that the optical light curve is highly variable.

The V and I band photometry taken 1.6 days later (for the light curve see fig 5.6) and can be seen folded in the same way in fig 5.7. Although this is only slightly later the light curves appear quite different. The V and I band fluxes are ~ 0.2 and ~ 0.3 mags brighter respectively. There are two cycles of coverage in both cases and some variability is seen, though not as great as that seen in the first case. The dip seen in the previous data at phase 1.05-1.2 is only present in the I band data.

**Comparison to EUVE Observation**

The light curves for the two SW observations and the deep survey instrument (for October 1993 only) were extracted and folded into 50 phase bins and can be seen in fig 5.8. The overall profile from both of the observation epochs is essentially similar. The morphology of
Figure 5.6: Light curves for each of the V and I bands (V band shifted upwards in each case). The upper panel indicates the data simultaneous with ROSAT. The lower panel was taken ~ 1.5 days later. The phase of the ROSAT dip is also indicated for comparison.
Figure 5.7: Upper Panel: The V and I band orbitally phased data (not binned up), for the data simultaneous with the PSPC. Lower Panel: The same plot except for data recorded \( \sim 1.6 \) days after the ROSAT observation. For better clarity the V band has been shifted up by 1.5 mags.
the underlying orbital modulation is double peaked with maxima phases 0.6 and 1.1 (these will be referred to as the secondary and primary maxima respectively). The secondary maximum is less pronounced in the August observation.

During the primary maximum ($\phi = 0.85 - 1.35$) there is a deep dip lasting 0.1 in phase and centered on $\phi = 1.05$. At the minimum of the dip there appears to be some evidence of flaring activity. Observations of the dip made by the Deep Survey instrument reveal how the dip profile changes (see fig 5.9). Firstly, the dip ingress occurs at a different phase in each of the three slots that cover it. Secondly, during dip minimum flaring activity is observed in three slots. Finally, the dip egress occurs at exactly the same phase in all the slots which cover this phase.

The double peaked morphology clearly contrasts with that of the earlier ROSAT observation, which only show one maximum, which occurs at the same phase as the EUVE primary maximum. The secondary maximum coincides with the ROSAT faint phase. As the WFC and EUVE instruments cover the same energy range the change in modulation is not due to an energy dependence.

The dip observed by EUVE occurs at the same phase as the dip observed in the PSPC and WFC light curves, though it is much broader in the EUVE than the PSPC ($\Delta\phi \approx 0.1$ and $0.025$ respectively). The egress of the EUVE dip is coincident with that of the PSPC and is consistent with the WFC egress.

5.3.2 ROSAT Spectral Analysis

Three spectral files were created for the data, a mean spectrum of the data and bright and faint phase spectra (phases 0.85-1.45 and 0.45-0.85 respectively). These all used the SASS recommended binning (see Appendix A) and included a systematic error of 2% to allow for uncertainty in the gain of the instrument. Fitting was performed using the XSPEC package (Shafer et al. 1991) and the most recent, appropriate response matrix (DRM 36). Initially fitting the mean spectrum with an absorbed blackbody gave a poor fit with a reduced $\chi^2$ of 4.0, inclusion of a bremsstrahlung component of temperature 10keV gave an acceptable fit with a reduced $\chi^2$ of 1.0, thus detecting for the first time the presence
Figure 5.8: Upper Panel: The EUVE SW (70 – 190Å) folded light curve (50 phase bins) recorded in August 1993. Middle Panel: The EUVE SW orbital light curve for October 1993. Lower Panel: The EUVE Deep survey instrument (70 – 360Å) folded light curve. These plots indicate that a second accretion region is now emitting between phases 0.4-0.8. The deep dip is again seen at the same phase as the dip observed by the PSPC (phase 1.05) though in this case it is much broader.
Figure 5.9: Individual observations of the dip phase made by the EUVE Deep survey instrument. The vertical dotted lines indicate the dip ingress and egress times. The presence of flares during dip minimum can be seen in observations 4, 6, 12 and 14. Ingress position also appears to vary, see slots 3, 4 and 12. The egress occurs at the same phase in all the observations.
Table 5.3: The mean, bright and faint phase spectral fits using an absorbed blackbody and Bremsstrahlung model.

<table>
<thead>
<tr>
<th>Model Parameter or Statistic</th>
<th>Average (errors are for $\Delta \chi^2 = 7.78$)</th>
<th>Bright Phase</th>
<th>Faint Phase</th>
</tr>
</thead>
<tbody>
<tr>
<td>$N_H (10^{19} \text{ atoms cm}^{-2})$</td>
<td>$0.00^{+4.00}_{-4.00}$</td>
<td>$0.00^{+3.90}_{-3.90}$</td>
<td>$0.00^{+23.9}_{-23.9}$</td>
</tr>
<tr>
<td>Blackbody $kT (\text{eV})$</td>
<td>$24.2^{+1.3}_{-1.4}$</td>
<td>$24.6^{+1.3}_{-1.4}$</td>
<td>$24.2^{+1.6}_{-1.4}$</td>
</tr>
<tr>
<td>$f_{\text{obs}} (0.1 - 2.5 \text{keV})$</td>
<td>$5.00^{+1.14}_{-0.51}$</td>
<td>$5.55^{+1.44}_{-0.77}$</td>
<td>$3.00^{+0.66}_{-1.02}$</td>
</tr>
<tr>
<td>Bremsstrahlung $kT (\text{keV})$</td>
<td>frozen at a value of 10</td>
<td></td>
<td></td>
</tr>
<tr>
<td>$f_{\text{obs}} (0.1 - 2.5 \text{keV})$</td>
<td>$0.50^{+0.06}_{-0.03}$</td>
<td>$0.51^{+0.08}_{-0.03}$</td>
<td>$0.56^{+0.12}_{-0.08}$</td>
</tr>
<tr>
<td>$\chi^2, dof$</td>
<td>26.9, 28</td>
<td>27.7, 22</td>
<td>12.4, 22</td>
</tr>
</tbody>
</table>

of a hard component in QS Tel. The fits for the mean and bright and faint phase spectra can be seen in table 5.3.

These results clearly support the lack of hardness ratio change with no discernible difference between the bright and faint phase spectra. As there is no difference between the bright and faint phases for further analysis I shall just deal with the mean spectrum.

The WFC S2b data point was extracted and then added to the spectrum to see if an improvement in the $N_H$ constraint can be achieved. The fit has a reduced $\chi^2$ of 1.2, which is greater than the PSPC result though still acceptable. The resulting spectrum can be seen in fig 5.10 and the fit results in table 5.4. The effect of adding in the WFC data was to allow a better constraint the $N_H$ and to reduce the blackbody temperature. This blackbody temperature is higher than that measured by EUVE (Rosen et al. 1995) which has a value of 14.7 eV. This is hard to explain as an instrument effect, although the PSPC and WFC combination has a lower resolution than the EUVE it does cover essentially the same energy range. However the large accretion mode changes may provide the explanation for this effect. From this spectrum the total X-ray luminosity ($L = \pi d^2 f$) at a distance of 150 pc (Schwope et al. 1995) is determined to be $2.9 \times 10^{31} \text{ erg s}^{-1}$. The
Figure 5.10: The detector space spectrum of the mean PSPC and WFC (s2b) spectrum, for the model given in table 5.3.

The total hard component flux is $1.5 \times 10^{-12}$ erg cm$^{-2}$ s$^{-1}$ with an estimated upper bound (assuming a temperature of 30keV) of $3.1 \times 10^{31}$ erg cm$^{-2}$ s$^{-1}$.

The narrow dip observed by all the instruments appears to be broader in the lower energy range of the WFC and EUVE. Modelling the the effects of $N_H$, determined an amount, required to reduce the WFC flux by 80%, of $6 \times 10^{19}$ atoms cm$^{-2}$. This amount of $N_H$ in the PSPC band would reduce this flux by 65%.

5.3.3 EUVE Spectral Analysis

A mean spectrum of the second EUVE observation (75-190Å) was accumulated (Rosen et al. 1995). To account for Fixed Pattern Noise (a modulation of the signal on the detector with a spatial scale of $\sim 17$ pixels, Vallerga et al. 1991) a 20% systematic error was added in
Table 5.4: Simultaneous spectral fit of the mean PSPC and WFC (02b) data using an absorbed blackbody and bremsstrahlung model.

<table>
<thead>
<tr>
<th>Model Parameter / Statistic</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>(errors are for $\Delta \chi^2$ of 7.78)</td>
<td></td>
</tr>
<tr>
<td>$N_H$ ($10^{19}$atoms cm$^{-2}$)</td>
<td>3.35$^{+1.04}_{-0.88}$</td>
</tr>
<tr>
<td>Blackbody kT (eV)</td>
<td>20.4$^{+1.4}_{-1.8}$</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2.5keV)$</td>
<td>4.39$^{+0.13}_{-0.12}$</td>
</tr>
<tr>
<td>($\times 10^{-12}$ergs cm$^{-2}$s$^{-1}$)</td>
<td></td>
</tr>
<tr>
<td>Bremsstrahlung kT (keV)</td>
<td>frozen at a value of 10</td>
</tr>
<tr>
<td>$f_{obs}(0.1 - 2.5keV)$</td>
<td>4.78$^{+0.06}_{-0.03}$</td>
</tr>
<tr>
<td>($\times 10^{-12}$ergs cm$^{-2}$s$^{-1}$)</td>
<td></td>
</tr>
<tr>
<td>$\chi^2$, dof</td>
<td>33.5, 29</td>
</tr>
</tbody>
</table>
quadrature to the data in each bin. Data were grouped to achieve a minimum of 100 counts per bin. Fitting with an absorbed blackbody model reveals a best of $kT = 14.8 \pm 0.6$ eV and $N_H = 4.4 \pm 0.03 \times 10^{16}$ atoms cm$^{-2}$ with a $\chi^2$ of 1.06 (for a $\nu$ of 139). This temperature is lower than that measured from the PSPC data. Fitting the ROSAT spectrum with the 15eV temperature from EUVE gives an unacceptable fit with $\chi^2 = 3.88$ (for a $\nu$ of 31).

5.4 Discussion

These observations have shown us many interesting features about the polar QS Tel, namely the difference in the orbital modulation between the ROSAT and the EUVE observations just a year later and the narrow dip observed at phase 1.1. The spectrum has been characterized as fairly typical for a polar, with a difference in the measured blackbody temperature between the PSPC and EUVE observations. I shall deal with each of the main points in turn.

Changes in the light curve morphology

The ROSAT observation of QS Tel shows a similar morphology to that observed during the ROSAT survey (Buckley et al. 1993, Beuermann & Thomas 1993) with bright and faint phases. However the relative strength of the bright phase has reduced since the survey observations. In contrast to this the EUVE observation shows a double peaked morphology. Table 5.5 summarises observations made of QS Tel. The table clearly shows the range count rates and states that QS Tel exhibits. The change from the bright-faint mode to two bright phase mode, with the comparable second maximum being approximately half a cycle from the primary maximum, can occur in just 10 months and similarly for the change from bright-faint to low state. The bright phase profiles vary in different observations, the lower count rate in the PSPC also means that it is difficult to differentiate between the bright and faint phases. Are these different light curve morphologies as a result of changes in the accretion geometry?

Accretion geometry changes have been observed in several polars, e.g. AM Her (Heise et
Table 5.5: This table details previous observations made of QS Tel. Count rates and states of the source are given for comparison.

