The Unusual Central Star of the Planetary Nebula Sh2-71

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Abstract

This thesis presents new photometric and spectroscopic observational results for the unusual central star of bipolar planetary nebula Sh2-71. The combined lightcurve, composed from the photometric datasets of three different telescopes, was in agreement with the reported ephemeris of the sinusoidal brightness variations with a period of 68 days. The two sharp brightness dips, indicated by the preliminary automated data reduction process, were confirmed. The presence of three additional dips tentatively suggested that the dips, possibly eclipses, are occurring periodically with a period of 17.2 days. The comparison between U and V lightcurves revealed that the 68 day brightness variations are accompanied by a variable reddening effect.

Spectroscopic observations revealed pronounced spectral variations, which were not correlated with the 68 days brightness phase. On the other hand, the high-cadence echelle spectra did not exhibit any variability on hourly timescales, which implied that the spectral variations must occur on timescales of a few days. Radial velocity measurements suggested an amplitude of ±40 km/s but were not correlated with the brightness phase. The measured average radial velocity of the observed star ~26 km/s was in near agreement with the reported mean radial velocity of the planetary nebula. As some doubt has been raised recently that the central star could be another field star, this near agreement between the radial velocities provided supporting evidence that the observed star actually is the central star of the planetary nebula. The comparison between the measured and synthetic spectra yielded stellar atmospheric parameters $T_{\text{eff}}$ 12000 K, log($g$) 4.0 cm/s$^2$, $v_{\text{rot}} \sin(i)$ 200 km/s with an indicated high value of metallicity. Fitted stellar parameters and the comparison with standard spectra classified the star as B8V.

The obtained spectrophotometric observations have been used to construct a model for the central star. A previously suggested cataclysmic binary model has been revisited. The required <1 day orbital period for the mass transfer to establish should be reflected in pronounced spectral profile and radial velocity variations on similarly short timescales. Instead, the high resolution 30 minutes cadence echelle spectra did not exhibit any variations in the timespan of 4.5 hours and thus rejected the cataclysmic model. From the various considered potential models, the spectrophotometric properties of the observed star were best reproduced with a precessing Be disc in a misaligned close binary model. This model could also provide the required collimation for the resulting bipolar shape of the planetary nebula. However, due to the lack of spectra with Hα and Hβ wavelength coverage with a daily cadence, the proposed model should be regarded as tentative.

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The Unusual Central Star of the Planetary Nebula Sh2-71
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The Author

The author gained his first research experience in an astronomical research project titled Measuring the Latitude by Observing the Duration of Sunset, which was carried out during the third year of high school in 2006/2007. The project was presented in a Slovenian nationwide research competition and won the first prize. In 2007/2008, he presented his own measurements of Venus’ parameters in a research project Venus’ Identity Card, which was again placed first on a nationwide competition. The results of both projects the author also published as two journal articles in a Slovenian astronomical journal Spika.

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1 Introduction

The aim of this thesis is to present newly obtained observational properties of the unusual central star of the bipolar planetary nebula (PN) Sh2-71. The provided discussion on the possible progenitors of the observed properties could help astronomers to better understand the origin of aspherical planetary nebulae (PNe).

This chapter presents basic information on PNe, properties of their central stars and possible reasons for aspherical shapes of the vast proportion of observed PNe. A general introduction to PNe is followed by a detailed example of the aspherical PN Sh2-71.

A planetary nebula is an expanding circumstellar cloud of ionized gas, which results from extensive stellar winds during the asymptotic giant branch phase of low to intermediate mass stars. The light from the central star ionizes the gas, which emits recombination and collisionally excited spectral lines (Kwok 2000). PN is a short phase in a star's lifetime lasting only a few tens or hundreds of thousands of years (Gurzadyan 1997). An example of a planetary nebula is shown in Fig. 1.1.

Figure 1.1: Planetary nebula M27. Ljubljana, Observatory Golovec, telescope's mirror diameter 70cm, f/8.3. 2×60s R, 2×120s V, 2×180s B. Image taken 22nd June 2011. Imaging and processing by Teo Močnik.
PNe have a variety of shapes, spanning from spherical through elliptical and butterfly-like shapes to point symmetrical. Binary central stars, magnetic fields and high rotation rates may all play some role in explaining the observed shapes of PNe (Nordhaus et al. 2006). However, mechanisms responsible for such a variety of shapes are still not well understood.

The second part of the Introduction is dedicated to the bipolar planetary nebula Sh2-71. The peculiar variable lightcurve and uncommon spectral variations of the nebula’s central star have attracted astronomers’ attention for decades in trying to understand the physics responsible for the observed unusual properties. In addition, Sh2-71 is one of very few PNe with a confirmed binary central star (De Marco 2006), and it might even be possible that the central star of Sh2-71 is an eclipsing binary. Therefore, Sh2-71 is an intriguing object, which can help us to better understand the shaping mechanisms in general.

1.1 Formation of Planetary Nebulae

Planetary nebulae are formed at the end stages of a star’s life. A typical solar mass star is in thermodynamic equilibrium for billions of years. Being in thermodynamic equilibrium means that H fusion into He in the stellar core is stable, and during this phase, the star is placed on the main sequence of the Hertzsprung-Russell (H-R) diagram (Choudhuri 2010). Eventually, H becomes exhausted in the stellar core and therefore the H fusion ceases in the core. The resulting inert He core contracts due to its own gravity and consequently heats up (Kippenhahn et al. 1991). The increased temperature of the central regions leads to higher reaction rates of H fusion in the H-burning shell around the core, causing the luminosity to increase by a several orders of magnitude. On the other hand, the increased temperature outside the He core produces an excess pressure in the H envelope, which then expands and therefore decreases the star’s effective temperature (Böhm-Vitense 1993). Because the luminosity increases and the effective temperature drops, the star moves up and right from the main sequence of the H-R diagram. Over time, the star becomes a red giant – a large, luminous star with an extensive stellar wind, which can carry away up to $10^8 \, M_\odot$ per year (Groenewegen 2012).

Since the core contracts, both, the density and temperature rise. When the He core reaches the temperature $\sim 10^8 \, K$ and mass $\sim 0.45 \, M_\odot$, the fusion of He into C begins (Kippenhahn et al. 1991). After the core runs out of the He fuel, it begins to contract
again and fusion in the surrounding shells pushes the outer layers even further out. At that stage, the star is in a so-called asymptotic giant branch stage (AGB) (Habing et al. 2003), which typically lasts for a few million years (Vassiliadis et al. 1993). The energy output of the AGB star is provided by the He-burning shell, which lies immediately above the C-O core, and the H-burning shell, which is separated from the He-burning shell by an inert He layer of variable thickness. Since the H-shell supplies He for the He-burning and because the increase in He-shell luminosity leads to near-simultaneous decline in H-burning luminosity, it seems as though the star is pulsating (Habing et al. 2003). The thermal pulsation due to the oscillating H- and He-shell burning increases the mass loss up to values of $10^{-4} \ M_\odot$/year (Zijlstra 2006) – also called the superwind phase (Renzini 1981).

This superwind quickly depletes the stellar envelope. When the mass of the envelope drops to $10^{-3} – 10^{-4} \ M_\odot$, the mass loss rate drops to $\sim 10^{-8} \ M_\odot$/year and the wind speed increases (De Marco 2009). At the same time the effective temperature of a small degenerated stellar core in the centre rises to about 100,000 K (Kovetz et al. 1981). We call this hot stellar remnant a white dwarf (WD).

The mass loss expansion velocities due to stellar wind and thermal pulsation are $\sim 10$ km/s. However, the speed of the fast wind from the hot central star is of the order of 1,000 km/s (Kahn 1983). According to the interacting stellar winds model, the fast ionized wind ploughs the slow neutral wind released during the AGB stage into a specific pattern, which depicts the distribution of the mass loss (Kwok et al. 1978) (see Fig. 1.2). At the same time, because the core of the central star becomes exposed, the UV flux increases, which causes the resulting circumstellar gases to become ionized and then fluoresce (De Marco 2009). At this stage, the object becomes a PN.

Typical optical spectral lines of PNe are H$\alpha$ (6563 Å – red), [OIII] (4959 and 5007 Å – green), [NII] (6548 and 6584 Å – red), H$\beta$ (4861 Å – green/blue), [SII] (6719 and 6730 Å – red) and HeII (4686 Å – blue) (NIST Atomic Spectra Database), each of which adds to the specific colour structure of a PN.
Figure 1.2: Cat's Eye Nebula. The image illustrates the interactions between slow and fast stellar wind. 1: Slow wind, $10^7 \text{M}_\odot$/year. 2: Slow wind, $10^5 \text{M}_\odot$/year. The onion-like structure might result from a disruptive binary mergers or the sudden emergence of a magnetic field (Balick et al. 2001). 3: Result of fast wind sweeping up slow wind. A spiral-like structure probably results from the precession of the central star in the binary system (Miranda et al. 1992). Adapted from Megeath (2011). Credit: NASA, ESA, HEIC and The Hubble Heritage Team (STScI/AURA).

1.2 Central stars

It has been suggested that PNe result from stars with initial masses between 0.8 and 8 $\text{M}_\odot$ (Maciel et al. 2009). The lower limit for initial stellar mass arises because a main sequence star with mass lower than 0.8 $\text{M}_\odot$ would not have had time to evolve beyond the main sequence in the lifetime of the Universe (Adams et al. 1997). On the other hand, the final mass of the central star must not exceed the Chandrasekhar limit of 1.44 $\text{M}_\odot$, which sets the initial mass upper limit roughly to 8 $\text{M}_\odot$ (Koester et al. 1996, also see equation 2). If the final central star’s mass exceeds the Chandrasekhar limit, it explodes as a supernova. The discoveries of PNe with total masses of several $\text{M}_\odot$ indicate that their central stars have had to suffer substantial mass losses. One of the latest robust relationships between the initial ($M_i$) and final ($M_f$) stellar mass was presented by Catalán et al. (2008):
\[ M_f = (0.096 \pm 0.005) M_i + (0.429 \pm 0.015) \]  

if \( M_i < 2.7 \, M_\odot \), whereas for \( M_i > 2.7 \, M_\odot \) the relationship becomes:

\[ M_f = (0.137 \pm 0.007) M_i + (0.318 \pm 0.018) . \]

The reported mean stellar mass value for the central star of a PN is 0.58 \( M_\odot \) (Schönberner 1981), which coincides with the mean stellar mass of DA white dwarfs (Koester et al. 1979), i.e. white dwarfs with hydrogen dominated atmospheres.

The physical and chemical properties of the PN and its central star depend on the initial mass of the progenitor star. A different initial mass defines different chemical composition of the surrounding expelled material as well as the composition (Iben et al. 1978), temperature and the cooling rate of the central star (Blöcker 1995). Fig. 1.3 shows the evolution of the luminosity-temperature dependence for white dwarfs with different masses.