<table>
<thead>
<tr>
<th>Date</th>
<th>Instrument</th>
<th>mag$^a$</th>
<th>X-ray/EUV$^b$ cts s$^{-1}$</th>
<th>State</th>
<th>1/2 pole</th>
<th>Ref</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td></td>
<td>bright phase counts</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Sept-Oct 1990</td>
<td>WFC</td>
<td>-</td>
<td>~ 0.65</td>
<td>high 1</td>
<td>1</td>
<td>1</td>
</tr>
<tr>
<td>Sept-Oct 1990</td>
<td>PSPC</td>
<td>-</td>
<td>15</td>
<td>high 1</td>
<td>2</td>
<td></td>
</tr>
<tr>
<td>Aug 1991</td>
<td>-</td>
<td>15.6 (B)</td>
<td>-</td>
<td>high</td>
<td>-</td>
<td>1</td>
</tr>
<tr>
<td>Oct 1991</td>
<td>-</td>
<td>15.6</td>
<td>-</td>
<td>high</td>
<td>-</td>
<td>1</td>
</tr>
<tr>
<td>July 1992</td>
<td>EUVE</td>
<td>17.2</td>
<td>0.0062</td>
<td>low</td>
<td>-</td>
<td>3</td>
</tr>
<tr>
<td>Aug 1992</td>
<td>-</td>
<td>17.4</td>
<td>-</td>
<td>low</td>
<td>-</td>
<td>4</td>
</tr>
<tr>
<td>Sept 1992</td>
<td>-</td>
<td>16.2</td>
<td>-</td>
<td>int</td>
<td>-</td>
<td>5</td>
</tr>
<tr>
<td>Oct 1992</td>
<td>PSPC</td>
<td>16.3</td>
<td>3</td>
<td>int</td>
<td>1</td>
<td>6</td>
</tr>
<tr>
<td></td>
<td>WFC</td>
<td>16.3</td>
<td>0.14</td>
<td>int</td>
<td>1</td>
<td>6</td>
</tr>
<tr>
<td>Aug 1993</td>
<td>IUE</td>
<td>-</td>
<td>-</td>
<td>high</td>
<td>-</td>
<td>7</td>
</tr>
<tr>
<td>Aug 1993</td>
<td>EUVE</td>
<td>15.1 (B)</td>
<td>0.28</td>
<td>high</td>
<td>2</td>
<td>5, 6</td>
</tr>
<tr>
<td>Sept 1993</td>
<td>EUVE</td>
<td>-</td>
<td>0.3</td>
<td>high</td>
<td>2</td>
<td>6</td>
</tr>
<tr>
<td>June 1994</td>
<td>IUE</td>
<td>-</td>
<td>high</td>
<td>-</td>
<td>7</td>
<td></td>
</tr>
</tbody>
</table>

$^a$ V band magnitude unless otherwise stated
$^b$ s2b count rate unless otherwise stated

References:
1 Buckley et al. 1993
2 Beuermann & Thomas 1993
3 Warren et al. 1993
4 Ferrario et al. 1994
5 Schwope et al. 1995
6 Rosen et al. 1995
7 de Martino et al. 1995
al. 1985), QQ Vul (Osborne et al. 1987), BY Cam (Ishida et al. 1991 and references therein),
RXJ1940-1025 (Watson et al. 1995) and BL Hyi (Beuermann et al. 1985, Beuermann &
Schwope 1989). The causes of the changes in these polars have been attributed to two
different mechanisms. In the case of AM Her and QQ Vul these changes are difficult to
explain, the 'reverse mode' light curve is best attributable to the accretion switching from
one pole to another. This pole switching can be caused by greater levels of accretion which
mean that the accreting material can penetrate deeper into the magnetosphere and thus
allow accretion at the second pole. In the case of BY Cam and RXJ19490-1025 these
accretion mode changes have been attributed to a slight asynchronism between the white
dwarf spin and the orbital motion. The EUVE observations presented in this chapter
provide a good example of accretion occurring at both poles simultaneously, the EUVE
light curves simply showing the resumption of mass transfer onto the second pole which
was inactive in the ROSAT observations. The enhancement of the secondary maximum in
the later EUVE observation may indicate the ongoing change in the geometry. Schwope
et al. (1995) found cyclotron emission from a second pole in QS Tel just 17 days before
the ROSAT observations presented here.

Whether accretion rate changes could cause the observed changes in QS Tel depends on
several factors relating to the penetration of the magnetosphere. If the infalling material
is largely homogeneous an increase in its density and therefore its ram pressure leads to
compression of the effective magnetospheric radius. For a stream composed of blobs, be­
haviour is related to the viscous drag time scale and thus the blob density (King 1993,
Wynn & King 1995). Penetration deeper into the magnetosphere results from an increase
in accretion rate which in turn causes the density of blobs to increase (a possible conse­
quence of increased mass transfer). The blobs can then persist for longer and pass further
into the magnetosphere traversing a larger range of azimuth before threading onto field
lines. Thus some material can thread onto field lines flowing to the second pole.

To examine the relative changes in the accretion rate between the ROSAT and EUVE
observations, the spectra were extrapolated from the soft X-ray to the EUV. This was
done by assuming the same temperature and column (set at that measured for the PSPC)
and then scaling this to the higher count rate observed in the PSPC during the survey. This
underpredicts the EUVE SW count rate by a factor of ~ 2, allowing for the accretion at two
poles during the EUVE observation. Alternatively comparing the total soft component from each of the observations, again appropriately scaled, shows that this brightening factor is as much as 5. Thus the accretion rate during the EUVE observation is 2-5 times higher than that during the ROSAT survey. Such an increase in accretion rate, for the homogeneous accretion scenario, results in a contraction of the magnetospheric radius. Using equation 1.18, assuming that threading occurs in this region, a simple estimate of the contraction of the magnetospheric radius can be gained. In this case an increase in accretion rate by factors of between 2-5 result in a contraction of the magnetospheric radius of just 18-37%. It seems unlikely that a route to the second pole could result from this level of constriction for any reasonable dipole configuration, even if not centered on the white dwarf.

Rosen et al. (1995) performed simple simulations to test the effects of enhanced accretion in the case of a structured stream. These were based on the treatment of such streams by King (1993) and Wynn & King (1995). Although the geometrical and dynamical properties of QS Tel are poorly known, they used a field strength of a few tens of MG, a dipole colatitude and longitude of 60° and 10° respectively. They find that it is possible for blobs to accrete onto the second pole if the accretion rate increases by a factor of ~ 5. Thus blob penetration of the magnetosphere may offer an explanation of the geometry changes seen in QS Tel. This has some implications. If, at high accretion rates, the blobs are dense enough to penetrate to the second pole then it follows that less dense blobs and tenuous interblob material is likely to accrete mainly at the primary pole facing the companion star. Thus the primary pole would be expected to be a stronger emitter of hard X-rays, from material shocking at the surface or low optical depths of the white dwarf, than the secondary. Beardmore et al. (1995) found such a phase shift in the hard and soft X-ray light curves in observations of QQ Vul. However the lack of medium X-ray observations during the EUVE observation means that this cannot be tested.

In the case of asynchronous systems the white dwarf is spinning slightly asynchronously where the difference between the orbital and spin periods is small (< 0.1P_{orb}). For these systems accretion geometry changes would be expected at the beat period (P_{beat}^{-1} - P_{orb}^{-1}) i.e. changing in tens of days. However X-ray and EUV measurements are not at present numerous enough to prove this to be the case for either BY Cam or RXJ1940-1025.
Optical measurements of RXJ1940-1025 have shown a systematic behaviour at the beat period between the spin and orbit periods resulting in a 'supercycle' in the photometry (Patterson et al. 1995). At this time the there is still not enough data collected on QS Tel to determine whether asynchronism could be causing the accretion mode changes, should this prove to be the case it would be expected that observation of accretion only onto the second pole would be observed.

One interesting feature from the optical spectroscopy recorded simultaneously with the EUVE observations (Rosen et al. 1995) have found no emission line components associated with a stream flowing towards the second pole. This implies that this second stream is not optically bright despite it providing enough accreting matter to produce a pole as bright as the primary in the EUV. Given the similar characteristics between the two poles in the EUVE data, it is unclear why this should be the case.

The Narrow Dip

The presence of the narrow dip at phase 1.1 is reminiscent of other similar dips observed in systems such as EK UMa (Clayton & Osborne 1994, and Chapter 2), UZ For (Osborne et al. 1988), QQ Vul, AN UMa and V834 Cen (Mason 1985). These have been attributed to occultation of the emission region by the accretion stream. Another mechanism for producing dips is: eclipse of the emission site by the secondary.

Eclipse by the secondary star would require a mass ratio, q (M_2/M_1) > 1, to produce an eclipse lasting ~ 0.1 in phase. Also the soft component is thought to be produced from a small region of the white dwarf surface, so the ingress and egress timescales should take much less than 40s. The ingress of the EUVE data takes ~ 300 s. The egress is short in the EUVE observation (30-40s). To get a large ingress timescale would require shadowing of an EUV emitting structure which is extended by ~ 10 white dwarf radii. It is unlikely that more than a few percent of the EUVE flux would be emitted so far from the white dwarf. Also the remaining flux in the EUV would have to eminate from the secondary, this luminosity is exceeds that expected for secondary stars. Thus eclipse by the secondary is unlikely to be the cause of the dip.
It seems reasonable at this stage to conclude that occultation by the accretion stream of the emission region is the cause of the dip observed in QS Tel. The dip has been observed on three occasions in different wavelengths and thus provides more information than the dip present in EK UMa and is therefore an interesting case study. The dip is observed at the same phase in all the observations, so it obviously appears to be unaffected by the accretion mode (except in cases of cessation of accretion). The breadth of the dip is observed to be greater in the EUV than the soft X-ray. The dip is even observed slightly in the optical though the mechanism may be different in this case.

There is no energy dependence observed during the dip in the EUVE data. This rules out photoelectric absorption as the absorbing mechanism. Obscuration by dense blobs, which are opaque in the EUV and soft X-ray, in a tenuous medium would not produce any energy dependence. Thus could account for this lack of observed energy evolution through the dip. It could also explain the presence of flares during dip minimum which might be associated with times, when due to statistical chance, the source is not entirely obscured by blobs. The observed soft X-ray excess in this system also supports this stream picture (Kuijpers & Pringle 1982; Frank, King & Lasota 1988). At longer wavelengths (e.g. UV) the blobs could become optically thin to bound-free and free-free absorption. By consideration of relative cross sections of the different absorption mechanisms for a dip to be present at longer wavelengths (see Chapter 6) the blobs require a very large optical depth in the EUV.

The difference in the breadths of the dips between the EUV and soft X-ray is hard to explain. To reduce the WFC flux by 86% (using photoelectric absorption) would require an $N_H$ of $6 \times 10^{19}$ atoms cm$^{-2}$, this value of $N_H$ in the soft X-ray band would produce a corresponding dip of 65%. From this it would be expected that the dip observed in the soft X-ray would be larger than observed (providing photoelectric absorption was the only mechanism for producing the EUV dip). It is true that the observation of the dip is not simultaneous between the EUV and soft X-ray, and this might explain the discrepancy. A flare occurring in the soft X-ray at the phase of the EUV ingress may be one explanation, this however would require a huge flare to produce a flux level observed over and above the 65% drop in flux, however during another observation slot (see fig 5.2) one such flare is observed. The EUVE dip observations do show us that the dip length can vary (refer
back to fig 5.9) therefore changes in the dip breadth are possible. Flares are also observed
during the dip minimum, but these are quite small and it is hard to believe that a similar
event in the X-ray could produce the observed profile. As this phase in the soft X-ray is
only covered once and the signal to noise of the WFC data is poor no real conclusions as
to the cause of the shorter dip can yet be reached.