![Figure 1.3: Luminosity-temperature evolution tracks for post-AGB stars. Tracks correspond to WD masses of 0.836, 0.625 and 0.605 M_\odot. Time marks are in units of 10^3 years. Credit: Blöcker (1995), page 761.](image)

We can see in Fig. 1.3 that central star luminosities are likely to exceed \( 10^3 \, L_\odot \), whereas temperatures can reach several \( 10^5 \, K \). However, even if the WD initial luminosities can be several orders of magnitude higher than those of the main sequence stars, it is
challenging to detect a WD spectral flux contribution in the optical part of the spectrum if a WD forms a binary system with a hot main sequence star (Smalley 1997). According to Planck’s law, the vast majority of the WD luminosity is emitted in the UV part of the spectrum. High initial temperatures of WDs are responsible for emission of highly energetic photons and the consequent ionisation of the circumstellar material.

1.3 Types of Planetary Nebulae

We can see from Figs. 1.1, 1.2, 1.7, 1.10 and 4.7 that PNe have a variety of shapes, which allow us to distinguish them by their appearance. Various classification schemes were introduced by different authors, for example: Hromov et al. (1968), Peimbert (1978), Balick (1987), Schwarz et al. (1992) and others. Some classification schemes distinguished PNe according to their degree of symmetry, others to chemical composition or different morphologies. Schwarz et al. (1992) introduced the morphological classification of PNe shapes in narrowband Hα images. They grouped PNe into five classes:

- elliptical: including all spherical and elliptical PNe. (Originating from a single central star with fast stellar rotation or strong magnetic field.)
- bipolar: elongated, axially-symmetric PNe defined by having an equatorial belt from which two faint, extended lobes extend in both polar directions. (Originating from a binary central star or possibly from a single star with fast stellar rotation.)
- point symmetrical: morphological components show point symmetry about the centre. (Originating from a binary central star.)
- irregular: irregular asymmetric shapes, which do not fall into the above classes. (Origin cannot be determined.)
- stellar: PNe which are too far to be resolved and thus have a stellar appearance in the CCD images. (Origin cannot be determined.)

Fig. 1.4 presents the above morphological classes illustratively.
1.4 Planetary Nebulae with Binary Central Stars

The theory of central star binarity arises to explain the observed aspherical and, in many cases, axisymmetric shapes of PNe. The most probable reason for any asphericity is a non-uniform mass loss in the AGB phase. The shapes of bipolar or elliptical PNe intuitively suggest that the slow stellar wind or other mass loss mechanisms produced an envelope with higher density in equatorial regions and lower in poloidal regions. The following fast wind then penetrates the poloidal regions more easily than the equatorial regions, creating an aspherical shape. The possible reasons for aspherical mass loss are binary star gravitation interactions, strong magnetic fields and fast stellar rotation (Corradi et al. 1995). Magnetic fields and stellar rotation are weak collimating mechanisms capable of producing an elliptical PN (Corradi et al. 1995). Bipolar shapes could arise only in most extreme rapid rotation rates (Garcia-Segura et al. 1999) or in the presence of strong magnetic fields of the order of $10^4$ G, which could only be sustained in binary systems, anyway (Nordhaus et al. 2006). On the contrary, binary central stars can induce gravitational interactions, which are a strong collimating mechanism, and thus capable of producing bipolar shaped PNe (Corradi et al. 1995). Therefore, it is crucial to consider that some PNe central stars are likely to be binary stars.

In 2006 Lada presented that about one third of stellar systems are binary or even multiple star systems. A binary star is a system of two stars, which are so close together that they are gravitationally bound. As a consequence, such stars are orbiting around their barycentre. Depending on the distance between the two stars, binary stars are divided into three classes (Kopal 1955):
• detached: the distance between each star and the inner Lagrangian point (L₁) of the Roche lobe is greater than the stellar radius (see Fig. 1.5).

• semidetached: the distance between one of the stars and L₁ point of the Roche lobe is the same as the stellar radius. Therefore mass transfer establishes, flowing from the star which fills its Roche lobe (donor) to the star which does not (acceptor). The inflowing matter usually forms an accretion disk around the acceptor. If the mass transfer rate is higher than the accretion rate, the accreting disk can overflow the Roche lobe through L₂.

• contact: both stellar radiiuses reach L₁ and so they both fill their Roche lobes.

Figure 1.5: Equipotentials in the orbital plane of a binary system consisting of point masses in circular orbits. The mass ratio is $M_1/M_2 = 4$. Credit: Iben et al. (1993), page 1374.

Binary systems in which the separation between the stars is smaller than ~100 AU are called closer binary systems (De Marco 2009). If the separation between the stars is so small that at least one star has filled or will fill its Roche lobe, then such a system is called a close binary system (Iben et al. 1993). Close and closer binary systems have smaller periods and greater gravitational interactions than widely separated binaries whose orbital periods can be even longer than the lifetime of the PN. Therefore, binaries with smaller separations have a greater influence on the shape of the outflowing matter and, eventually, on the shape of a PN (De Marco 2009).
Another common distinction between binaries is made by the discovery method (De Marco 2006):

- visual binaries: the angular separation of the stars is greater than the telescope's angular resolution, which allows us to see the visual binary as a double star.
- astrometric: the sequence of position measurements of one star reveals that it orbits around a barycentre and one can conclude such a star is a part of a binary system.
- spectroscopic: the sequence of radial velocity measurements in the stellar spectra reveals periodic red and blue Doppler shifts. The other possibility is to fit the system's spectra with the composition of two individual spectra corresponding to spectral classes of the individual stars.
- photometric: the sequence of photometric measurements reveals periodicity in stellar magnitude. The drops in magnitude can correspond to a certain system configuration or actual eclipses of the two stars. It is obvious that photometric eclipses can be applied only if the inclination of the system is close enough to 90° (edge on).

See Fig. 1.9 for statistics on the discovery methods of PNe binary central stars.

White dwarfs are not the only possible ionizing source of PNe. In close binary systems, it is possible that due to gravitational interactions a red giant will lose its hydrogen envelope before the helium ignites in the core (Jeffery 2005). Such stars are called subdwarfs, whose luminosities are spanning from few to 100 L⊙ (Napiwotzki 2011) and temperatures between 2·10⁴ and 10⁵ K (Oreiro Rey et al. 2004). Such parameters justify that hot subdwarfs can ionize the circumstellar material in the same way as WDs (Smalley 1997). Therefore, if the central star of the PN is identified to be a close binary star, it is important to consider the possibility that the ionizing source is a hot subdwarf instead of a WD.

From the present samples of observed morphologies of PNe it appears that the pronounced aspherical shapes of PNe with a binary central star result from a common envelope around a close binary central star (Zijlstra 2007). A common envelope is a short lived phase of the binary stellar evolution in which a donor giant star overflows L₁ while it is expanding. Consequently, the mass transfer accelerates and the binary orbit shrinks. As a result, the timescale for mass transfer is considerably shorter than the timescale in which the acceptor can adjust thermodynamically to the inflowing matter.
The accreting matter will heat up, expand and fill the Roche lobe of the acceptor. The matter originating from the donor then flows into a common envelope encompassing both stars. Such a system is visually similar to a contact binary, but unlike a contact binary system, the secondary star in a common envelope experiences drag forces and spirals in (Iben et al. 1993). If the common envelope then overflows L₂, it creates a dense equatorial disk around the binary system. And that disk leads to a bipolar PN (Cuesta et al. 1993). Nordhaus and Blackman (2006) showed that even if the common envelope does not overflow L₂, it still gets distributed with a density contrast between higher density in equatorial and lower in poloidal regions. The density contrast allows the PN to have a bipolar shape. Three collimating scenarios are possible for common envelopes (Nordhaus et al. 2006):

- **Mutual drag transfers** the angular momentum from the secondary star to the common envelope. Consequently, the secondary spirals in and the common envelope is spun up. The secondary provides enough kinetic energy to unbind the envelope by itself. The envelope is ejected in the equatorial plane. Due to the higher density in equatorial regions, the resulting PN is elongated in poloidal regions (see Fig.1.6a).

- **The secondary spirals in,** the envelope is spun up and begins to rotate differentially. When the differential rotation is combined with the presence of a deep convective zone, a dynamo is generated in the envelope. The dynamo mechanism is similar to that studied in the Sun, but here the magnetic fields are considerably stronger. Magnetic fields drive the envelope's poloidal material in poloidal directions. As a consequence, the density of poloidal regions becomes smaller compared to equatorial regions (Fig. 1.6b).

- **The secondary spirals in.** Because the secondary does not provide enough kinetic energy to eject the common envelope, it eventually gets too close to the core of the primary star. The tidal forces shred the secondary into an accretion disc around the core, which drives an outflow in poloidal directions (Fig. 1.6c).

Whichever the collimating scenario, the resulting bipolar PN will always be elongated in poloidal directions. An example of a bipolar PN is the Butterfly Nebula or NGC 6302 (see Fig. 1.7).
Figure 1.6: Collimating scenarios of a binary central star with a common envelope. a – the secondary star ejects the envelope by itself; b – the secondary star induces large scale magnetic fields, which eject the poloidal regions of the common envelope in poloidal directions; c – the secondary star is shredded in a disk, which collimates the outflow. Credit: Nordhaus et al. (2006), page 2008.

Figure 1.7: Butterfly Nebula. This PN is a prototypical bipolar PN, also called a butterfly PN. The poloidal lobes are collimated by the dense equatorial disk of gas and dust. Credit: NASA, ESA and the Hubble SM4 ERO Team.
Point symmetric PNe are also believed to originate from binary central stars. Several researchers have proposed that point symmetric PNe result from periodic ejection of a precessing two sided jet from one of the stars in the binary system (Livio et al. 1996).

Different fractions of central star binarity have been reported. The most robust measurement was carried out by Miszalski et al. in 2009 who reported a PNe binary fraction of 12–21%. On the other hand, the fraction of bipolar morphologies is roughly 15% (Frew 2008) and the fraction of point symmetric PNe 4% (Livio et al. 1996). Therefore, we may argue that statistically binary central stars could be the progenitors of bipolar and point symmetric PNe (Miszalski et al. 2009).

1.5 Planetary Nebulae with Fast Rotating Central Stars

Fast stellar rotations could also produce bipolar or elliptical PNe. Garcia-Segura et al. (1999) proposed that stars above 1.3 $M_\odot$ can achieve near critical rotation during their superwind phase at the tip of the AGB. They showed that the equatorially confined winds can produce bipolar or elliptical PNe. Which of the two types will form depends on the central stellar mass. Corradi et al. (1995) analysed 400 high quality optical images and tentatively suggested that initial masses of elliptical PNe central stars are expected to be less than 1.1 $M_\odot$, while bipolar progenitor stars should have masses above 1.5 $M_\odot$. However, their strongest claim was that bipolar PNe are formed from the most massive stars of all PNe.

Another stellar rotation mechanism for collimating the PNe was introduced by Georgiev et al. (2011). They suggest that stellar rotation could divert spherical flux into poloidal collimated flux due to gravitational brightening. The authors speculate that the core of an AGB star is contracted at the end of the thermonuclear reactions. This stellar remnant partially conserves its angular momentum and increases its rotational velocity by about 2 orders of magnitude. They claim that the rotation velocity is therefore sufficient to differentiate poloidal from equatorial flux. As the flux is reduced in the equatorial regions, so is the ionization rate and thus the resulting PN is bipolar.
1.6 Statistics on Observed Planetary Nebulae

The expected number of PNe in the Galaxy is \( \sim 46,000 \) (Moe et al. 2006). However, only about 3,000 Galactic PNe have been discovered (Weidmann et al. 2011). Fig. 1.8 shows the distribution of morphological types among the discovered PNe. The number of applied observation techniques for discovered closer binary central stars is presented in Fig. 1.9.

Figure 1.8: Proportions of PNe morphological types. The PNe are classified as round if the difference between major and minor semiaxes is less than 5%. The diagram does not include stellar like PNe because these are just unresolved PNe of the remaining morphological classes (Frew 2008, Livio et al. 1996, Parker et al. 2006, The MASH Catalogue of Planetary Nebulae).

Figure 1.9: PNe with closer binary central stars – number of applied observation techniques. As of April 2009, the total number of confirmed closer binary central stars was 42. The number of discovered binaries is smaller than the sum of specific observation techniques because some binaries have been confirmed with multiple techniques (De Marco 2009).
1.7 Sh2-71

1.7.1 Discovery and Basic Data

PN Sh2-71 was discovered by Minkowski in 1946 who classified it as a diffuse and peculiar nebulosity. In 1959, Sharpless included this object in the second and final version of the Sharpless catalogue of emission nebulae. At that time, the object was classified as a possible PN. With the aid of the development of imaging, modern catalogues unequivocally classified the object as a PN (Catalogue of Galactic Planetary Nebulae, Kohoutek 2001). Sh2-71 is shown in Fig. 1.10.

Figure 1.10: Planetary nebula Sh2-71. Crete, Skinakas Observatory, telescope's mirror diameter 129cm, f/7.7. 5×90s R, 6×900s Hα, [SII] and [OIII]. Image taken 27th–29th July 2010. Imaging and processing by Makis Palaiologou, Stefan Binnewies and Josef Pöpsel. The use of this image was kindly permitted by Mr Josef Pöpsel.

The basic parameters of Sh2-71 and its central star are listed in Table 1.1. Due to inaccurate distance estimations (± 30%) it is difficult to determine distance-dependant variables, like the age of the PN or the luminosity of the central star. Therefore, some published parameters have higher uncertainties.

<table>
<thead>
<tr>
<th>planetary nebula</th>
<th>name of the PN</th>
<th>Sh2-71, PN G035.9-01.1</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>RA, dec [ICRS J2000.0]</td>
<td>19:01:59.3 +2:09:18</td>
</tr>
<tr>
<td></td>
<td>constellation</td>
<td>Eagle</td>
</tr>
<tr>
<td></td>
<td>dimensions [Δα × Δδ]</td>
<td>1.7' × 3'</td>
</tr>
<tr>
<td></td>
<td>type</td>
<td>bipolar</td>
</tr>
<tr>
<td></td>
<td>distance</td>
<td>1,006 pc</td>
</tr>
<tr>
<td>central star</td>
<td>name of the central star</td>
<td>V1710 Aql</td>
</tr>
<tr>
<td></td>
<td>RA, dec [ICRS J2000.0]</td>
<td>19:02:00.3 +2:09:11</td>
</tr>
<tr>
<td></td>
<td>type</td>
<td>binary (A7V-F0V &amp; WD or sdO)</td>
</tr>
<tr>
<td></td>
<td>average apparent magnitude</td>
<td>U 14.97</td>
</tr>
<tr>
<td></td>
<td></td>
<td>B 14.59</td>
</tr>
<tr>
<td></td>
<td></td>
<td>V 13.75</td>
</tr>
</tbody>
</table>

#### 1.7.2 Observations and Current Explanations of the Observed Properties

The first extensive photometric observations of Sh2-71's central star were performed by Kohoutek between 1977 and 1979. From 263 photometric measurements in U, B and V filters, he discovered flux variations and reported the variability of more than 0.7 mag with indications of a smooth sine-like lightcurve. He used the central star's colour indices to initially classify the central star as a B8V. Kohoutek also used the central star's spectra obtained by Glushkov et al. in 1975 and noted that the central star of spectral type B8V cannot be responsible for the observed high excitation in the surrounding PN. Therefore, he suggested that the central star is a binary system in which the secondary star is hot enough to ionize the nebula. However, Cuesta et al. (1993) identified with the aid of the composite spectra that the spectral type of the central star is likely to be A7V, which in fact even strengthened the assumption of a hot binary companion.

The suggestion of a binary with a hot ionising secondary star was further supported by Feibelman in 1999 who processed all five available UV spectra taken by IUE satellite and discovered a flux variability in the MgII emission line at 2800 Å. Referring to the variability similarities with the planetary nebula LoTr5, he tentatively attributed the variable MgII line to the excitation of the cooler component’s chromosphere by the hot companion.
Preite-Martínez et al. spectroscopically determined in 1989 that the effective temperature of the ionizing source is as high as 129,000 K. The high temperature was later confirmed by Kaler et al. (1990) and Bohigas (2001).

The results from the next set of extensive photometric observations between 1990 and 1992 were published by Jurcsik in 1993. 2,001 photometric measurements in filters U, B, V, RC and IC revealed a 68.06 day period in the lightcurve variations.

The largest photometric study to date was presented by Mikulašek et al. in 2007. They obtained 2,004 photometric measurements in V, RC and IC band, and combined them with existing photometric data from Kohoutek and Jurcsik. From combined 4,268 measurements, spanning over 28 years, they refined the period to (68.101 ± 0.010) days (see individual lightcurves in Fig. 1.11) and presented an ephemeris for timing of the light maxima:

\[ JD_{\text{max}} = (2449862.02 \pm 0.32) + (68.101 \pm 0.010) \cdot n. \] (3)

Beside those periodic light variations, Mikulašek et al. (2007) also noted long term non-periodic variations of the lightcurve shape (see Fig. 1.11).

![Figure 1.11](image-url): Long term non-periodic variations of the central star's lightcurve shape. Credit: Mikulašek et al. (2007), page 470.

Mikulašek et al. (2007) interpreted short-term (68 days) periodic variations as due to the orbital motion of stars in a binary system. As the stars orbit one another, the light from
the brighter central star could experience different optical paths in an optically thick inner nebula. On the other hand, the long-term lightcurve shape variations were attributed to an unstable mass transfer between the stars in a binary system.

Mikulašek et al. (2007) supported the interpretation of the short-term periodic variations by the observed differences in the spectra taken at two different phases of the lightcurve (see Fig. 1.12). They claimed that rapid changes in ionization conditions arise because we are looking at the central star through the inhomogeneous dense inner nebula and the optical path varies due to the orbital motion of stars in a binary system. They also used the intercomparison between different [OIII] spectral lines to confirm that the density of the inner nebula is very high.

![Figure 1.12: Spectra of the core of Sh2-71 at different phases. Credit: Mikulašek et al. (2007), page 471.](image)

In addition to the unusual lightcurve and spectroscopic variations, Sh2-71 also has an unusual outflow, manifested in the kinematic structure. Cuesta et al. (1993) have undertaken slit spectra of Sh2-71 at various position angles to investigate the kinematic structure of the PN (see an example of one position angle in Fig. 1.13a).

Cuesta et al. (1993) noted that the spectral lines experience Doppler shift, which is dependent on relative position, and a rapid 36 km/s switch in velocities when traversing the nucleus (see Fig. 1.13b). Interpolated kinematic structure from slit spectra at various position angles revealed that the orientation of the [NII] outflow kinematic structure as well as the largest velocity shifts were measured in the direction SE-NW (similar to the position angle of the slit in Fig. 1.13a). We can see in Fig. 1.13a that the kinematic structure of the [NII] outflow is not aligned with the overall orientation of the nebula. Besides the unusual [NII] kinematics, we can also see in Fig. 1.13b and 1.13c
Figure 1.13: Slit spectra of Sh2-71. a – position of the slit; b – spectra along the slit. Featured spectral lines are Hα (6563 Å) and [NII] (6548 and 6584 Å). Note that spectral lines are not straight but experience different Doppler shifts along the slit. Rapid velocity switch in the centre can also be noted. c – spectra of the nucleus (same as Fig. 1.13b at relative position 0°). Credit: a – La Palma, Liverpool Telescope, telescope's mirror diameter 2.0m, f/10. 2×10s R. Image taken 11th June 2012. Imaging by Myfanwy Lloyd, Don Pollacco and Rachel Street. Processing by Teo Močnik. b – Cuesta et al. (1993), page 382. c – Cuesta et al. (1993), page 383.

the huge velocity contrast between Hα and [NII] spectral lines. Full width at half maximum (FWHM) of [NII] lines are 30 km/s, and change only slowly with distance from the nucleus. On the other hand, the two Hα emission peaks in Fig. 1.13c are separated by 230 km/s, whereas the basal width of the line in the nucleus is 1,026 km/s and becomes comparable to the width of [NII] only a few arcsec away from the nucleus. At the same distance, [NII] emission begins to dominate over Hα. These spectroscopic measurements imply that Sh2-71 has a slow-expanding disk, which collimates a high velocity central outflow. A fast bipolar outflow could explain the two peaks of Hα emission in Fig. 1.13c (Cuesta et al. 1993).

To support their idea, Cuesta et al. (1993) proposed a model for the progenitor of Sh2-71 (see Fig. 4.1). According to that model, the central star of Sh2-71 is a close binary system of A7V-F0V and an ionizing star, in which the material is transferred
from the main sequence star to the ionizing star, where the matter overflows the outer Lagrangian point \( L_2 \). As a result, the overflown matter forms a slowly expanding equatorial disk, which collimates high velocity outflow from the ionizing star. The shape of the nebula implies that the outflow and also the binary system are likely to be nearly edge-on. Therefore, according to the model presented by Cuesta et al. (1993), the binary system must have a very short (\( \ll 68 \) days) orbital period in order to establish mass overflow and because the system is nearly edge-on this means that we should be able to see eclipses or partial eclipses with periods much shorter than 68 days. Until now, such eclipses have not been reported. Because of the disagreement between the observed lightcurve period and their estimation of orbital period and also because of the lack of data, especially the central star’s spectra, Cuesta et al. (1993) admit that their model should be regarded as tentative.

Miranda et al. (2005) proposed a different view. They interpreted their spectral kinematics analysis and narrow band images as a result of three episodic fast wind eruptions, which were collimated in the immediate vicinity of the central binary instead of an equatorial disk. And as the collimation axis was changing through time, episodic fast winds resulted in three edge-on bipolar lobes (see Fig. 1.14).

![Figure 1.14: Three pairs of bipolar lobes of Sh2-71. Credit: Miranda et al. (2005), page 97.](image)
Several authors of scientific discussion about Sh2-71 have somewhat different interpretations related to the lightcurve and spectroscopic variations, different collimating mechanisms or even proposing a different field star as a progenitor (Frew et al. 2007). The variety of interpretations implies that Sh2-71 is an intriguing and, in many ways, unusual PN that is still far from being well understood.

1.8 Conclusion and Thesis Outline

The science related to PNe has made significant and notable steps forward in the past few years. The progress was made mainly due to the improved computer simulations and even more importantly, vastly improved statistics for PNe central binary stars. From the current knowledge, it appears that close binary central stars are the most effective collimating mechanism that is required for shaping bipolar PNe. However, no general consensus has yet been made for progenitors of aspherical PNe and therefore improving our knowledge of certain representative, or unusual, PNe helps us to better understand PNe in general.