The spectrum

The combined PSPC and WFC spectrum results in a blackbody kT of ~ 20 eV and a
corresponding \( N_H \approx 3.3 \times 10^{13} \) atoms cm\(^{-2}\). This is typical of many other polars (Cropper 1990). The observation also for the first time significantly detects a hard component
with a flux of \( \sim 5 \times 10^{-13} \) ergs cm\(^{-2}\)s\(^{-1}\) (0.1-2.5 keV) which corresponds to a flux of
\( \sim 7 \times 10^{-13} \) erg cm\(^{-2}\)s\(^{-1}\) in the 2-10 keV band. This result is entirely consistent with the
upper limit on this flux obtained from GINGA (\( \lesssim 2 \times 10^{-12} \) erg cm\(^{-2}\)s\(^{-1}\) in the 2-10 keV
band; Buckley et al. 1993).

No spectral difference is observed between the bright and faint phases, this is unexpected
if the sources of the bright and faint phases are supposed to arise from different regions.
The faint phase is either a weaker second pole or intrinsic emission from the white dwarf
(this is more likely as the faint phase count rate has not changed between the this and
the survey observations). These would be expected to produce different emission, the low
count rate during the faint phase may account for this, as poor statistics may not allow
us to determine the spectral nature well enough to distinguish it from the bright phase.

Ramsay et al. (1994) calculated a soft X-ray excess for QS Tel (based on the pointed
ROSAT data presented here) of 2.2 (with an upper limit of 10.6. Recent studies of the
relationship between the soft X-ray excess and magnetic field has found that for increasing
magnetic field the soft X-ray excess increases (Ramsay et al. 1994). However, for the high
magnetic field of QS Tel, one would expect a soft X-ray excess of > 50 (see fig 5. Ramsay
et al. 1994).

The EUVE spectrum reported by Rosen et al. (1995) determined a blackbody temperature
of 14.7 eV and an \( N_H \) of \( 4.29 \times 10^{19} \) atoms cm\(^{-2}\). The inferred bolometric luminosity of
each the two poles is $1.5 \times 10^{33} \text{ erg s}^{-1} \text{ cm}^{-2}$ and $2.1 \times 10^{33} \text{ erg s}^{-1} \text{ cm}^{-2}$ for the primary and secondary poles respectively (assuming a distance of 190 pc, Schwope et al. 1995). The blackbody temperature is lower than that measured by ROSAT a year earlier. Changes in accretion rate, which is thought to cause the change in accretion mode, may account for this difference.

Geometry

A geometrical study similar to that made for EK UMa in chapter 2 can be performed for QS Tel. QS Tel contains the same features which can be used to determine limits, namely:

1. Bright phase length, can be related to $i$ and $\delta$ by equation 1. However for QS Tel things are more uncertain as there is difficulty in determining the bright phase length. For the purpose of this study I shall use a value of $\Delta \phi = 0.55$ with a range of 0.5-0.6 also shown.

2. The lack of an eclipse by the Roche lobe filling secondary places an upper limit of the inclination of 73.5° (see section 2.4 for more details about the calculation).

3. The presence of an absorption dip suggests that the inclination is greater than the magnetic colatitude.

Fig 5.11 shows these limits and the region devoid of hatching shows the allowed region for the inclination of geometry. For a longer bright phase length the larger the allowed region. The allowed inclination and colatitude for a bright phase length of 0.55 is $i \sim 67 - 73.5°$ and $\delta \sim 62 - 72°$ for a bright phase length of 0.6 $i \sim 60 - 73.5°$ and $\delta \sim 42 - 72°$, for a bright phase length of 0.5 no region is allowed. An extended soft X-ray emission region produces a longer bright phase length. Thus the bright phase relation can only provide a lower limit to the inclination and magnetic colatitude.

Schwope et al. (1995) determined inclination limits as a function of the mass ratio from the measured radial velocity of the heated atmosphere of the secondary star. They found for a nominal white dwarf mass of 0.6$M_\odot$ the nominal value of the inclination is 35° and cannot
Figure 5.11: The geometrical constraints for QS Tel. The hatched regions are areas that are not allowed. The 'dot-dash-dot' lines indicate bright phase length limits, the upper corresponding to $\Delta \phi = 0.5$ and the lower $\Delta \phi = 0.6$. 
be greater than 50°. This is clearly in contrast with my findings where the lower limit for the inclination is ~ 60°. The parameters that determine my lower limit are just the presence of the absorption dip and the bright phase length which are both independent of orbital parameters. To achieve an inclination of ~ 50° requires a bright phase length of greater than 0.75 and, for $i = 35°$, the emitting pole must be virtually constantly in view which is clearly not allowed by the data. From the analysis of Schwope et al. (1995) an inclination of 60° requires a white dwarf mass in the range $M_2 = 0.3 - 0.55 M_\odot$ (3σ) with the optimum value at 0.44 $M_\odot$.

Hameury, King & Lasota (1988) determined that, accounting for uncertainties in models of the secondary, and specifically uncertainties in the entropy and opacity structure, an upper limit for the periods of systems below the period gap was 131 mins for a mass, of 1.44 $M_\odot$, if the system evolved from a period of > 4 hours. The low mass determination for QS Tel conclusively shows that QS Tel could not have evolved into the period gap as this would require a mass much greater than 0.6 $M_\odot$ and hence QS Tel must have been born inside the period gap.

Conclusions

These observations have shown two accretion geometries in QS Tel, i.e. one pole and two pole modes. They also have the first observation of an accretion stream dip for this system in both X-ray and EUV wavebands. The X-ray and EUV spectra have been characterised as fairly typical for a polar. Tighter geometrical constraints have also been determined.

Other systems show accretion geometry changes but these tend to be infrequent in these systems, in comparison QS Tel exhibits probably the most variable accretion geometry of all the polars thus making it a highly interesting subject of study. As a bonus accretion stream dips can also be studied. The spectrum is typical of polars, however more EUV/soft X-ray measurements will allow studies of temperature variation as a function of accretion mode to be made for the first time.
Chapter 6

A Hubble Space Telescope Observation of QS Tel.

6.1 Introduction

QS Tel has already revealed itself to be a highly interesting polar (see previous chapter). Recently de Martino et al. (1995) reported on some IUE observations of this object. They found strong variability in the major emission lines C IV (1550Å), He II (1640Å) and N V (1240Å) at the orbital period. Line velocities and phases were measured and found to be consistent with those observed in the optical. They propose that the line emission originates in the magnetically confined accretion stream as in many other polars. Indication that the far UV flux increased at superior conjunction of the white dwarf lead them to suspect a contribution from the heated accretion region.

This chapter reports on a HST Faint Object Spectrograph observation of QS Tel performed during June 1994. The wavelength coverage and resolution is very similar to the IUE instrument, however increased signal to noise and time resolution allow a more detailed analysis of the UV emission.
Table 6.1: Observation log of the HST FOS observation of QS Tel.

<table>
<thead>
<tr>
<th>Slot</th>
<th>No. of spectral file</th>
<th>Date</th>
<th>Start time (UT)</th>
<th>Duration (s)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>1-17</td>
<td>27/6/94</td>
<td>20^11m</td>
<td>327</td>
</tr>
<tr>
<td>2</td>
<td>18-95</td>
<td>27/6/94</td>
<td>21 04</td>
<td>1501</td>
</tr>
<tr>
<td>3</td>
<td>96-152</td>
<td>27/6/94</td>
<td>21 34</td>
<td>1096</td>
</tr>
<tr>
<td>4</td>
<td>153-287</td>
<td>27/6/94</td>
<td>22 41</td>
<td>2850</td>
</tr>
<tr>
<td>5</td>
<td>288-381</td>
<td>28/6/94</td>
<td>00 17</td>
<td>1808</td>
</tr>
<tr>
<td>6</td>
<td>382-423</td>
<td>28/6/94</td>
<td>00 52</td>
<td>808</td>
</tr>
<tr>
<td>7</td>
<td>424-475</td>
<td>28/6/94</td>
<td>01 54</td>
<td>1000</td>
</tr>
</tbody>
</table>

6.2 Observation

The HST observation of QS Tel took place on June 27th 1994 (after the COSTAR servicing mission) and lasted ~6hrs, with an exposure of ~2.6 hrs. As the HST is a low Earth orbiting satellite the observation was divided into 7 slots, a log of the observation can be seen in table 6.1. The data were recorded using the blue digicon on the Faint Object Spectrograph (FOS) (Kinney 1994), operated in rapid mode to obtain repetitive 19s exposures with minimal dead time (<1s) between integrations. The source was observed through the 0.9" aperture and the light dispersed through the low dispersion G160L grating to give spectra in the 1154-2508Å range with a resolution of 6.8Å. The HST spectra supplied by STScI were pre-calibrated in both flux and wavelength, with the wavelength scale accurate to ~0.3Å (~60 km s⁻¹ at 1550Å).

6.3 The mean spectrum

A mean spectrum can be seen in fig 6.1. The emission lines and continuum regions were identified for the purposes of the further analysis. The four emission lines that I will be studying in further detail are C IV, He II, Si IV and N V.
Figure 6.1: The mean HST FOS spectrum of QS Tel taken in June 1994 at 6Å resolution. The principal emission lines are identified. The feature labeled c is a blend of O I, Si II and Si III.
Table 6.2: Line luminosities ($L_i = 4\pi d^2 f$) calculated for $d=170$ pc (Schwope et al. 1995) for the four principal emission lines.

<table>
<thead>
<tr>
<th>Line</th>
<th>FWHM*</th>
<th>Central wavelength</th>
<th>$\pm 3\sigma$ range</th>
<th>Luminosity de Martino (1995)</th>
<th>Luminosity</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>(Å)</td>
<td>(Å)</td>
<td>(Å)</td>
<td>(erg s$^{-1}$)</td>
<td>(erg s$^{-1}$)</td>
</tr>
<tr>
<td>CIV</td>
<td>$\sim 11.2$</td>
<td>1549$^b$</td>
<td>1529-1565</td>
<td>$6.7 \times 10^{30}$</td>
<td>$6.4 \times 10^{30}$</td>
</tr>
<tr>
<td>HeII</td>
<td>$\sim 11.2$</td>
<td>1640</td>
<td>1623-1656</td>
<td>$2.3 \times 10^{30}$</td>
<td>$1.6 \times 10^{30}$</td>
</tr>
<tr>
<td>SiIV</td>
<td>$\sim 15.7$</td>
<td>1397$^b$</td>
<td>1373-1421</td>
<td>$2.2 \times 10^{30}$</td>
<td>$1.6 \times 10^{30}$</td>
</tr>
<tr>
<td>NV</td>
<td>$\sim 11.2$</td>
<td>1240</td>
<td>1223-1259</td>
<td>$1.8 \times 10^{30}$</td>
<td>$1.0 \times 10^{30}$</td>
</tr>
</tbody>
</table>

* FWHM is the intrinsic value of the line, decoupled from the instrumental resolution.

The continuum flux ($f$), calculated over the entire (1154-2508 Å) band with interpolation over the emission line regions, is $1.7 \pm 0.1 \times 10^{-14}$, which assuming a distance of $d=170$ pc (Schwope et al. 1995), leads to a luminosity of $L_c = 4\pi d^2 f = 5.9 \pm 0.4 \times 10^{31}$ erg s$^{-1}$. This is consistent with the IUE measurement of $5.7 \times 10^{31}$ erg s$^{-1}$.

The fluxes have been calculated for each of the lines by fitting the line with a gaussian and summing up the flux within $\pm 3\sigma$ and then subtracting the continuum contribution. The luminosity ($L = 4\pi d^2 f$) is calculated for the distance of 170 pc so as to make a direct comparison to values obtained by de Martino et al. (1995). Table 6.2 details these results.

The resonance line ratios of NV/CIV and NV/SiIV are consistent with those of other polars (see fig 6.2, Bonnet-Bidaud & Mouchet 1987). Fig 6.2 (adapted from Bonnet-Bidaud & Mouchet 1987) also shows the predicted line ratios (as calculated by Kallman 1983) from collisional excitation, with ion abundances resulting from photoionization by blackbody spectra. From this QS Tel is illuminated by a blackbody source of temperature in the range 15-25eV, which is consistent with earlier ROSAT/EUVE observations (Schwope et al. 1995, Rosen et al. 1995 & chapter 5).
Figure 6.2: This figure shows the measured line ratios of a sample of polar systems taken from Bonnet-Bidaud & Mouchet (1988), the line ratios measured for QS Tel are also shown. The lines indicate predicted ratios from collisionally excitation with ion abundances resulting from photoionisation by blackbody spectra. The solid, dotted and dashed lines indicate blackbody source temperatures of 15, 25 and 45 eV respectively.
Table 6.3: Wavelength Ranges used for determining the continuum.

<table>
<thead>
<tr>
<th>Wavelength Range (Å)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1260-1270</td>
</tr>
<tr>
<td>1430-1470</td>
</tr>
<tr>
<td>1590-1610</td>
</tr>
<tr>
<td>1700-1820</td>
</tr>
<tr>
<td>1900-2280</td>
</tr>
<tr>
<td>2310-2500</td>
</tr>
</tbody>
</table>

6.4 The photometric behaviour

6.4.1 Continuum

A light curve of the continuum was extracted using the summed flux from 6 wavelength regions between the prominent emission lines, see table 6.3. The light curve was heliocentrically corrected and then folded into 200 phase via the linear ephemeris determined by Schwope et al. (1995) (see chapter 5). This ephemeris will be used in all future orbital analysis. The result of this fold can be seen in fig 6.3, phase coverage is low, with most phases being covered just once, from five separate orbital cycles. Clear orbital modulation is evident with the light curve having a broad minimum at phase ~ 0.7 – 0.8. The fractional amplitude of the modulation is ~ 30%. The light curve is not smooth, 10% flickering/flaring is present with a timescale ~ 10 mins. The orbital modulation profile is different to that observed by ROSAT, the minimum being between phases 0.7-0.8 as compared to 0.3-0.8 in the ROSAT band. The bright phase exhibits a slower rise and decline compared to the EUV and soft X-ray bands. At phase 1.00-1.05 a dip (with a fractional depth of ~ 24%) is present, similar to the dip seen at other wavelengths (see chapter 5).

The continuum was divided into 4 different bands, which do not contain emission lines, and light curves were extracted to see whether the modulation in QS Tel was wavelength
Figure 6.3: The orbital period fold of the continuum emission (using the ranges shown in table 6.3), using the ephemeris of Schwone et al. (1995). Modulation is clearly present, with a fractional modulation of $\sim 30\%$ of the mean level. A dip is also seen at phase 1.0-1.06.
Table 6.4: This table shows wavelength ranges used for the extraction of four continuum bands, the percentage modulation in each band and the percentage depth of the dip.

<table>
<thead>
<tr>
<th>Band No.</th>
<th>Wavelength Range (Å)</th>
<th>Fractional Modulation (%)</th>
<th>Depth of Dip (%)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>1900-2400</td>
<td>23 ± 8</td>
<td>26 ± 7</td>
</tr>
<tr>
<td>2</td>
<td>1700-1800</td>
<td>30 ± 10</td>
<td>36 ± 10</td>
</tr>
<tr>
<td>3</td>
<td>1590-1610</td>
<td>36 ± 13</td>
<td>39 ± 11</td>
</tr>
<tr>
<td>4</td>
<td>1430-1470</td>
<td>38 ± 14</td>
<td>35 ± 10</td>
</tr>
</tbody>
</table>

The four wavelength regions used are detailed in table 6.4. Fig 6.4 shows the orbital light curve in the four different bands. The morphology in the different bands is similar, the better signal to noise in the longest wavelength band showing flaring activity. The modulation fractions of each of the bands, measured by eye (because other methods underestimate the modulation level) are shown in table 6.4. This shows that there is a slight indication of a 'turn-up' in the spectrum during the bright phase. The dip is present in each band, it does not significantly vary in depth between the different bands (see table 6.4).

6.4.2 Emission lines

Count rates covering the four major emission lines, namely C IV, He II, Si IV and N V, were then extracted using the wavelength ranges of table 6.2 and heliocentrically corrected to form light curves. They were folded into 200 phase bins (see fig 6.5). Orbital modulation is again present and is in fact much more prominent than in the continuum. There are clear bright and faint phases, the bright phase occurring from ~ 0.2 – 0.6, and the faint phase ~ 0.6 – 1.2. These are 0.3 later in phase than the EUV and soft X-ray bright and faint phases described in chapter 5. At phase 0.95-1.15 the flux drops further than the phases preceding it (except for the Si IV line). This may be due to a broader counterpart to the dip observed in the continuum at phase 1.00-1.05. The line flux does not go to zero.
Figure 6.4: Orbital period folds for each of the continuum bands. Morphology of the different bands is similar.
6.5 Radial velocity measurements

Phase resolved spectra were extracted for the two brightest emission lines, C IV and He II. The data were binned into 20 and 10 phase bins for the C IV and He II lines respectively. Trailed spectrograms for these two lines can be seen in figures 6.6 and 6.7. These show clear radial motion in both the lines. Extraction of individual profiles revealed that the emission lines can be decomposed into two components, a broad and a core component. Each of these line profiles were then fitted with double Gaussians to determine the centroid wavelength of each of the core and broad components (an example can be seen in figure 6.8). The two components probably arise from different regions of the stream as seen in the optical (due to the low wavelength resolution of this observation, the high velocity component and narrow component that are witnessed in the optical cannot be separated, but appear combined in the core component).

The centroid wavelengths were then converted to velocities using the known rest wavelength of the lines and using the relation

\[ \text{Velocity} = (\frac{\lambda_{\text{obs}} - \lambda_{\text{rest}}}{\lambda_{\text{rest}}}) \times c \]

(It has to be noted that the C IV line consists of a doublet at 1548.19Å and 1550.76Å with a flux ratio of 2:1 so therefore a weighted mean of 1549.05Å was used for the subsequent calculation.) The resulting radial velocities can be seen in figure 6.9. A constant (systemic velocity) plus a sinusoid was fit to each of the velocity data and resultant fits can be seen in table 6.5. The fit to the C IV core component is poor, probably due to the better quality data, indicating that there may be contamination by an unresolvable narrow component. The systemic velocity was heliocentrically corrected. It has to be noted that the systemic velocity will include both the z motion of the gas flow as well as some component from the binary motion.

Using relations in Rosen, Mason & Cordova (1987) and Schneider & Young (1980a, b), the systemic velocity, velocity amplitude, inclination and magnetic colatitude can be used...
Figure 6.5: Orbital period folds of the four major emission lines. Modulation is clearly evident, though it has a different morphology to the continuum. If a dip is present in the line modulation it is much broader than the continuum, spanning phases $\sim 0.95 - 1.15$. Each of the lines have a similar modulation morphology.
Figure 6.6: Trailed spectrogram for the C IV emission line.
Figure 6.7: Trailed spectrogram for the He II emission line.
Figure 6.8: A C IV line profile for one phase slice ($\phi \sim 0.8$). The Gaussians indicate how the core (dashed) and broad (dotted) components are fitted.

Table 6.5: The results of sinusoidal fits to the radial velocity data for the two components of the C IV and He II lines. The velocities quoted are for the semi-amplitude of the sinusoid. The rest wavelength of the C IV line is the weighted mean value for the doublet. The systemic velocities are heliocentrically corrected. Errors are quoted for a $\Delta \chi^2$ of 3.5.

<table>
<thead>
<tr>
<th>Line Component</th>
<th>Rest Wavelength (Å)</th>
<th>Rest Velocity (km s$^{-1}$)</th>
<th>Systemic Velocity (km s$^{-1}$)</th>
<th>Velocity Amplitude (km s$^{-1}$)</th>
<th>Blue to red Phase</th>
<th>$\chi^2/\nu$</th>
</tr>
</thead>
<tbody>
<tr>
<td>C IV core</td>
<td>1549.05</td>
<td>72.7±13.2</td>
<td>326±27</td>
<td>0.87±0.1</td>
<td>0.87±0.14</td>
<td>4.6</td>
</tr>
<tr>
<td>C IV broad</td>
<td></td>
<td>408±163</td>
<td>242±73</td>
<td>0.82±0.03</td>
<td>2.75±0.04</td>
<td>0.75</td>
</tr>
<tr>
<td>He II core</td>
<td>1640.4</td>
<td>26±14</td>
<td>752±409</td>
<td>0.80±0.15</td>
<td>0.80±0.15</td>
<td>0.2</td>
</tr>
<tr>
<td>He II broad</td>
<td></td>
<td>56±46</td>
<td>752±409</td>
<td>0.80±0.15</td>
<td>0.80±0.15</td>
<td>0.2</td>
</tr>
</tbody>
</table>
Figure 6.9: The radial velocity motion for the CIV and HeII lines. The top two panels indicate the core (left) and broad (right) motions for the CIV line. The bottom two panels are for the HeII lines core (left) and broad (right) components. Sinusoidal fits are performed for each of the motions and maximum red shift occurs at $\phi_{orb} \sim 0.1$. 
to find the location of the line emitting region of the stream. However due to the low resolution of the data and the lack of accurate knowledge of the geometry of QS Tel, this method cannot produce any useful limits on the location of the emitting region.

6.6 The dip: the occulted source spectrum

A more detailed spectral analysis of the dip has been carried out. Spectra during dip minimum ($\phi = 0.00 - 0.05$) and the immediate post-dip region ($\phi = 0.10 - 0.15$) were extracted. The dip spectrum was subtracted from the post-dip spectrum. If it is assumed that whatever obscures the source does not introduce a colour dependent flux reduction, the resulting 'difference' spectrum is that of the occulted source. A discussion of absorption processes and any colour dependencies that they may have can be seen in section 6.7.1.