The unusual bipolar PN Sh2-71 has been a target of different surveys for decades. Large photometric datasets enabled astronomers to accurately determine the period for the observed 0.8 magnitude brightness variations and also to construct detailed but unfortunately tentative models for the central star, as discussed in subsection 1.7.2. There are many different explanations for the progenitor of the Sh2-71, some of them even being contradictory to the other. But one opinion seems to be in common to all of the involved researchers: the need for further observations.

The aim of this project is to reduce and analyse new photometric and spectroscopic observations of Sh2-71’s central star. The stimulation for the work was the lightcurve, which was obtained in 2011 by the automated data reduction of photometric measurements from the Byrne Observatory at Sedgwick (BOS). The data revealed two possible eclipses events, which were separated by ~16 days (see Fig. 2.1). In order to confirm the eclipses, two new photometric datasets were obtained with the Faulkes Telescope South (FTS) and Liverpool Telescope (LT). Furthermore, six spectra at different phases of the lightcurve were acquired by the LT. An additional set of 25 archived and unpublished spectra from the Isaac Newton Telescope (INT), taken over a two year period, was used to extend the sample of central star’s spectra. These spectra were used for the verification of the correlation between the spectral variability and star’s 68 day brightness variations. On the other hand, hourly timescales of
spectroscopic variability was checked by considering eight archived echelle spectra with 30 minute cadence obtained by the Anglo-Australian Telescope (AAT). The first objective of this project is to investigate the presence of the eclipses and possibly to model them in order to derive binary parameters. Secondly, the spectroscopic dataset will be used to provide the most accurate verification to date of the correlation between the spectral features and the lightcurve phase. Spectra will also be analysed in order to determine radial velocities and stellar atmospheric parameters. The final goal of the project is to present the most likely model for the progenitor of the Sh2-71.

This thesis proceeds as follows: Chapter 2 presents photometric measurements, data reduction process and final photometric results; Chapter 3 is dedicated to spectroscopic data reduction, results and the correlation of the spectra with the lightcurve phase; Chapter 4 discusses possible models for the progenitor of the planetary nebula based on the observed properties, which are presented in this thesis.
2 Photometry

The first photometric dataset for this study of Sh2-71’s central star was obtained with 0.8 m telescope at Byrne Observatory at Sedgwick (BOS) in California between 26\textsuperscript{th} June 2011 and 28\textsuperscript{th} July 2011. The object was observed at that time, because it was identified in 2004 to possess the most pronounced variability of spectral features among 33 PNe central stars that were included in a survey of central star binary fraction (Sorensen et al. 2004). The observations in B filter, obtained with BOS telescope, were later reduced by an automated data reduction process. The resulting lightcurve of Sh2-71’s central star shows a long term trend of increasing brightness corresponding to the 68 day brightness variations, discovered by Kohoutek in 1979. But in addition, the automated data reduction also indicated the presence of two sharp brightness dips in the lightcurve (see Fig. 2.1). The two dips, possibly eclipses, are separated by ~16 days, whereas the overall extend of the monitored lightcurve is 32 days.

Figure 2.1: Preliminary lightcurve obtained by BOS telescope. The lightcurve indicates two possible eclipses, separated by ~16 days. The overall increase in brightness is a part of the 68 days brightness variations. The lightcurve is the result of an automated data reduction process. Credit: Pollacco et al. (2011), page 3.
The presence of possible eclipses in the lightcurve, obtained by the automated data reduction process, stimulated additional follow-up observations with other telescopes. The primary goal was to confirm the presence of eclipses and possibly to model the eclipses in order to obtain the basic binary parameters. In addition, photometric measurements were taken simultaneously with the spectra, so that the spectra could be accurately compared to the actual phase and current shape of the lightcurve. Simultaneous photometry could also verify whether the eclipsing events are causing any spectral variations.

The first additional photometric dataset in B and V filters was obtained with 2 m Faulkes Telescope South (FTS) in Australia. The second additional and also the largest dataset was supplied by the FTS’s twin telescope, 2 m Liverpool Telescope (LT) on La Palma, Canary Islands. The latter observations were performed in U and V filters. Table 2.1 presents an overview of the basic observational details for the three telescopes.

<table>
<thead>
<tr>
<th>telescope</th>
<th>duration of observations</th>
<th>filter</th>
<th>exposure time</th>
<th>S/N</th>
<th>number of data points</th>
</tr>
</thead>
<tbody>
<tr>
<td>BOS (0.8m)</td>
<td>26th Jun 2011 – 28th Jul 2011</td>
<td>B</td>
<td>600s</td>
<td>543</td>
<td>21</td>
</tr>
<tr>
<td>FTS (2m)</td>
<td>4th Oct 2011 – 23rd Mar 2012</td>
<td>B</td>
<td>15s</td>
<td>92</td>
<td>7</td>
</tr>
<tr>
<td></td>
<td></td>
<td>V</td>
<td>15s</td>
<td>157</td>
<td>7</td>
</tr>
<tr>
<td>LT (2m)</td>
<td>31st Mar 2012 – 20th Jul 2012</td>
<td>U</td>
<td>3×24s</td>
<td>64</td>
<td>34</td>
</tr>
<tr>
<td></td>
<td></td>
<td>V</td>
<td>3×3s</td>
<td>171</td>
<td>36</td>
</tr>
</tbody>
</table>

Table 2.1: Basic observational details for photometry. The table lists the names of the telescopes and the primary mirror’s diameters, the dates of the first and last observation, applied filters, exposure times, mean signal to noise ratios for the central star and the final number of useful photometric data points.

Note the differences in the signal to noise ratios in Table 2.1, which result mainly due to different exposure times and different selections of the filters. Signal to noise ratio was calculated for each date individually using equation (4) (Howell 1989):

$$\frac{S}{N} = \frac{N_s}{\sqrt{N_s + n_{pix}(N_{sky} + N_{dark} + N_{RON})}},$$  \hspace{1cm} (4)$$

where $N_s$ denotes the number of electron counts in the photometric aperture originating from the source alone, $n_{pix}$ is the number of pixels in the aperture, $N_{sky}$ is the mean number of counts per pixel in the annulus, $N_{dark}$ number of counts per pixel due to dark
current, and $N_{\text{RON}}$ is the readout noise per pixel. The deviation of the measured signal to noise ratio from the ideal Poisson statistics $N_s/\sqrt{N_r}$ was small and in the worst cases reached 20% for LT when using U filter. Therefore, even for the smallest signal to noise ratios, photometric noise corrections due to sky, dark current and readout noise did not significantly contribute to the noise. Photometry was clearly governed by Poisson statistics for all the photometric observations.

The following subsections present the data reduction process and the final lightcurves for each of the three telescopes. Combined lightcurves and outlined results are presented in subsection 2.4.

## 2.1 Byrne Observatory at Sedgwick

The first photometric dataset for this project was obtained with 0.8 m telescope as a part of Byrne Observatory at Sedgwick (BOS) between 26$^{\text{th}}$ June 2011 and 28$^{\text{th}}$ July 2011. Soon after that, the data were reduced by the automated data reduction process, which yielded a lightcurve, shown in Fig. 2.1. The first objective of this project is to verify the result provided by the automated photometry reduction pipeline. Therefore, I manually re-reduced the BOS data to obtain a more reliable and accurate lightcurve. Photometric and spectroscopic data reductions were performed with several application packages, provided by Starlink.

Images, obtained by all three telescopes, BOS, FTS and LT, had already been pre-processed, i.e. bias and dark subtracted, flat field corrected. The task was therefore only to align the images and perform differential photometry. An example of an image, taken by BOS, is shown in Fig. 2.2.

Before the photometry could be performed, all the images had to be aligned. Images which are aligned in the pixel coordinate frame, enable the user to input the same pixel positions of the stars for all the observation dates. This simplifies and speeds up the photometric reduction process. The BOS pre-processing pipeline supplied the images with accurate equatorial coordinate frames. So, in order to align the BOS images, I only had to align the equatorial coordinate frames. This was done by using a command `wcsalign` as a part of the Starlink package KAPPA.
Figure 2.2: First image of Sh2-71 obtained by BOS on 26th June 2011. The central star of the nebula is labelled with number 2. Stars 1, 3, 4 and 5 are non-variable comparison stars. The star labels are consistent for all three telescopes. A faint nebula is also visible. The field of view is 14.6’ × 9.7’.

Some of the images were taken during the same night and for such cases it is worth considering the possibility of making a combined image in order to increase the signal to noise ratio. Having the images aligned, it is straightforward to combine them. However, because of the relatively long time sequences between the images, and consequent noticeable differences in counts values due to variable airmass, seeing and weather conditions, it was better to perform photometry on each individual image. In addition, individual photometric data points can trace the reliability of the reduction process. Furthermore, the BOS individual images already had by far the highest signal to noise ratios in comparison with any other photometric observation (see Table 2.1).

After all the individual images had been aligned in the pixel coordinate frame, I was able to proceed with the photometric data reduction. For this task, I used the command autophotom provided by the package PHOTOM. The command places the photometric aperture and the sky annulus on the specified coordinates and performs centering. After that, the command sums the data counts in the aperture and subtracts the sky contribution, which is estimated from the 2 sigma clipped mean count value in the annulus. I used circular apertures, because the stars’ eccentricities were very small and the photometric results obtained with elliptical apertures differed negligibly from circular apertures. The aperture radius, position and width of the annulus must also be specified by the user. Fig. 2.3 shows an example of the photometric aperture and
annulus. The resulting photometric measurement is then given as the number of source electron counts with the corresponding error. The alternative is that the data counts are converted into instrumental magnitudes. The advantage of this command is that it enables the user to supply a list of pixel coordinates for all the stars of interest and the resulting photometric measurements are then provided for all the stars from that list. In addition, the command can be run from a c shell script. Compared to completely manual photometry, running `autophotom` command from a c shell script is faster and minimises the probabilities of human errors.

![Photometric aperture and annulus configuration](image)

**Figure 2.3:** Photometric aperture and annulus configuration. The aperture is centred on the Sh2-71’s central star. Aperture radius is equal to 1 averaged stellar full width at half maximum (FWHM), the annulus extend between 2 and 3 FWHM. The image was taken on 26th June 2011.

When considering the radius of the photometric aperture, it is important to include as much stellar signal as possible, but on the other hand minimise the sky contribution by making the aperture smaller. Theoretically the ideal aperture radius with the largest signal to noise equals to 0.68 full width at the half maximum (FWHM) of the averaged stellar point spread function (PSF), i.e. the seeing (Mighell 1999). However, if the centering is inaccurate, bigger errors arise in the case of smaller apertures. Therefore, a recommended practical compromise is to choose a radius of 1 FWHM (Mighell 1999). But because brighter stars generally have larger optimal aperture size than faint stars, the best way to choose an optimal aperture radius is to plot a curve of growth and read the appropriate radius. The curve of growth gives the relation between the sky subtracted data counts or instrumental magnitudes and aperture radius (see Fig. 2.4).
The optimal aperture radius is then the smallest possible radius where the most of the signal is captured in the aperture, i.e. the radius where the rate of increase in the signal for a given increase in aperture radius declines rapidly (Stetson 2013).

Figure 2.4: Curve of growth in number of electron counts for star 2 (black, left vertical axis) and star 3 (red, right axis). The curve provides the relation between the sky-subtracted aperture counts versus aperture radius. Note that the slope of the curves for both stars changes most rapidly near the radius 4.5 pixels. This is the smallest radius where most of the signal is captured in the aperture. The optimal choice of aperture radius maximises the signal to noise ratio. The data points, shown in this and on some following photometric plots, are connected with lines for easier comparison between different stars or other data types.

Figure 2.5: Curve of growth in instrumental magnitudes. Note that both vertical axes have the same scale. Thus, the differences between the two instrumental magnitudes do not depend strongly on the selected aperture radius.
Figs. 2.4 and 2.5 show that optimal aperture radius for image taken on 26th June 2011 is close to 4.5 pixels. Because it is very impractical to read an optimal radius for every observation date from the curve of growth, it is suitable to relate the radius to seeing. For this task, I used a command psf, which fits a list of bright, unsaturated stars with Gaussian functions and computes a mean FWHM in pixels. Fig. 2.6 shows an averaged normalized Gaussian fit for 26th June 2011.

Figure 2.6: Stellar PSF fit for 26th June 2011. The shown PSF is the averaged normalized fit of stars 2, 3, 4 and 5. Star number 1 has been excluded because it was exceeding the linear range of the full well capacity of the CCD. The calculated FWHM was 4.27 pixels, provided by the command psf.

We can see from Fig. 2.6 that for 26th June 2011 the seeing was 4.3 pixels, close to estimated optimal aperture radius from the curve of growth in Fig. 2.4. Therefore, the curve of growth and PSF fit justify the choice of 1 FWHM aperture radius, as recommended by Mighell (1999). Hereinafter, I was able to perform a PSF fit for every observation date and use the computed FWHM to set the aperture radius.

Another important parameter for photometry is to set the position of the annulus. Because we want to subtract the contribution of the sky that affects the region of the source, it is reasonable to place the annulus as close to the source as possible. On the other hand, if the annulus is too close to the source, the estimated sky value becomes contaminated by the source itself. The suggested gap between the aperture and the annulus is 1 FWHM (Munari 2012). We can see from Fig. 2.6 that the contamination of the source becomes negligible at 2 FWHMs (i.e. 8.5 pixels for 26th June 2011). Fig. 2.7
shows a part of the final lightcurve, when the 1 FWHM wide annulus is placed between 2 and 3, 3 and 4 and finally between 4 and 5 FWHM distance from the source.

![Figure 2.7: The impact of the annulus position on the final lightcurve. Vertical axis shows the differences between the instrumental magnitudes of the central star and star number 3. According to the legend, each curve corresponds to a certain position of the annulus, e.g. the black lightcurve corresponds to aperture of radius 1 FWHM, whereas annulus is placed between 2 and 3 FWHM. Four observing nights are shown.](image)

Since Fig. 2.7 proves that the annulus has not been contaminated by the stellar light in any of the attempted annulus positions, I decided to use the annulus, which extends between 2 and 3 FWHM.

However, it is also possible that the central star was contaminated by the light from the planetary nebula. In this case, the differential photometry would have a certain dependence on the aperture radius, i.e. the seeing. The scatter, present on Fig. 2.8, proves that any eventual nebulosity contamination was negligible. If the contamination would have been present, the solution would be to choose an aperture with the same radius for all the observing nights. This would reduce the impact of the nebular contamination on the differential photometry at the expense of the lower signal to noise ratio.
The differential photometry with a seeing-dependant aperture radius shows no correlation with the seeing. Therefore, the nebular contamination of the central star’s photometry is negligible.

Having the images aligned, the photometrical parameters selected and having contamination-free data counts, it is possible to present the final lightcurve of the central star. If the absolute magnitude would have been required, it would be necessary to extrapolate the measured instrumental magnitudes to asymptotic values of the curves of growth (Howell 1989). In addition, the provided dataset did not include any photometric standard stars and additional filters for colour corrections that would normally be required for accurate absolute magnitude measurements. However, the differential magnitudes provide all the information required for the tasks of this research project.

The lightcurve, shown in Fig. 2.9, presents the differences between the instrumental magnitudes measured by autophotom. The black lightcurve shows the magnitude differences between the central star and comparison star number 5. The red lightcurve plots the differential magnitudes with comparison star number 3. The upper panel of Fig. 2.9 shows the differential magnitudes between the two comparison stars. The smaller the variations of the differential magnitudes of the two non-variable comparison stars, the more reliable is the differential photometry of the central star. The horizontal axis is given in Julian date. Note that these data have not been corrected to heliocentric Julian date as the time corrections (≤8 min) are insignificant for the purposes of this project.
Figure 2.9: The final lightcurve of the BOS data. Differential magnitudes in B filter are the differences between the instrumental magnitudes of the central star and star number 5 (black line), or star number 3 (red line). The reliability of the photometry can be estimated from the variations of the differential magnitudes between the two comparison stars (upper panel). The three vertical axes are at the same scale.

The most obvious difference between the manually obtained photometrical results, presented in Fig. 2.9 and automated data reduction pipeline in Fig. 2.1 is the absence of the deepest data point in the second brightness dip. During the manual data reduction it turned out that the image, corresponding to that data point, was actually a different region of the sky, which did not even contain the central star. However, the manually obtained lightcurve still exhibits brightness dips that are significantly more pronounced than the variations of the differential lightcurve of the two comparison stars. Therefore, the lightcurve in Fig. 2.9 tentatively confirms the presence of the two dips, previously seen on the lightcurve of the preliminary results.

The brightest comparison star, star number 1, which could provide the highest signal to noise ratios, was outside of the linear range of full well capacity of the CCD for some of the observing nights. Therefore, the differential photometry with this star was not reliable and was removed from the photometric analysis of the BOS data.
2.2 Faulkes Telescope South

The next photometric dataset was supplied by the 2 m Faulkes Telescope South (FTS) located in Siding Spring in Australia. The telescope observation time was allocated in semester 2011B, which covered 1st Oct 2011 – 31st Mar 2012, immediately after obtaining the preliminary results with the BOS telescope. However, in the middle of this period, the nebula was in its closest position to the Sun, which prevented observing. Therefore, the data were taken only at the beginning and end of the semester 2011B. A total of 13 observing nights supplied 13 images in V and 13 images in B filter.

The initial quality check revealed that during three observing nights the provided images exhibited nothing but noise, probably due to clouds. One observing night supplied a pair of images with a shifted field of view, on which the central star of Sh2-71 was not present. A further number reduction of useful data points occurred due to a blot that was present on two pairs of images (see Fig. 2.10). The FTS support staff explained that the blot sometimes appeared because of unpredicted heating of the CCD and consequent condensation of the CCD’s window. The light was refracted on the ring of condensation and resulted in features similar to Fig. 2.10. The only way of overcoming this problem would be to apply a proper flat field correction. But because the blot appeared randomly and because the shapes of the blot were not always the same, the flat field would have to be taken very often, preferably after every image. Since additional flat fields were not provided, the blots could not be removed. Photometry, performed on affected images, clearly demonstrated that the photometric results were corrupted by the blots, since the differential magnitudes were deviating from expected values by even more than one magnitude. The differential magnitudes were affected regardless of which comparison star was used in such images. It is most probable that the light from the central star itself was also refracted. Therefore, I have decided to exclude the two pairs of images, which were affected by the blot. Thus, the final number of useful images, obtained by FTS was 7 for each filter. Due to the small number of data points, the FTS data were used only as the verification of the 68 day lightcurve phase and as an additional constraint for determining the lightcurve shape.
Figure 2.10: Condensation blot, present in some of the FTS images. The shown example is the image taken on 24th March 2012 in V filter. Two pairs of images had to be removed from the data analysis, because the photometric measurements of these images were affected by the blot.

The remaining 7 useful images were photometrically reduced in the similar way as images obtained with BOS. The only difference was that the FTS’s pre-processing data reduction pipeline did not supply the images with accurate equatorial coordinate frames. Thus, the images could not be aligned in their equatorial coordinate frames. Instead, I used a sequence of the four Starlink commands that aligned the images, based on the intercomparison of stellar patterns. The command findobj locates and centroids image features, such as stars. The resulting list of stellar coordinates was then further processed by findoff, which computes the offsets between the lists. Offset lists were then used by register, which determines the required transformations between the images, so that the transformed images would be aligned. For this task, I allowed register to use translations and rotations. Magnification was not required, because the CCD detector and all the settings that could affect the magnification were the same for all the observations. The transformations, determined by register, were then used by tranndf, which transforms the pixel coordinate frames. If the transformations, determined by register, are accurate, all the resulting images are accurately aligned.
in their pixel coordinate frames. All four commands are a part of the Starlink CCDPACK application package.

The following photometric data reduction process of the aligned images is the same as for the BOS images. Fig. 2.11 shows the differential magnitudes in V filter, whereas in Fig. 2.12 the differential magnitudes are given for B filter.

In contrast to the BOS images, star number 1 was not exceeding the linear range of the full well capacity in the FTS images and thus it could be used as a comparison star. Note in Figs. 2.11 and 2.12 that the error bars corresponding to comparison star 1 are almost 50% smaller, compared to the 3 magnitudes dimmer star number 3. Nevertheless, the overlap of the lightcurves of different comparison stars is well within the error bars. As mentioned before, due to the very small number of data points, the FTS data was only used as an aid for the characterisation of the 68 day lightcurve.

![FTS - V Lightcurve](image)

Figure 2.11: The lightcurve of the FTS in V filter. As for every lightcurve plot in this thesis, the scales of all vertical axes are the same.
Figure 2.12: The lightcurve of the FTS in B filter. Note how the smaller signal to noise in B filter affects the error bars, compared to V filter.

2.3 Liverpool Telescope

The final and also the largest photometric dataset was obtained by the Liverpool Telescope (LT), located on the island of La Palma on the Canary Islands. Thirty observations were carried out from 31st March 2012 to 20th July 2012, and resulted in a final number of 36 and 34 data points in the V and U filters, respectively. In addition, the dataset also contained images for 7 data points in R filter, which were originally required for the pointing purposes for spectroscopy of the central star. Because of the very small number of data points and incompatibility with other filters, the dataset in R filter was not used in the final photometric analysis.

The LT images had a smaller field of view, compared to the BOS or FTS. And because the smaller field of view was also coupled with a significant pointing inaccuracy (see Fig. 2.13), the stars 1, 2 and 3 were the only labelled stars from the LT dataset, which were present in every image. Therefore, differential photometry could only be provided for the comparison stars 1 and 3.
Figure 2.13: Field of view and pointing accuracy of the LT dataset. The image was taken on 31st March 2012 in V filter. The field of view is 4.7’×4.7’ (for the comparison with the BOS, see Fig. 2.2), whereas the pointing accuracy can be estimated from the gray padding, which represents the pointing extent of the entire dataset for V filter. Some images were also rotated. The only stars with relatively high $S/N$, visible in every image, were the central star and comparison stars 1 and 3. The bright star near the right edge of the image was present only in 14 out of 36 combined images in V filter.