The difference spectrum can be seen in fig 6.10. This spectrum has a 'turn-up' towards the blue end of the spectrum, the only line feature visible is a feature resembling a P Cygni profile of the CIV line, probably as a result of slightly different radial velocities of this line between the two spectra. On first examination this 'turn-up' is reminiscent of the Rayleigh-Jeans tail of a blackbody. Fitting the shape of this spectrum with a blackbody requires a temperature of $> 6 \text{eV}$ (no useful upper limit can be placed on the temperature). The best fitting absorbed blackbody model found for the mean EUVE data taken in September 1993 ($kT \approx 15 \text{eV}, N_H \approx 4.3 \times 10^{19} \text{ atoms cm}^{-2}$; Rosen et al. 1995) was then extended out to HST wavelengths, see fig 6.11. This revealed that this model overestimates the HST flux by a factor of $\sim 4$, but is broadly consistent in shape with the HST data. Two factors could contribute to the mismatch of the HST flux: (i) if the absorption is not complete this will result in an under estimate of the flux when the dip spectrum is subtracted from the post-dip spectrum (though for the blue 'turn-up' to be due to emission from the occulted source the absorption has to have no colour dependence). (ii) The flux of QS Tel in the EUV is known to vary by factors of up $\sim 400$ within the timescale of a few months (Rosen et al.1995). As this observation occurred a year later than the EUVE observation it is reasonable to suppose that the flux could have changed by a factor of 4. The spectrum may also have changed in this time. Note, this variability
Figure 6.10: The 'post-dip minus the dip spectrum', indicating the spectral shape of the occulted source (assuming the absorption mechanism is not spectrally dependent). The feature at ~ 1550Å resembles a P-Cygni profile of the C IV line, which is most probably an artifact of the subtraction caused by the radial velocity shift between the post-dip and dip phases.

also means that firm conclusions should not be made from matching these two spectra.

6.7 Discussion

This HST FOS observation has given us a far more detailed view of the UV emission from QS Tel than the previous IUE data has. Although the mean spectrum is consistent with the earlier IUE observation both in flux and shape, this new observation accurately measures the UV continuum and emission line orbital modulation. A narrow dip is observed at phase 0.0-0.05 (phase 0.0 corresponds to the blue/red crossing phase of the optical narrow
Figure 6.11: The HST spectrum of the region occulted by the stream at phase 1.0-1.06 (red) plotted with the EUVE spectrum from Sept 1993 (blue). The model shown (black) is the best fit model to the EUVE data alone (blackbody temperature ~ 15 eV, $N_H \sim 4.3 \times 10^{19}$ atoms cm$^{-2}$).
emission line, Schwope et al. 1995), which is coincident with a dip observed both in the EUV, soft X-ray and optical (Rosen et al. 1995). The lines have a prominent orbital modulation. The bright and faint phases at \( \phi \sim 0.2-0.6 \) and 0.6-1.2 respectively. If the dip is present in the lines then it is much broader, spanning phases 0.95-1.15. Radial velocities of the CIV and He-II lines (core and broad components) have been determined. The spectrum of the occulted source has also be examined.

Although QS Tel was probably in a bright state, it is unclear whether it was accreting onto one or two poles. The flux measured is consistent with that observed by IUE a year earlier. This IUE observation took place just two weeks before an EUVE observation found QS Tel in a two pole accretion mode. If QS Tel was in a two pole accretion mode during this observation, then the lack of double peaked continuum modulation may suggest that the UV light is not dominated by emission from the polar accretion region, or one of the poles does not emit in the UV, or that both poles are visible most of the time. In order to make some determination of what is causing the continuum modulation, spectra were summed for the peak of the bright phase flux (\( \phi = 0.15 - 0.40 \)) and the minimum of the faint phase flux (\( \phi = 0.7 - 0.9 \)). The subtraction of the faint phase from the bright phase can be seen in fig 6.12. However due to the presence of strong variable lines, any continuum change is hard to discern. A ratio of the bright to the faint flux was then produced (see fig 6.12). This indicates that there is a difference in the line fluxes which is as expected, as they are emitted from a different region and have a different phasing (see fig 6.5), but there is also a slight difference in the continuum shape, with a rise below 1400Å. This has the same profile as that observed in the occulted spectrum which was consistent with the Rayleigh Jeans tail of the blackbody from the emission region. This may also be the same feature indicating that the continuum modulation is partially attributable to the accretion region.

The accretion region is supposedly optically thick. If so, the Rayleigh Jeans tail is also optically thick and therefore it must introduce a variation. If two poles were present (as for EUVE), then two peaks from the Rayleigh Jeans tail would be present in the light curve below 1400Å. If this is not observed then, if the ratio does show emission from the accretion region, this implies one of three situations; either QS Tel is in a one pole accretion mode, or the second pole is not emitting in the UV or, as in the case of AM Her (Mazeh, Kieboom, Heise 1986), the one and two pole accretion modes are indistinguishable using
their UV light curves.

The fluxes and velocities of the strongest emission lines (CIV, HeII, SiIV and NV) vary in a similar fashion. Radial velocity measurements found maximum red shift (and thus inferior conjunction of that part of the stream (providing the emission is infall dominated) at phase 0.1, which is also the phase of minimum line flux. The maximum line flux occurs at $\phi_{\text{orb}} = 0.3 - 0.5$, around the phase of first quadrature. This is consistent with the stream being the source of the line emission if the line emission is optically thick. The variations in flux are most likely due to changing projected areas. For a uniformly emitting straight cylindrical stream the emission line variation would be expected to form a double peaked lightcurve with the maxima occurring at the two quadrature phases, as the stream is being viewed side on. The lack of a second maximum may indicate that the stream is preferentially irradiated on one side. A curved stream could produce this. As the stream leads the secondary star (which is expected from binary dynamics i.e the effects of the coriolis force), at the first quadrature phase the concave side of the stream is in our line of sight, at the second quadrature phase it is the reverse, convex part of the stream that is being observed. It is reasonable to assume that the concave face of the stream would intercept more flux than the convex, and hence the second quadrature phase would have less intensity than the first quadrature phase. The phase of the dip occurs during line flux minimum, where the stream is crossing our line of sight and the projected area of the stream is small, this is consistent with the dip being a result of occultation of the emission region by the stream. The total flux emitted by the CIV, HeII, SiIV and NV lines is $3.8 \times 10^{-12}\text{erg cm}^{-2}\text{s}^{-1}$ compared to a value of $3.1 \times 10^{-12}\text{erg cm}^{-2}\text{s}^{-1}$ as measured by IUE (de Martino et al. 1995). This emitted flux can be used to derive an estimate of the irradiation of the stream (Mukai 1988). The observed flux from the main pole (assuming $kT=25\text{eV}$) above $50\text{eV}$ (a mean value for the temperature required to photoionize the gas stream) is $9.3 \times 10^{-11}\text{erg cm}^{-2}\text{s}^{-1}$, therefore 4% of the flux is intercepted by the stream (de Martino et al. 1995) also find a similar value). This indicates that the area intercepting such radiation is small and either a large distance from the white dwarf or the emission just intercepts a small part of the stream.

Rosen et al. (1995) completed a similar analysis of the radial velocity motion for the H$\beta$ (4861Å) optical emission line. Again they could not deconvolve the core component
Figure 6.12: Top panel: The bright-faint phase difference spectrum. Bottom Panel: The ratio of the bright phase spectrum and the faint phase spectrum. This clearly shows a 'turn-up' below 1400Å.
into the narrow emission line component and the high velocity component. Schwope et al. (1995) only detail the results for the narrow emission line, so no comparison to their results can be made. Our measurements of the velocity amplitudes of the core component for the CIV and HeII lines, are slightly higher than that for the Hβ line. The broad component velocity amplitudes are consistent for all the lines as are the blue to red crossing phases. The CIV core component systemic velocity is consistent with that measured for the Hβ line, however the other components, perhaps as a result of poorer line decomposition during the fitting process, are not. It has to be noted that the phase convention employed refers to the blue/red crossing phase of the narrow emission line (Schwope et al. 1995). If this component originates from the secondary, this phase marks inferior conjunction of the secondary. This apparently occurs before inferior conjunction of the stream, and suggests that the stream lags the secondary, which is in conflict with the conventional picture in which, due to the coriolis force, the stream leads the secondary. This may indicate that the phasing of the narrow emission line is not known as accurately as suggested by Schwope et al. (1995). Alternatively Schwope et al. (1995) incorrectly ascribed the narrow emission line component to the secondary star. It might also indicate that the accreting pole lags the secondary and emission comes from close to the white dwarf. If this were the case the previous discussion of equivalent widths is invalid.

The narrow dip was examined. The difference spectrum of the post-dip to dip spectra was consistent with the Rayleigh-Jeans tail of a blackbody of temperature > 6eV. This analysis did require the assumption that the absorbing material did not have any colour dependence. The stream is assumed to occult the accretion region and various mechanisms could be involved in such absorption. I now discuss the various mechanisms and limits (if any) we can place on $N_H$.

6.7.1 Absorption mechanisms

There is only a slight difference in the flux of the lines between the post-dip spectrum and the dip spectrum, indicating that the emission lines are produced in a different region to the source that is occulted, so to simplify the further analysis the lines are removed. In the absence of knowledge about the exact spectral shape of the post-dip continuum, I crudely
Figure 6.13: The post-dip spectrum ($\phi = 0.10 - 0.15$), with the lines removed for ease of analysis, crudely fitted with a powerlaw model of index 1.4.

represented the spectrum with a powerlaw model with an index of 1.4, see fig 6.13.

This prescription of the spectrum is highly simplified. The UV spectrum will consist of the Rayleigh Jeans tail at short wavelengths and at least one other component which dominates at wavelengths of $\gtrsim 2400\text{\AA}$. The other components could originate from the white dwarf, the stream and cyclotron emission. Over the wavelength band studied here it is impossible to deconvolve these other components from the component of interest (the Rayleigh Jeans tail), so they are not included in further analysis. Thus all the next discussion is rather speculative, but crude qualitative statements can be made.

This powerlaw was then fixed and to achieve the dip spectrum various absorption mechanisms were applied. The required $N_H$ was determined by varying the absorption until the model count rate equaled the observed count rate. Examination of the resulting spectral shape indicates whether the model is reasonable.

**Free-free absorption**

Free-free absorption (or thermal bremsstrahlung absorption) occurs when radiation is absorbed by free electrons in a field of ions. The absorption coefficient for this process is
where $\xi$ is the velocity averaged gaunt factor, $\alpha_{\text{eff}}$ is in units of cm$^{-1}$, $T$ is the temperature of the absorbing medium (K), $E$ is the energy of the incident radiation, $h$ is the Planck constant, $Z$ is the atomic number of the material, $n_e$, $n_i$ are the densities of the electrons and ions respectively and $k$ is the Boltzmann constant. This was coded to give a multiplicative model which could be used within the XSPEC spectral fitting package. The resulting model calculated the multiplicative factor as a function of the absorption coefficient and the density through the material. The model had three parameters, namely; $N_{\text{H}}$, temperature of the absorbing gas and the path length through the stream. As the $N_{\text{H}}$ and mean free path are not independent one of them has to be fixed during fitting. In this case the path length through the stream was fixed at the upper limit to the size of the gas stream ($10^9$ cm). Fig 6.14 shows variation of the transparency of the absorber as a function of wavelength (the two vertical lines indicating the HST observation window) for various values of $N_{\text{H}}$, temperature and mean free path. This clearly shows that the absorption is highly colour dependent, with greater absorption occurring at redder wavelengths.