As for the other two datasets, the first step in reducing the LT data was again to align the images. Similar to the BOS, the LT’s pre-processing pipeline provides accurate equatorial coordinate frames to the images, which enable a very straightforward alignment with the command wcsalign. However, for some images, particularly for the ones with the smallest signal to noise ratios, the pipelined stellar registration failed to detect a sufficient number of stars. As a result, the comparison with the position catalogues was not successful and such images were not provided with accurate equatorial coordinate frames. Therefore, I had to align the LT images with the intercomparison of stellar patterns. This was achieved in the same way as for the FTS images, with a set of four Starlink commands: findobj, findoff, register and tranndf.
The exposure times of the images provided by the LT were very short, 3 and 24 seconds for V and U, respectively. Because of these short exposure times the resulting signal to noise ratios were relatively small. However, the LT dataset was acquired in sequences of three images, which were combined in order to increase the signal to noise ratio. The simplest way of combining the images is to add the intensities of the individual images. Adding \( N \) images together can theoretically improve signal to noise ratio by a factor of \( \sqrt{N} \) (IRAF Tutorial). A similar increase of signal to noise ratio is achieved by a median, which in contrast to adding does not increase the signal, but instead reduces the amount of noise by a factor of \( \sqrt{N} \). Besides the noise reduction of the source and sky intensity values, the readout noise is also reduced by the same factor (IRAF Tutorial). In addition, producing a median image, instead of adding, removes all the cosmic rays and any residual hot pixels and avoids the combined intensities from exceeding allowable intensity values (e.g. 65535 for 16 bit data class). Therefore, I decided to combine the images with their median intensity values. Every sequence of the three aligned images was combined using the CCDPACK command `makemos`, with a specified parameter `method=median`.

After the images were aligned and combined, the rest of the data reduction was the same as the photometric data reduction process for the BOS.

Because of the lower signal to noise ratio, particularly in the U filter, and because of the big, 3 magnitudes difference in brightness between the stars 1 and 2, it is worth verifying the curve of growth and aperture radius. Figs. 2.14 and 2.15 show the curve of growth for U filter.

The average stellar FWHM, measured for the example shown in Figs. 2.14 and 2.15, was 4.0 pixels. The aperture radius, set to one FWHM still roughly corresponds to the minimal radius where most of the signal is included and where the corresponding photometric signal to noise ratio is the highest (Mighell 1999). But on the other hand, due to the low signal to noise ratio, the differences in brightness cause noticeable differences in the shapes of the curves of growth for U filter. However, these differences only result in a 0.05 magnitude offset in differential magnitudes and less than 0.01 magnitude increase of uncertainties of the final lightcurve. Therefore, because of this very small contribution to the final error budget for the LT U photometry, I decided to be consistent with the other datasets and to set the aperture radius to one FWHM. As indicated earlier, the discrepancies between the curves of growth were the largest for the LT U dataset, which had the lowest signal to noise ratio (see Table 2.1). The differences between the curves of growth for the LT images in V filter were negligible.
After examining the impact of the aperture radius for the images with the lowest signal to noise ratios and after performing the photometric data reduction process, described in subsection 2.1, the final LT lightcurves were obtained. Fig. 2.16 shows the differential magnitudes for the filter V, whereas Fig. 2.17 presents the lightcurve for U filter.
Figure 2.16: The final lightcurve for the LT V filter. The lightcurve consists of 36 data points, obtained during 30 observing nights between 31st March 2012 and 20th July 2012.

Figure 2.17: The final lightcurve for the LT U filter. The lightcurve consists of 34 data points. Note that the scale of vertical axes is smaller than for the lightcurve in V filter. Due to the lowest signal to noise ratio, the reliability plot, shown in the upper panel, exhibits the largest scatter among the obtained datasets.
We can see in Figs. 2.16 and 2.17 that both lightcurves exhibit one full brightness variation period. Fig. 2.18 compares the measured lightcurve with the predicted lightcurve, reported by Mikulašek et al. (2007). The measured lightcurve, plotted in Fig. 2.18, is the lightcurve with the highest signal to noise ratio, which was attained by the differential magnitudes between the central star and the comparison star 1 imaged in V filter. The predicted lightcurve is represented by equation (5) as a sinusoidal 0.8 V magnitude brightness variation, with timing maxima as shown in equation (3) (Mikulašek et al. 2007):

\[ V = -2.13 + 0.4 \cdot \cos \left( \frac{JD - 2449862.02}{68.101 \cdot 2\pi} \right), \]  

(5)

The offset value -2.13 in equation (5) was adjusted accordingly to the average differential magnitude between the central star and star number 1.

Figure 2.18: The comparison between the LT lightcurve and predicted lightcurve. The predicted lightcurve, plotted by a red line, is a sinusoidal 0.8 V magnitude brightness variation, as parameterized in equation (5) by Mikulašek et al. (2007).

Fig. 2.18 shows a good agreement with the predicted ephemeris of the light maxima, as well as with the previously observed amplitude of 0.8 magnitude. However, the clear discrepancies in the shape of the lightcurve, especially between the first maximum and the second minimum brightness, is most probably due to the deviation of the current lightcurve shape from an ideal sine function (see Fig. 1.11). We can also see three
narrow dips in the lightcurve, similar to the dips noted in the lightcurve of the BOS. The comparison and correlation between the dips for both telescopes is further discussed in subsection 2.4, also see Fig. 2.21.

The smaller scale of the vertical axes on Fig. 2.17, compared to Fig. 2.16, suggests that the amplitude of the brightness variation in U filter is larger than in V filter. Fig. 2.19 presents a more accurate comparison between the two different filters.

We can see in Fig. 2.19 that the shapes of lightcurves in different filters are similar. On the other hand, the difference between the axes scales suggests that the brightness variation amplitude is larger for U filter and therefore that the central star exhibits the reddening effect along with the brightness variations. The amount of additional periodically varying reddening can be estimated from the difference between the lightcurve amplitudes $U_{\text{amp}} - V_{\text{amp}}$ in Fig. 2.19:

$$\Delta E(U - V) = U_{\text{amp}} - V_{\text{amp}} = V_{\text{amp}} \cdot \frac{U_{\text{axis}}}{V_{\text{axis}}} - V_{\text{amp}} = 0.382,$$

where $V_{\text{amp}}$ is the 0.8$m$ lightcurve amplitude in V filter, whereas the amplitude $U_{\text{amp}}$, can be calculated from the ratio between the axes scales $U_{\text{axis}}/V_{\text{axis}} = 1.33/0.9$, used to overlay the two lightcurves in Fig. 2.19. The estimated amount of reddening in equation
could correspond to the observed amplitude of brightness variations due to periodically varying amount of gas and dust absorption (see equation (10)).

Another discrepancy between the lightcurves of the two different filters emerges from the brightness dips that appear shallower in U filter. The obvious contribution to this observed discrepancy is the different axes scaling due to reddening. Thus, even if the dips have the same depth for both filters, they appear shallower in U filter. It is clear however, that due to relatively large uncertainties and small number of data points, it is not possible to make an accurate comparison between the LT V and U lightcurves.

The following subsection extends the lightcurve analysis to all three telescopes.

2.4 Combined Results

This subsection combines the lightcurves, obtained by the three different telescopes. The first important goal is to extend the comparison with the lightcurve prediction, provided by Mikulašek et al. (2007), to the entire combined lightcurve and possibly to determine the shape of the lightcurve. Secondly, the combined lightcurve is used to verify and determine any potential periodicity in the brightness dips, noted in the BOS and LT data.

The accuracy of the combined lightcurve is challenged by several aggravating circumstances. First of all, the lightcurves were obtained by three different telescopes; secondly, three different filters were used, among which, none of them was used by all of the three telescopes; and lastly, the datasets do not chronologically overlap (see Table 2.1). Although the reliability of the combined lightcurve is diminished by the number of factors listed above, it is still possible to examine the combined lightcurve for the objectives of this project.

A set of three individual lightcurves was chosen from the number of different filters and different comparison stars in a way to obtain the highest signal to noise ratio and compatibility of the final combined lightcurve. For this reason, the combined lightcurve consists of the lightcurves with comparison star number 3 in B filter for the BOS and star number 1 in V filter for the FTS and LT. Because of this incompatibility between the first and other two telescopes, I had to offset the BOS differential magnitudes. On the other hand, the FTS and LT datasets were directly compatible because of the same filter and the same comparison star. Fig. 2.20 shows the combined lightcurve, which is compared to the predicted lightcurve, parameterized in equation (5) by Mikulašek et al. (2007).
Figure 2.20: The comparison between the combined and predicted lightcurves. The configuration of the combined lightcurve is given in the title of the plot. For compatibility reasons, the BOS differential magnitudes were offset by -3.4\text{m}. The red line represents the predicted lightcurve, given in equation (5). The data points are connected with lines for easier presentation purposes.
Fig. 2.20 shows an overall agreement between the combined and predicted lightcurve. Our data are largely consistent with the amplitude and the period of the brightness variations, reported by Mikulašek et al. in 2007. The comparison confirms that the photometric nature of the central star remained the same since the last Mikulašek observation in 2002. The disagreement between the observed and predicted decreasing part of the lightcurve is most likely due to the current deviation of the lightcurve’s shape from the assumed sinusoidal brightness variations. The LT and FTS datasets suggest that the current lightcurve shape exhibits slightly flattened lightcurve minima, which were occasionally observed in other surveys (see Fig. 1.11), and were reported as the non-periodic long-term lightcurve shape variations (Mikulašek et al. 2007, also see subsection 1.7.2). The direct comparison with the BOS dataset is more uncertain because of a different filter. However, the BOS brightness increase still coincides with the predicted lightcurve. Clearly, many more data points would have been required for a more accurate comparison with the predicted lightcurve.

The combined lightcurve also provides a platform for examining any potential correlation between the brightness dips, observed in the lightcurves of the BOS and LT. Fig. 2.21 shows the combined lightcurve with engraved periodic marks, centred on the brightness dips.
Figure 2.21: Periodicity of the brightness dips. The engraved vertical lines are separated by 17.2 days and shifted for the best possible alignment with the observed dips.
Fig. 2.21 indicates that the brightness dips are periodical with a period of \((17.2 \pm 0.1)\) days. The identified period coincides with the BOS preliminary results, which suggested a period of \(\sim 16\) day. Thus, the analysis of the combined lightcurve tentatively confirms the preliminarily suggested periodicity of the brightness dips and refines it to 17.2 days.

The following two figures show the periodicity analysis of the 0.8 V magnitude brightness variations and the periodicity analysis of the brightness dips. Periodogram of the combined lightcurve is shown in Fig. 2.22. Fig. 2.23 shows the periodogram of the combined lightcurve, after it was subtracted by the 0.8 magnitude brightness variation, which is expressed in equation (5). Subtracting was required in order to extract a lightcurve which corresponds only to the brightness dips. The periodogram of the combined lightcurve (Fig. 2.22) was obtained with a Lomb-Scargle type of periodogram analysis. This type of periodogram is calculated based on the least-square fitting of sine functions. The main advantage over the fast Fourier transform technique is that the Lomb-Scargle technique is designed for unevenly sampled data (Birney et al. 2006). A similar Lomb-Scargle analysis of the subtracted lightcurve yielded no useful signal corresponding to the sharp brightness dips, because the dips could not be well fitted with a sine function. Therefore, the second periodogram, shown in Fig. 2.23, was instead calculated by the classical Fourier transform technique. Both periodograms were obtained with application package PERIOD.

![Figure 2.22: Lomb-Scargle periodogram of the combined lightcurve. The spike centred at the frequency 0.0141 days\(^{-1}\) corresponds to a period of (70.9 \pm 3.2) days. The period of 17.2 days, identified in Fig. 2.21, might correspond to a small frequency feature near 0.058 days\(^{-1}\).](image)
Despite the small number of data points, the long known 68 day brightness variation period can be seen clearly in the periodogram in Fig. 2.22. On the other hand, the periodicity of the brightness dips, with a period of 17.2 days, can only be identified as a weak and unreliable frequency feature, present in both periodograms.

If the observed 17.2 days periodicity is real, it is possible that the dips are a result of an eclipsing binary. However, to obtain a sensible eclipsing lightcurve, the shape of the subtracted lightcurve has to be determined accurately. Due to the low number of data points the shape of the 68 days lightcurve could not be modelled accurately and therefore the subtracted lightcurve contains large uncertainties, which are reflected in the inconclusive periodogram in Fig. 2.23. It was also impossible to accurately fold the subtracted lightcurve and to model the potential eclipses.

The photometric analysis has proven that for an accurate determination of the shape of the lightcurve and for a reliable detection of any potential eclipsing events, many more data points would be required, preferably by the same photometric filter. The eclipses modelling would also require a higher signal to noise ratio.

Nevertheless, photometric observations firstly confirmed the period and phase, reported by Mikulašek et al. (2007) (see Figs. 2.20 and 2.22). Secondly, the FTS and LT
datasets helped to constrain the shape of the lightcurve, which was examined for the purposes of the accurate positioning of the LT spectra on the brightness lightcurve. Thirdly, the analysis of the LT U and V lightcurve identified the periodically variable reddening effect, with an estimated amplitude of $\Delta E(U-V)=0.38$ (see Fig. 2.19 and equation (6)). And finally, the correlation between the brightness dips, noticed in the BOS and LT datasets, suggested that they appear periodically with a period of 17.2 days (see Fig. 2.21), which gives rise to a tentative possibility of eclipsing events.
3 Spectroscopy

A detailed study of the stellar spectra can provide a wealth of information, such as kinematic properties, stellar atmospheric parameters, as well as the information about the circumstellar absorption or emission material. The goal of the analysis of the central star’s spectra was to examine any potential spectral variations and to verify if the variations are correlated with the phase of the 68 day brightness variations. Information derived from the spectroscopic observations could help to indentify the cause of the observed unusual properties of the central star and the PN, which are outlined in subsection 1.7.2.

The primary spectroscopic dataset of six spectra, taken at different brightness phases, was provided by the Liverpool Telescope (LT). The spectra were taken in two different wavelength ranges, which jointly covered two of the most intriguing spectral features of this object: Hα and Hβ. However, the initial quality check of the acquired spectra revealed that in the case of three spectral pairs the telescope was not pointing at the central star. In addition, one spectrum at shorter wavelengths had a very low signal to noise and could not be used. Because of this greatly reduced number of useful spectra, I have extended the spectral sample with 25 archived spectra obtained with the Isaac Newton Telescope (INT) between 1996 and 1999. However, these spectra do not cover the first two Balmer lines. The second extension of the spectral sample was provided by eight archived echelle spectra taken during the same night with Anglo-Australian Telescope (AAT). AAT spectra were used to check spectral variations on hourly timescales. Table 3.1 provides an overview of the spectroscopic observations. The position on the brightness phase curve of every spectrum is given in Fig. 3.1.

<table>
<thead>
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<th>telescope</th>
<th>duration of observations</th>
<th>wavelength range [Å]</th>
<th>exposure time</th>
<th>S/N</th>
<th>number of spectra</th>
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<td>3820-4800</td>
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<td>161</td>
<td>25</td>
</tr>
<tr>
<td>AAT (3.9m)</td>
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<td>3700-5100</td>
<td>1800s</td>
<td>20</td>
<td>8</td>
</tr>
<tr>
<td>LT (2m)</td>
<td>16th Jun 2012 – 19th Jul 2012</td>
<td>3900-5100 5900-8000</td>
<td>1200s</td>
<td>31</td>
<td>3</td>
</tr>
</tbody>
</table>

Table 3.1: Overview of spectroscopic observations. The INT exposure time is given as a median value. Mean signal to noise ratios are calculated for central wavelengths, except for AAT, where it is provided for Hβ region.
Figure 3.1: Brightness phase for each of the spectra. The phase was calculated based on the lightcurve, parameterized by Mikulašek et al. (2007) (see equation (5)).
3.1 Isaac Newton Telescope

The earliest and also the largest spectroscopic dataset consists of 25 archived spectra, provided by the Intermediate Dispersion Spectrograph (IDS), mounted on the Isaac Newton Telescope (INT), La Palma, Canary Islands. The spectra of the central star were taken between 1996 and 1999 as a part of a larger survey for determining the binary fraction of PNe central stars (Sorensen et al. 2004). The survey revealed the spectral variability of Sh2-71’s central star, but did not provide any further discussion on the origin of the observed variability, nor extract any additional parameters. Until today, the spectra had not been examined in detail.

To be able to accurately analyse and obtain reliable spectra, I had to perform an accurate spectroscopic data reduction. In contrast to photometric observations, INT spectra have not been automatically pre-processed. Therefore, the first step in INT data reduction was to crop the raw spectra, create and subtract master bias frames, prepare master flat frames, rotate the spectra to the horizontal orientation and finally remove the cosmic rays. The exposure times of the spectra were relatively long, usually 1500 seconds, and consequently, the cosmic ray contamination was substantial. Starlink command bclean compares the intensity of each pixel with its neighbours and if the specified threshold parameters are attained, such pixels are identified as cosmic rays. Their values are then replaced by the interpolated intensities of the neighbouring pixels, which effectively removes the cosmic rays from the image of the spectrum. bclean and other commands, related to spectroscopic data reduction, are part of Starlink application package FIGARO. The pre-processing had to be performed for the spectra of Sh2-71’s central star, as well as for the wavelength calibration spectra. Fig. 3.2 shows an example of the raw spectrum before and after pre-processing.

The pre-processed images were then further processed with Starlink package ECHOMOP, which extracts the spectra from the pre-processed spectral images. ECHOMOP was developed specifically for the spectroscopic data reduction of AAT echelle spectra. However, the package can be used for any other echelle spectra, as well as for single order spectra such as INT/IDS. The first step in data extraction process with ECHOMOP is to locate the initial positions of all the dispersion orders on the pre-processed spectrum. Because the INT spectra are single order type of spectra, only one order had to be located. ECHOMOP then traces the orders’ positions along the entire dispersion axis with a specified order of polynomial. A user also specifies the region of
Figure 3.2: Spectral pre-processing. Fig. 3.2a is the raw spectrum taken on 10th July 1996 near Hδ, which can be seen as an absorption feature in the centre of the figure. Fig. 3.2b is the same spectrum, shown with the same greyscale range, after applying the pre-processing sequence. The comparison demonstrates the effect of bias subtraction and cosmic ray removal.

the object and sky for every dispersion order. In the next step, a flat field correction is applied to the spectrum. Finally, ECHOMOP adds the pixels’ intensity values along a predefined object region. At the same time, a median sky value is calculated from the sky region and subtracted from the intensity sum of the object. Gain and readout noise are also taken into account, which converts the analogue to digital units (ADUs) into number of electrons. The extracted spectrum is provided in electron counts versus pixel position along the dispersion axis. The same process is repeated for wavelength calibration spectra.

Because the exposure times of the central star’s spectra were long – close to half an hour – the temperature variations of the spectrograph and the changing altitude of the object might result in an inaccurate wavelength calibration if only one calibration spectrum was obtained. Therefore, each INT spectrum was supplied with two wavelength calibration spectra, one taken immediately before the spectrum of the central star and the second immediately after. The spectral lines of each of the extracted CuAr wavelength calibration spectra were identified with FIGARO command arc, which fits the entered wavelengths of the identified calibration spectral lines with a polynomial. arc command uses the best fit of the wavelengths entries in order to add a wavelength axis to the calibration spectra. Every pair of wavelength axes of the calibration spectra was then interpolated to their intermediate wavelengths. The interpolation of two arc spectra was performed with a command xcopi, which also copies the interpolated wavelength axis to the object’s spectrum. However, the differences between using an interpolated wavelengths and a single wavelength calibration spectrum were small, of the order of 0.02 Å, and similar to the standard
deviation of the wavelength fit. Small corrections imply that INT spectrograph was thermally stable and that the changes of the object’s altitude did not strongly affect the measured wavelengths.

After the wavelength axes have been added to the extracted spectra, it is necessary to rebin the spectra to a linear wavelength scale. This was done using a Starlink command `scrunch`.

Rebinned spectra are the end result of INT spectroscopic data reduction. The reduced spectra were hereafter plotted, manipulated and analysed with Starlink spectral analysis tool SPLAT. At this stage, the spectra are given as intensities in number of electron counts versus wavelength (see Fig. 3.3). Because the spectra of spectrophotometric standard stars were not provided within the INT dataset, the spectroscopic intensities of the central star could not be flux calibrated. However, the best way of visualising and comparing a set of spectra is to normalise their intensity values, regardless if the intensities are flux calibrated or not. Normalisation was performed in SPLAT by dividing the spectral intensities with a black-body shaped curve. An example of the normalised spectrum is shown in Fig. 3.4. Spectral lines were identified with the aid of The Interactive Database of Spectral Standard Star Atlases (SpectroWeb 1.0) and NIST Atomic Spectra Database. In addition to ordinary spectral lines, a wide absorption feature, centred at 4428 Å has been identified as a diffuse interstellar band (DIB) (Herbig 1995).
Figure 3.3: INT spectrum of Sh2-71’s central star. Intensities are given in arbitrary units. Labelled spectral lines’ wavelengths are given in Å. An unsuccessfully removed cosmic ray is labelled with CR. The spectrum was taken on 10th July 1996.

Figure 3.4: Normalised INT spectrum. The drop of continuum intensities at shorter wavelengths is caused by the overlap of the gravitationally broadened Balmer absorption lines.
The INT spectroscopic dataset comprised 25 spectra. The spectra were obtained at various position angles of the slit, as the slit position angle has always been set close to the parallactic angle to minimise the atmospheric effect of differential refraction. But since the object was generally observed close to the celestial meridian, the position angle of the slit was usually close to 0, i.e. in the orientation N-S. Slit length was 3.3', whereas the mean slit width was 1.2". Table 3.2 lists all the observing dates, along with the seeing and central wavelength’s signal to noise ratios.

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Table 3.2: List of INT observing dates, seeing and S/N ratios. Seeing is given as the FWHM of the spectra in number of pixels. S/N ratios are calculated for the central wavelengths.