This colour dependence can produce the blue 'turn-up' observed in the 'difference' spectrum, but as free-free absorption is a strong function of wavelength ($\propto \lambda^2$), a factor of $\sim 8$ difference between the blue and red ends of the HST FOS spectrum would be expected (if other components are present this is reduced).

Fitting the dip spectrum data with the previously determined power law and free-free absorption (the powerlaw parameters frozen at the post-dip values) requires an $N_{\text{H}}$ of $4.4 \times 10^{22}$ atoms cm$^{-2}$ (assuming a temperature of the absorbing gas of 10,000K (because the presence of emission lines requires temperatures of 5,000-35,000 K (Ferrario & Wickramasinghe 1993)) and a path length through the stream of $1 \times 10^9$ cm). This can be seen in fig 6.15. This clearly does not have the correct shape to fit the data with the red end of the spectrum being over absorbed and the blue end not being absorbed enough.
Figure 6.14: The multiplicative factor plotted against energy is shown for the free-free absorption mechanism for various values of temperature (T), \( N_H \) and path length through the stream (\( pl \)). These are calculated for two different values of \( N_H \). For each value of \( N_H \) the solid lines indicate a T of the medium of 10,000K and a \( pl \) of 10^6 cm, the dashed line a T of 20,000K and \( pl \) of 10^6 cm and the dotted line a T of 10,000K and \( pl \) 5 x 10^8 cm.

Figure 6.15: The dip spectrum (\( \phi = 0.00 - 0.05 \)) with a free-free absorbed powerlaw model. The powerlaw component is frozen at the post-dip value and the free-free absorption is increased until the required dip count rate is achieved, in this case an \( N_H \) of 4.4 \times 10^{23} \) atoms cm\(^{-2}\).
Bound-free

Bound-free absorption (or photoionisation) was initially tested using a crude model within XSPEC developed by Done et al. (1992). This model does not include any effects occurring below 5eV (>2500Å), therefore the Balmer edge and He edges are not taken into account. A fit to the data results in an $N_H$ of $8.6 \times 10^{10}$ atoms cm$^{-2}$ (with $T=10,000K$) see fig 6.16. To get the required reduction in count rate an absorption of $N_H$ of $1.7 \times 10^{20}$ atoms cm$^{-2}$ is needed (assuming a temperature of the absorbing gas of 10,000K) see fig 6.16. This model has some absorption edges. There is a slight indication of an edge at ~ 2050Å (from excited HeII) in the dip spectrum which is not as obvious in the post-dip spectrum. This model could produce a blue 'turn-up', but a strong edge at ~ 1400Å (from excited CIV) should also be present. However as this model does not include edges from longer wavelengths it will be producing much larger edges in this wavelength band to produce the required reduction in count rate. Optical observations at the Balmer edge and He edges would measure the level bound-free absorption more accurately, and hence determine whether bound-free is the significant absorption process. At present these models and data cannot show which type of absorption process is occurring.

The theoretical opacities for bound-free and free-free absorption by Hydrogen, are:

$$\kappa(H_{\text{bf}}) = \alpha_s \lambda^3 \sum_{n=2}^{\infty} \frac{g_n'}{n^3} e^{-\left(\kappa/\kappa_T\right)}$$  \hspace{1cm} (6.1)

where $\alpha_s = 1.044 \times 10^{-26}$ for $\lambda$ in Å and $g_n'$ is the Gaunt factor ($\approx 1.0$), $\chi$ is the excitation potential ($= 13.60(1 - 1/n^2)$ eV),

and

$$\kappa(H_{\text{ff}}) = \alpha_s g_r \lambda^2 \log_e \frac{\rho}{2\Theta} \times 10^{-31}$$  \hspace{1cm} (6.2)

where $g_r$ is the Gaunt factor for free-free absorption ($\approx 1.512$), $\Theta = 5040/T$ and $I=13.6$ eV. At 1500Å and a temperature of 10,000K these opacities are $3.3 \times 10^{-23}$ and $2.4 \times 10^{-25}$ cm$^2$/H for bound-free and free-free absorption respectively. This is roughly consistent with the observed picture in which the bound-free model requires less absorbing
Figure 6.16: The two panels show the dip spectrum (with the lines lines removed) with the powerlaw plus bound-free absorption model. The powerlaw in both cases is frozen at the post-dip level. The top panel shows the best fit value for the $N_H$ ($8.6 \times 10^{19}$ atoms cm$^{-2}$) and the bottom panel shows the model for the value of $N_H$ required to produce the observed count rate ($N_H = 1.7 \times 10^{20}$ atoms cm$^{-2}$).
material to produce the required reduction in count rates. However much more complex models and better data will be required to make quantitative measurements.

Partial covering

The gas stream, as pictured in the 'blob' accretion model (Kuijpers & Pringle 1982), consists of a series of opaque blobs in a tenuous medium. Absorption by such a stream would be a combination of complete absorption by the blobs and either bound-free or free-free absorption by the tenuous inter-blob medium. The quality of the data do not allow us to deconvolve these. For the 'blobs' to be opaque to UV radiation an $N_H \geq 6 \times 10^{22}$ atoms cm$^{-2}$ is required (using a multiplicative factor of 0 for the free-free model used earlier). The colour dependence of this picture depends on the relative strengths of the two components, the absorption by the 'blobs' having no colour dependence and the tenuous medium having a colour dependence dependent on density. In this picture it might be expected that density variation along the stream may allow emission to seen during the dip, appearing like flares. Flares are seen during dip minimum in the EUVE observation (see chapter 5, Rosen et al. 1995). This is a more complex picture of absorption but may be a more realistic description, but this cannot be tested with this data and models presented here.

6.7.2 Conclusions

This HST observation of QS Tel showed clear orbital modulation in the lines and continuum. Radial motion measurements are as expected for lines produced within the accretion stream. A dip is observed in the continuum and possibly a much broader component in the lines. The post-dip minus dip difference spectrum reveals a blue 'turn-up', which is consistent with a Rayleigh Jeans blackbody tail of temperature $> 0.6\text{eV}$, suggesting that the occulted source is the polar accretion region. A similar 'turn-up' is witnessed in the ratio between the bright and faint phase spectra, also indicating that at least some part of the orbital modulation is due to the polar accretion region.

An examination of the various absorption mechanisms that may occur during the dip are
discussed, though no quantitative results can be gained due to the complicated nature of the UV continuum. However, it can be said that bound-free absorption is likely to dominate over free-free absorption as the opacity is approximately 100 times greater. Better quality data during the dip have the potential of providing detailed information about the nature of the gas stream, and thus, direct observational evidence of the existence (or not) of inhomogeneous accretion.
Chapter 7

Concluding Remarks.

The observations in this thesis have shown the great amount of information about these cataclysmic variables which can be gained from satellite data. The ROSAT data of the polar EK UMa showed the best orbital light curve at this time, revealing features that had not previously been observed, namely the complete light curve and the deep, narrow dips attributed to occultation of the emission region by the accretion stream. The soft X-ray spectrum was also characterised for the first time, and was found to have a high blackbody temperature and a large soft X-ray excess.

The observations of the proposed IPs SW UMa and 1H0709-360 did not have sufficient signal to noise for any conclusions as to their class to be made, thus leaving their IP status open to debate. However, modulation limits and the spectral nature were determined for SW UMa.

The ROSAT data of the peculiar IP AE Aqr revealed a wealth of information. Detailed measurements of the white dwarf spin modulation were made and compared to the un-modulated flux level. X-ray flaring behaviour was observed for the first time. The flares were found to be simultaneous with the optical and UV. AE Aqr has shown itself to be unique amongst IPs in having a soft two temperature optically thin plasma emission spectrum. Comparisons to other IPs were made, as well as to coronally active systems, as the spectrum is reminiscent of such systems. The recent model of Wynn, King & Horne
(1995) was also discussed. Although some coronal activity on the secondary could not be ruled out, the 'propeller' model of accretion (Wynn, King & Horne 1995) is the most promising model for accretion in AE Aqr.

Many observations have been made of QS Tel since its discovery. The ROSAT, EUVE and HST observations reported in chapters 5 and 6 have shown this to be a highly interesting polar. Firstly QS Tel has shown three distinctly different accretion morphologies (two reported here) ranging from complete cessation of accretion, to one and two pole accretion modes. Also present in the photometric data from the ROSAT, EUVE and HST observations is the presence of a deep, narrow dip attributed to occultation of the emission region by the accretion stream. Spectrally QS Tel is unremarkable, with absorbed blackbody \( N_H \sim 3 \times 10^{19} \text{ atoms cm}^{-2} \), \( K_T \sim 20 \text{ eV} \) plus bremsstrahlung emission.

These observations have shown that multiwavelength studies are vital for the understanding of magnetic cataclysmic variables, each wavelength providing valuable information to compliment other wavelengths. However due to the variability of CV behaviour (recall for example the light curves of AE Aqr and QS Tel) simultaneous measurements in other wavelengths are of much greater use (if not vital). Thus, it would be helpful for future satellite missions to have at least optical bolometers on board so one other wavelength region can be covered, however, no missions planned to date have such an optical bolometer.

The range of unanswered questions and lack of precise measurements also shows that despite great improvements in satellite technology and data in the last ten years, better resolution instruments are still required. These instruments are already on there way, or are in use, such as ASCA (in use), Jet-X, XMM and AXAF.

As new satellites and detectors come into use atomic theory must be constantly improved. The excellent spectral resolution of ASCA is already revealing problems with the atomic data which is currently in use. Another feature of increased spectral resolution is that the simple spectral models we use today will no longer work with such good data. Refining the models to include more realistic situations is vital (e.g. including varying densities in absorption column models, rather than using a constant slab of material).

The HST data used in Chapter 6 shows how good such data is even at low resolutions.
Studying accretion stream dips with the HST at higher spectral resolutions would provide a wealth of information of the structure of the stream. Other improvements in the study of accretion stream dips, could be made by the use of codes such as 'cloudy' as these model the effects of absorbing materials in much greater detail.
Appendix A

Instrumentation.

A.1 Introduction

In this appendix I will describe the two satellites from which the majority of the data used in this thesis has been gained. These are the ROSAT satellite, which covers the EUV and soft X-ray, telescope and the Hubble Space Telescope which allow optical and UV measurements to be made. Where appropriate, details of analysis procedures are included.

A.2 ROSAT

ROSAT, the ROentgen SATellite (see fig A.1), is an X-ray observatory developed through a cooperative programme between Germany, the United States and the United Kingdom. It is described by Trumper 1983; Pfefferman et al. 1987; Briel et al. 1995; Aschenbach et al. 1985. The satellite was designed by Germany and was launched into a low Earth orbit (see table A.1) from Cape Canaveral, USA, on June 1st 1990. The satellite operations are controlled from the German Space Operations Center (GSOC) ground station, Oberpfaffenhofen, Germany. There are five or six ground contacts every day.
Figure A.1: A schematic View of the ROSAT Spacecraft
ROSA T is a three axis stabilized satellite using 4 gyros and 3 reaction wheels. The high torque of the momentum wheels allows fast slews (180° in ~ 15 mins) so that ROSAT can observe two targets on opposite hemispheres each orbit. The pointing is accurate to ~ 1 arcmin with a ~ 10 arcsec jitter radius.

The scientific payload consists of:

- The X-ray telescope (XRT). The XRT has a mirror assembly consisting of four nested Wolter-1 mirrors. The focal plane instruments consist of a carousel on which there are two position sensitive proportional counters (PSPCs) and a high resolution imager (HRI).
- The wide field camera (WFC). The WFC mirror assembly consists of three nested Wolter-Schwarzchild mirrors, which are co-aligned to the XRT. At the focal plane the instrumentation is made up from a curved microchannel plate (MCP) detector and a carousel containing eight filters, of which six are science filters.

A.2.1 The XRT and PSPC

XRT

The XRT consists of

- the X-ray mirror assembly (XMA)
o a magnetic deflector

o two South Atlantic Anomaly Detectors (SAADs)

o a focal plane turret containing the two PSPCs and the HRI

The XMA consists of four nested grazing incidence Wolter-1 mirrors with a maximum aperture of 83.5 cm and a focal length of 240 cm. All 8 mirror shells (four paraboloid/hyperboloid pairs) are constructed of Zerodur, a glass ceramic with an almost negligible thermal coefficient, and are coated with a thin layer of gold to enhance the X-ray reflectivity. Typical grazing angles are between 1° and 2° depending on the subshell considered. The focal plane plate scale of the XRT is 11.64 μm arcsec⁻¹.

The effective area of the XRT is a function of the off-axis angle and energy due to combined effects of:

- The reflectivity of the gold coating on the XRT mirrors is a decreasing function of energy with a number of minor absorption edges.

- Off-axis rays strike the mirror surface at a shallower angle such that the projected geometric area decreases. Also the average reflectivity of gold is reduced. In addition off-axis rays suffer a higher degree of obscuration due to the radial struts of the XMA. Collectively these lead to an energy dependent decrease in effective area with increasing off-axis angle, known as vignetting.

- Small scale irregularities in the mirror surface give rise to small angle scattering, the level of which is energy dependent.

**PSPC**

Two redundant Position Sensitive Proportional Counters (PSPCs) are mounted on the focal plane of the XRT, two flight-spare detectors were also constructed and used in ground calibration measurements. The PSPCs are multiwire proportional counters, each consisting of essentially two separate counters: the anode A1 and the cathodes K1 and K2.
Figure A.2: Schematic cross-sectional view of the PSPC, showing the arrangements of the anodes and cathodes.

The grid system is accommodated in a gas filled counter housing containing a mixture of 65% argon, 20% xenon and 15% methane. The PSPC entrance window consists of a ≈ 1 \( \mu \)m foil of polypropylene ([C\(_2\)H\(_4\)]\(_n\)) coated with ≈ 50 \( \mu \)m cm\(^{-2}\) graphite and...
An X-ray photon, passing through the thin plastic window will be photo-electrically absorbed by the counter gas producing a photo-electron. The primary electron(s) is thermalised and in the process causes the ionisation of other gas atoms producing a secondary electron cloud. The positive ions (including the initial ion) can also add to this cloud by Auger and shake-off processes when energetically allowed. The mean energy required to create a secondary electron is approximately constant over the entire energy range of the PSPC. The probability distribution function of the number of secondary electrons can be described by a Gaussian whose width is given by the modified Fano factor. The num-
The electron cloud drifts through the cathode K1 towards the anode A1, near A1 the electric field strength is sufficiently high that the charge cloud is amplified by a (gain) factor of $5 \times 10^4$. The avalanche of charge onto the anode leads to a charge signal at the anode and an induced signal at the cathodes, which can be sensed by charge-sensitive preamplifiers. The cathode grids are used to determine the position of the incident X-ray, meanwhile the anode grid determines its energy.

Each cathode grid is divided into 25 cathode strips, each strip consisting of seven cathode wires, and is connected to a preamplifier. Dependent upon the energy of the photon, the cathode signal is induced on 3-5 strips. These signals are digitized after passing through pulse shapers and peak detectors by analog-to-digital converters (ADCs). After checking the validity of the event, the position is calculated. To get the position in both co-ordinates, the two cathode grids are arranged perpendicular to each other.

The wires of the anode grid are connected in parallel to one preamplifier with pulse shaper, peak detector and ADC. Its signal is used to obtain the energy of the X-ray photon and also as a trigger signal for the read-out electronics. The time required by the electronics for conversion, validity checking and position calculation is $< 1$ ms, leading to a maximum event rate of $> 10^3$ s$^{-1}$. The energy resolution of the PSPC is $\frac{\Delta E}{E} = 0.43(E/0.93)^{1/2}$ (FWHM).

Besides the signal from the 'veto' anode (A2), the signals of the outer cathode wires K1 and K2 are fed to an anti-coincidence circuit. The gas depth in the veto counters is sufficient such that ionising particles, as well as resonant scattered X-rays (coming from the frames and counter body), are detected. The background rejection efficiency for the PSPC is around 99.8%.

Four different components have been identified to the non-cosmic background of PSPC observations. The two best understood components are the particle background and the solar X-rays scattered into the field of view by the Earth's atmosphere. The other components have been identified through their temporal variations, as observed in the ROSAT sky.
survey (Snowden & Schmitt 1990, Voges 1992). One of which is short term enhancement ($\ll 1$ to 2 orbits) which includes auroral X-rays and the other is long term enhancement (1-2 days). Other background components are from cosmic rays and charged particles in the local space environment, these have been modelled by Snowden et al. (1992) and Plucinsky et al. (1993) for master veto rates of $< 170$ cts $s^{-1}$.

Whether the cosmic X-ray background is a source of contamination or an interesting signal is obviously dependent on one's scientific point of view. In most observations and for all energies less than $\sim 2$ keV it is typically the dominant background component. The cosmic X-ray background is both spatially and spectrally anisotropic with independent variations of greater than a factor of five in surface brightness in the hard and soft bands. The structure is due to both variations in emission and foreground absorption. The angular scale of the structure varies from the coarse all-sky negative correlation between galactic $N_H$ and the 1/4 keV background to few arc-minute scale emission by clusters of galaxies (most obvious at higher energies). Thus, there is no "typical" cosmic background flux or spectrum.

A filter wheel with four positions is mounted in front of the detector. The open and closed filter wheel positions are used for "standard" observations and for monitoring the particle background. The third position contains a Boron filter, the insertion of which allows an increase of spectral resolution at low energies. The PSPC filter wheel carries three calibration sources in the fourth position, which produce aluminium (1.49 keV) fluorescent K-shell photons by X-ray excitation from Fe-55. These X-rays are used to illuminate the PSPCs in order to measure in-orbit temporal variations in the PSPC gain.

The combination of the XRT and PSPC

The XRT and PSPC provide a 2° field of view (fov) for sources with emission in the 0.1-2.4 keV energy range.

The combined point spread function (PSF) of the XRT and PSPC is the convolution of five components:
1. The XRT mirror assembly (XMA) scattering angle
2. The off-axis blur of the XMA
3. The intrinsic spatial response of the PSPC
4. Focus and detector penetration effects
5. A widening of the PSPC resolution due to the existence of 'ghost images' (for PHA channels < 20 only)

The mirror dominates the PSPC components of the PSF outside a radius of ~ 14 arcmin. The PSF is thus a very strong function of off-axis angle and energy. For off-axis angles greater than 25 arcmin the PSF is almost energy independent, with a FWHM of 180 arcsec at a radius 40 arcmin off-axis. On-axis, the PSF FWHM is 47 arcsec (25 arcsec) at an energy of 0.2 keV (1.5 keV). Examination of a bright point which is observed off-axis reveals asymmetry in the distribution of the source counts which becomes easily noticeable once the source is observed outside the central ring of the PSPC.

The PSPC cannot always properly position X-ray events with very low energy. This results in an extension to the image (ghost image) caused by the detector response to very low energy X-ray photons, but one which will not fit any PSF model. The effect is very pronounced below channel 15 (~ 0.15 keV) but is not evident above channel 20 (~ 0.2 keV).

In the vicinity of the center of the field of view the focus of the X-ray beam, from a point, is so sharp there is significant shadowing behind the coarse mesh. A slow spacecraft wobble (amplitude 3 arcmin, period 400s) is therefore performed in pointed observations in a direction diagonal to the mesh structures. In this fashion, the coarse mesh grid will be smoothed out when projected onto sky co-ordinates, hence permanent occultation of X-ray sources behind the coarse mesh is not possible. This results in a modulation of the source flux at this period and its harmonics. To diminish such modulation a source can be positioned 40 arcmin off-axis, where the point spread function is much larger, and the shadowing by the wires is reduced.
Figure A.4: The effective on-axis area of the ROSAT XMA+PSPC combination with (dashed line) and without (solid line) the boron filter.

The quantum efficiency of the PSPC is dominated by the presence of a carbon edge at 0.28 keV due to material in the detector window. Below the carbon edge the window transmits about 50% of the incident X-ray flux. At energies of just above the carbon edge (0.3 keV) the window is opaque, whereas at higher energies (up to 2 keV) the efficiency is close to 100%. The overall effective area of the XRT and PSPC is a convolution of the XRT effective area (dependent on energy and off-axis angle) and the quantum efficiency of the PSPC, see fig A.4.
Analysis

On-board electronics digitise pulse height information into 256 raw channels over the entire energy range of $\sim 0.1 - 2.4$ keV$^1$. The time resolution of the data is 150 $\mu$s. The data is preprocessed in Germany (the SASS processing). The processing includes corrections for spatial and temporal gain variations, resulting in pulse-height-invariant channels (PIs). The PI channels which are valid for use (since October 1991) are 11-256. Times of high master veto rate ($> 170$ cts s$^{-1}$) are excluded from the further analysis. Data are supplied to the user in the form of 3 good event lists (1 total energy range, the other to be soft and hard energy bands) and various housekeeping and response files.

Analysis proceeds further using the Starlink ASTERIX analysis package (Saxton 1992). The image is displayed and a circular region is selected around the source of interest, whose size is dependent on the psf in that region of the detector, being careful to allow for ghost imaging in soft sources. To provide background region either an annulus is chosen around the source (can only be done if this region is free of sources) or another circle (preferably at the same off-axis angle so as that the same vignetting correction can be applied). The counts in each of the source and background regions are then summed at a specified time binning. The particle rates are then calculated (as a function of PI channel number and time) and subtracted from the background count rate, this is typically a small value of $\sim 1 - 3\%$ of the total. If necessary the background counts are corrected for vignetting to the source position. The background counts are then normalized for the difference in the source and background regions areas. The background counts are then subtracted from the source. Corrections are then applied for vignetting, exposure, dead time and wire obscuration.

For spectral analysis essentially the same procedure is applied as above, except the counts are just binned into one time bin, but several energy bins. Analysis proceeds using the XSPEC package (Shafer et al. 1991). The energy bins are binned into a form recommended by the Germans, called the SASS binning. The 256 raw channels are binned into 34 spectral channels. The response matrix depends on the period of observation. For observations

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$^1$The PSPC gain was lowered during October 1991, this resulted in the upper limit increasing to $\sim 3.5$ keV. However the PSPC effective area is not well calibrated for energies $E \gtrsim 2.0$ keV.
Table A.2: The HRI performance summary

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Field of View</td>
<td>38 arcmin (square)</td>
</tr>
<tr>
<td>Spatial Resolution</td>
<td>1.7 arcsec (FWHM)</td>
</tr>
<tr>
<td>Quantum Efficiency</td>
<td>30% at 1 keV</td>
</tr>
<tr>
<td>Window Transmission</td>
<td>75% at 1 keV</td>
</tr>
<tr>
<td>Background</td>
<td>$1.0 \times 10^{-6}$ internal</td>
</tr>
<tr>
<td>(cts arcmin$^{-2}$ s$^{-1}$)</td>
<td>$(1.7-8.3) \times 10^{-3}$ external</td>
</tr>
<tr>
<td></td>
<td>$(3.5-13.8) \times 10^{-4}$ X-ray background</td>
</tr>
<tr>
<td></td>
<td>$3.8 \times 10^{-3}$ Typical Total</td>
</tr>
<tr>
<td>Temporal Resolution</td>
<td>61 µsec</td>
</tr>
<tr>
<td>Dead Time</td>
<td>0.36-1.35 msec</td>
</tr>
</tbody>
</table>

since October 1991 (which all the observations in this thesis are) the matrix DRM 36 is used. As there are still uncertainties in the PSPC gain, a systematic error of 2% is included in the data analysis.

A.2.2 The HRI

Also in the focal plane of the XMA is the High Resolution Imager (HRI), which is practically identical to the Einstein Observatory HRI and is comprised of two cascaded microchannel plates (MCPs) with a crossed grid position readout system.

The performance of the HRI is characterised by several parameters. These are the field of view, the detector quantum efficiency, the window transmission, the sources of background, the spatial and temporal resolution and the overall detector plus electronics event capacity. These properties are briefly summarized in table A.2.
A.3.3 WFC

The WFC (Wells et al. 1990; Simms et al. 1990; Pye, Watson & Pounds 1991) is an autonomous instrument having its own star tracker, thermal control system and command and data handling system. The WFC does rely, however, on the ROSAT spacecraft for power, on-board data storage and telemetry. A schematic of the WFC can be seen in fig A.5. The WFC optics consist of a nested set of 3 Wolter-Schwarzchild Type I mirrors, fabricated from aluminum and coated with gold for maximum reflectance. The mirrors provide a geometrical collecting area of 475 cm² with a focal length of 525 mm. The grazing incidence angles chosen (typically ≈ 7.5°) allow the collecting area to be optimised whilst retaining a wide (2.5° radius) circular field of view and a low energy reflectivity cut-off at 0.21 keV. The on-axis resolution is ≈ 2.3' HEW, but the response degrades to ≈ 4.4' HEW at 2.5° off-axis due to inherent optical aberrations.

To take full advantage of the telescope's resolution the pair of MCPs in the detector are both curved, to match the optimum focal surface. A CsI photocathode is deposited directly onto the front face of the front MCP to enhance the EUV quantum efficiency. The detector resolution is better (by a factor of > 2) than that of the mirror nest and consequently does not contribute significantly to the WFC performance. During flight one of two identical detector assemblies could be chosen.

The filter wheel assembly consists of eight filters, six of which are science filters, see table A.3. The science filters consist of two pairs of redundant 'survey' filters and two pointed phase filters. The pointed filters are smaller than the survey filters, and a large amount of vignetting is present in the outer parts of the field when using these pointed filters.

Fig A.6 shows the on-axis effective area for the complete WFC for each filter design. The effective area decreases with increasing off-axis angle, dropping linearly to 67% of its on-axis value at 2.5° off-axis.

The point response function of the WFC is dominated by the performance of the mirrors. Figuring errors determine the on-axis response but aberrations inherent in the optical design dominate at the edge of the field of view. As the energy of the incident photons
Figure A.5: Schematic of the WFC, showing its major components
Table A.3: The filters used on the WFC

<table>
<thead>
<tr>
<th>Filter name/type</th>
<th>Survey/Pointed</th>
<th>Mean Energy (eV)</th>
<th>Bandpass (eV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>S1a: C/Lexan/B</td>
<td>S+P</td>
<td>124</td>
<td>90-185</td>
</tr>
<tr>
<td>S1b: C/Lexan</td>
<td>S+P</td>
<td>124</td>
<td>90-210</td>
</tr>
<tr>
<td>S2a: Be/Lexan</td>
<td>S+P</td>
<td>90</td>
<td>62-111</td>
</tr>
<tr>
<td>S2b: Be/Lexan</td>
<td>S+P</td>
<td>90</td>
<td>62-111</td>
</tr>
<tr>
<td>P1: Al/Lexan</td>
<td>P</td>
<td>69</td>
<td>58-83</td>
</tr>
<tr>
<td>P2: Sn/Al</td>
<td>P</td>
<td>20</td>
<td>17-24</td>
</tr>
<tr>
<td>OPQ: opaque</td>
<td>(S+P)</td>
<td>-</td>
<td>-</td>
</tr>
<tr>
<td>UV: UV interference</td>
<td>-</td>
<td>-</td>
<td>-</td>
</tr>
</tbody>
</table>

Figure A.6: The WFC on-axis effective area for each of the science filters
increases the scattering wings become more pronounced. At 277eV and 40.8eV the PSF are 2.9 arcmin (half power width) and 1.7 arcmin respectively, degrading to 4.4 arcmin at 2.5° off-axis.

Data analysis is performed in a similar manner to the PSPC. Sorting software is used to select source areas and background areas. The background is corrected to the vignetting at the source and then after normalisation to the same area subtracted from the source counts. The source is then exposure and vignetting corrected.

In January 1991, the ROSAT satellite attitude control failed and pointing was uncontrolled for a period of time. This caused the loss of one of the PSPCs (detector C) which was replaced by the second detector and observations have proceeded with this. However the WFC was more badly effected, and suffered a serious degradation in efficiency (currently at around 0.18%). This efficiency change has to be now included in the corrections applied to the data.

### A.3 Hubble Space Telescope

The Hubble Space Telescope (HST) (Madau 1995) was launched by the Space Shuttle Discovery in April 1990. Fig A.7 is a schematic of the telescope showing its major components. The HST's scientific instruments (SIs) are mounted in bays behind the primary mirror. The Wide Field Planetary Camera 2 (WFPC2) occupies one of the radial bays, with an attached 45° pickoff mirror that allows it to receive the on-axis beam. Three SIs (Faint Object Spectrograph (FOS), Faint Object Camera (FOC) and the Goddard High Resolution Spectrograph (GHRS)) are mounted in axial bays and receive images several arcmins off-axis.

During the servicing mission in December 1993, the astronauts installed the Corrective Optics Space Telescope Axial Replacement (COSTAR) in the fourth axial bay (in place of the High Speed Photometer). COSTAR deployed corrective reflecting optics in the optical paths in front of the FOS, FOC and GHRS, thus removing the effect of the primary mirror's spherical aberration. In addition the Wide Field and Planetary Camera (WF/PC) was
Figure A.7: The HST. Major components are labeled and definitions of V1, V2, V3 spacecraft axes are indicated.
replaced by WFPC2, which contains internal optics to correct the spherical aberration.

The fine guidance sensors (FGSs) occupy the other three radial bays and receive light 10-14 arcmin off-axis. Since at most two FGSs are required to guide the telescope, it is possible to conduct astrometric observations with the third FGS. Their performance is unaffected by COSTAR.

The HST receives electrical power from two solar arrays, which are turned so that the panels face the incident sunlight. Nickel-hydrogen batteries provide power during the orbital night. The two high gain antennas provide communications with the ground. The support systems module, which encircles the primary mirror, carries out the power, control and communications functions.

The HST is in a low orbit, whose nominal parameters are shown in table A.4.

Optical Characteristics

Because the primary mirror has about one-half wave of spherical aberration, the Optical Telescope Assembly (OTA) did not achieve its design performance until after the December 1993 servicing mission. Table A.5 shows the summary of the optical performance that is now achieved.

Table A.4: The HST nominal orbital parameters (epoch 1994.2).

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Semi-major axis</td>
<td>6972 km</td>
</tr>
<tr>
<td>Altitude</td>
<td>601 km</td>
</tr>
<tr>
<td>Rate of descent</td>
<td>1.8 km yr$^{-1}$</td>
</tr>
<tr>
<td>Inclination</td>
<td>28.5°</td>
</tr>
<tr>
<td>Nodal period</td>
<td>96.4 min</td>
</tr>
<tr>
<td>Orbital precession period</td>
<td>56.1 days</td>
</tr>
</tbody>
</table>
Table A.5: HST optical characteristics and performance.

<table>
<thead>
<tr>
<th>Characteristic</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Aperture</td>
<td>2.4 m</td>
</tr>
<tr>
<td>Wavelength Coverage</td>
<td>1100 Å to 1 mm</td>
</tr>
<tr>
<td>Focal ratio (without COSTAR)</td>
<td>f/24</td>
</tr>
<tr>
<td>Plate Scale (on-axis, without COSTAR)</td>
<td>3.58 arcsec mm⁻¹</td>
</tr>
<tr>
<td>FWHM of WFPC2 images (at 6328 Å)</td>
<td>0.053 arcsec</td>
</tr>
<tr>
<td>FWHM of FOC images (at 4860 Å)</td>
<td>0.042 arcsec</td>
</tr>
</tbody>
</table>

A.3.1 Instruments

All of the SIs are permanently mounted at the HST focal plane, so that all except the WFPC2 receive light that is slightly off-axis. The following are a list of scientific instruments available on the HST:

- **Wide Field Planetary Camera 2 (WFPC2).** The WFPC2 is designed to provide digital images over a wide field of view. It has three 'wide-field' CCDs and one high resolution CCD that are sensitive from 1200-11,000 Å. All four CCDs are exposed simultaneously. A variety of filters may be inserted into the optical path, including a polarizer for polarimetry.

- **Faint Object Camera (FOC).** The FOC is intended to provide high-resolution images of small fields. The camera reimages the focal plane to provide two different scales. After the installation of COSTAR the focal ratios are f/151 and f/75. A variety of filters, prisms and polarizers may be placed in the optical beam.

- **Faint Object Spectrograph (FOS).** The FOS performs low resolution spectroscopy (R ≈ 250 and 1300) of faint source in the wavelength range 1150-8500 Å. A variety of different apertures of different sizes and shapes are available to optimize the throughput and spatial or spectral resolution. Linear and circular spectropolarimetry are available longward of 1650 Å with gratings G190H and G270H. In the R = 250 mode the dispersers are two gratings and a prism, while six grating are available for
the $R = 1300$ mode to cover the full wavelength range. The detectors are two 512-
element Digicons, one operating 1150-5500 Å (blue), and the other from 1620-8500 Å (red). The FOS can acquire data in the accumulation, rapid-readout and periodic modes. Time resolution as short as 30 ms can be achieved. The electron image may be magnetically stepped through a programmed pattern during observations, in order to provide oversampling, compensation of sensitivity variations along the Digicon array, sky measurements, and/or measurement of orthogonally polarized images.

- Goddard High Resolution Spectrograph (GHRS). The GHRS is used for spectroscopy at resolving powers of $R = 2,000$ (1100-1900 Å), 25,000 (1150-3200 Å), and 80,000 (1150-3200 Å). There are four gratings used at medium resolution and echelles combined with cross dispersers for the high resolution modes, as well as one low resolution grating for the 1100-1900 Å range. The detectors are 500-element Digicons which are sensitive from 1100-1900 Å and 1150-3300 Å. The wavelength coverage per exposure ranges from 9 Å at the highest resolution to 48 Å at medium resolution and 285 Å at the lowest resolution. Data is recorded either in accumulation or rapid-readout mode.

The data used in this thesis was from the FOS detector. Analysis methods relevant to the FOS are detailed in Chapter 6.
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