The whole set of 25 normalised INT spectra is shown in Fig. 3.5. The spectra are ordered chronologically, provided with the largest intensity offset for the earliest spectra and the smallest for the latest spectra. Fig. 3.6 shows the same set of normalised spectra, except that the spectra are ordered according to their 68 days brightness variation phase (see Fig. 3.1).
Figure 3.5: A set of normalised INT spectra, ordered chronologically. Note the variability of the spectra. The most variable spectral features include the occasional absence of metal absorption lines, variable emission component of the Balmer lines and unstable equivalent width of HeI line at 4471 Å.
Figure 3.6: A set of normalised INT spectra, ordered by their phase. Brightness phases were calculated according to the brightness variation parameters, published by Mikulašek et al. (2007). The phases of individual observing nights are provided under each corresponding date (also see Fig. 3.1). Note that the spectral variability is not correlated with the brightness phase.
Fig. 3.5 provides evidence for a very pronounced spectral variability, which is mainly manifested in the variable shapes of the Balmer lines due to the variable circumstellar Balmer emission and/or self-absorption component. In addition, the strength of metallic absorption lines and the 4471 Å HeI line exhibit easily noticeable variations. In that respect, the spectral variability, noted by Mikulašek et al. (2007), is confirmed. However, in contrast to their spectroscopic results, the INT spectra do not exhibit any [OIII] lines (see Fig. 1.12). The most likely reason for Mikulašek’s detection of [OIII] lines is that their pointing accuracy was insufficient and the resulting spectrum became contaminated by the nebular [OIII] lines. On the other hand, the spectral variability could also arise from the variable seeing. Larger seeing reduces the stellar intensity, whereas the intensity of the nebula remains the same. Thus, depending on the spectrograph’s slit width, the larger seeing might increase the ratio between the nebular and stellar contribution in the final spectra. Spectral variability, observed by the INT, does not show any correlation with the seeing, given in Table 3.2. Different signal to noise ratios also did not contribute to the observed variability. The only spectrum which was affected by the low signal to noise ratio, was taken on 22nd April 1997 (see Fig. 3.5 and Table 3.2).

Surprisingly, Fig. 3.6 does not provide any reliable correlation between the spectral variability and the brightness phase. It must be taken into account, however, that due to the variable lightcurve shape, the central star, observed at the same phases of the different brightness cycles, could have different brightness. Even so, due to the many discrepancies among the spectra with similar phases, it is reasonable to reject any correlation of the spectroscopic variability with the brightness phase. A detailed examination of Fig. 3.5 also excluded the periodicity of 17.2 days, which was identified as a possible eclipsing period (see subsection 2.4). The cadence of the INT spectroscopic observations was too low and the spectra were consequently too sparse to determine any potential unknown periodicity of the spectral variations.

3.2 Anglo-Australian Telescope

The archive of the Anglo-Australian Telescope (AAT) provided eight high resolution ($\lambda/\Delta\lambda=45000$) echelle spectra. The spectra were obtained by the University College London Echelle Spectrograph (UCLES) during the single observing night of 21st May 2000. Until now, these spectra have never been reduced (Pollacco 2013, private communication). The interval between the individual spectra was close to 30 minutes,
and therefore, the set of AAT spectra could be used to verify any spectral variations on hourly timescales.

The AAT spectra have been reduced with a similar spectroscopic data reduction process as the INT spectra. However, the selection of dispersion orders was different. The INT spectra had only one dispersion order, whereas the AAT echelle spectra consisted of 35 cross dispersed orders that had to be located accordingly in the ECHOMOP extraction process.

The second and also the only remaining difference from the INT data reduction process was that the AAT spectra were supplied with only one wavelength calibration spectrum for the entire AAT spectroscopic dataset. Therefore, the interpolation of AAT ThAr arc wavelengths was not required. The wavelength axis was copied to the central star’s spectra with the command xcopy, which in contrast to xcopy does not interpolate the wavelengths. Since the entire AAT spectroscopic dataset was supplied with only one wavelength calibration file, it is anticipated that the reliability of the calibration is reduced. Furthermore, the direction of the arc dispersion axis was reversed with respect to the star’s spectra, possibly due to improper settings for the acquisition of the arc spectrum. The data reduction revealed that even after the reversed direction has been taken into account, the arc wavelengths exhibited an order-dependent offset with respect to the spectra of the central star (see the positions of Hβ and Hγ lines in Figs. 3.7 and 3.8, respectively). However, the offset and unreliable calibration of the wavelength axes did not have any negative impact on the study of the spectral features’ variations.

One of the most easily noticeable spectral variations in the INT spectra was a variable circumstellar emission and/or self-absorption component of the Balmer series. Therefore, the spectral variability is most easily observed in the Balmer lines with a strong emission component. The Balmer line with the strongest emission component, Hα (6563 Å), lies outside the AAT wavelength coverage, whereas Hβ (4861 Å), the line with the second strongest emission component, is present in the AAT spectral wavelength range, which covers 3700 – 5100 Å. In addition, Hβ lies within the wavelength range with one of the highest signal to noise ratios of the AAT spectra. Therefore, among all the spectral features covered by the AAT spectra, the Hβ line provides the most accurate verification of the spectral variations. Fig. 3.7 shows the normalised set of fully processed AAT spectra in the wavelength region, which encompasses the Hβ line. For comparison, the spectrum of Hγ is shown in Fig. 3.8.
Figure 3.7: Normalised set of AAT Hβ profiles. The AAT echelle spectra were obtained on 21st May 2000 between 15:07:20 UT and 19:34:37 UT. The spectroscopic set with a timespan of 4.5 hours does not reveal any noticeable unequivocal variations of the Hβ profile’s shape. Note a blueshifted ~3 Å offset in the wavelengths, most likely due to improperly obtained arc spectrum.
Figure 3.8: AAT spectrum of Hγ. This shape of the Hγ profile has been previously seen in certain INT spectra (see Fig. 3.5). The spectrum was taken on 21st May 2000 at 15:07:20 UT. In contrast to the blueshifted offset of the Hβ region, the Hγ (4340.47 Å) region exhibits a redshifted wavelength offset, again this is due to a systematic calibration error.

The INT spectra exhibited pronounced spectral variability (see Fig. 3.5), but the analysis did not provide any evidence for periodicity, nor determined the timescale in which the variations occur. The AAT spectroscopic dataset excludes the possibility that the spectral variations occur on the hourly timescales because the set of AAT spectra, shown in Fig. 3.7, does not exhibit any pronounced spectral variations in the 4.5 hours timespan. According to results, extracted from the INT and AAT observations, it is reasonable to expect that the spectral variations must occur on timescales of the order of days.

3.3 Liverpool Telescope

The Liverpool Telescope (LT) spectra were originally obtained to provide a reliable and accurate verification of the correlation between the spectral variations and brightness phase. The LT spectroscopic dataset consists of six spectra, taken at six different brightness phases, including the maximum and minimum brightness. At the same time, simultaneous photometric observations were taken to provide the actual shape of the lightcurve and to detect any potential brightness dips, which could influence the observed spectra. The spectra were taken in two different wavelength ranges, which
jointly covered two of the most intriguing spectral features of this object: Hα and Hβ. The resolving powers of the red and blue spectral components were 5300 and 5500, respectively.

The spectra, obtained by the LT spectrograph, Fibre-Fed Robotic Dual-Beam Optical Spectrograph (FRODOSpec), are in contrast to other two spectroscopic datasets, multifibre type of spectra. The LT fibre bundle consists of 12×12 fibres, with a total field of view of 10"×10". Fig. 3.9 shows the intensity sum of every fibre for all six observing nights.

![Image](https://example.com/image.png)

Figure 3.9: The LT pointing verification. The images represent the intensity sum in each of the spectrograph’s fibres for all six observing nights. Upper six images correspond to the red part of the six spectroscopic pairs, whereas the bottom images represent the blue parts. Note the absence of the central star during the first three observing nights due to the wrong pointing of the telescope. Also note the low S/N for the blue spectrum taken on 16th June 2012. The pointing and the S/N issues reduced the number of usable individual spectra from 12 to 5.

Fig. 3.9 demonstrates the incorrect telescope pointing during the first three observing nights. Therefore, the first three spectral pairs had to be excluded from the LT dataset. In addition, the blue part of the spectral pair, taken on 16th June 2012, exhibited a very low signal to noise ratio, which resulted in exclusion of another spectrum. Therefore, the final LT set of usable spectra includes only three red and two blue components of the obtained spectral pairs. An example of a usable raw LT multifibre spectrum is shown in Fig. 3.10.
Figure 3.10: Zoomed H\(\alpha\) region of a raw LT multifibre spectrum of the central star. The central bright spectral feature corresponds to the H\(\alpha\) emission (6563 Å), whereas the two narrow features on either side of the H\(\alpha\) emission correspond to the nebular [NII] emission (6548 and 6584 Å). The spectrum is a red part of a spectral pair obtained on 19th July 2012.

The raw spectrum, represented in Fig. 3.10, demonstrates that the circumstellar H\(\alpha\) emission dominates the nebular [NII] emission only within the few central fibres. Every fibre contributes a circular 0.82'' field of view to the total 10''×10'' field of view of the entire bundle. With a known fibre’s field of view, it is possible to estimate the extent of the central H\(\alpha\) emission region. Fig. 3.10 suggests that the core region of the hydrogen emission extends roughly over 4 fibres on either side of the nucleus. This corresponds to an estimated angular radius of 4×0.82''=3.3'', which implies a compact H\(\alpha\) emitter. The estimated radius is in a near agreement with the value of 3.5'', which was reported by Cuesta et al. (1990).

The data reduction of the LT multifibre spectra differs from the data reduction of the two preceding spectroscopic datasets in only one step. Instead of extracting a spectrum from the dispersion orders, it is necessary to add the intensities of the brightest fibres. The rest of the data reduction process is similar to the process, discussed in subsection 3.1. However, the manual data reduction of the LT spectra did not provide any improvement when compared to the LT automated spectroscopic data reduction. In addition, the spectra, automatically reduced by the FRODOSpec level 2 (L2) data reduction pipeline, exhibited even higher S/N ratios than manually reduced spectra, most likely due to the higher number of added fibres. Therefore, I have decided to use the fully automatically reduced spectra. The only disadvantage of the L2 automatic data
reduction is its inability of removing the cosmic rays. However, the presence of cosmic rays does not affect the analysis of this project. As for the previous two spectroscopic datasets, the LT observations did not supply flux calibration files. Thus, the intensities of the LT spectra could not be flux calibrated and are given in arbitrary units.

The whole wavelength range of the fully processed red and blue spectral components is shown in Figs. 3.11 and 3.14, respectively. A selection of intriguing spectral features is plotted with larger wavelength scales in Figs. 3.12, 3.13, 3.15 and 3.16. Because the spectra had different intensities, they have been multiplied to comparable intensities, and then offset for clarity. The corresponding brightness phases are plotted in Fig. 3.1.

Figure 3.11: The red components of the LT spectra. The spectra have been reduced by the L2 FRODOSpec automatic data reduction pipeline. For comparison reasons, the spectral intensity values have been multiplied and offset as indicated in the title of the plot. The wavelength of the OI line (7773.8 Å) is given for the unresolved OI triplet (NIST Atomic Spectra Database).
Figure 3.12: Hα wavelength region of the LT spectra. Note the differences in the Hα emission profile. Small emission features on either side of the Hα feature correspond to the nebular [NII] emission (6548 and 6583 Å).

Figure 3.13: Telluric O₂ (7605 Å) and stellar OI (7774 Å). The apparent redshift changes of the OI absorption triplet (heliocentrically corrected 120 km/s) mostly result from the variable OI profile’s shape. The static O₂ telluric spectral feature is shown for comparison.
Figure 3.14: The blue components of the LT spectra. The small wavelength scale enables us to see the stellar gravitationally broadened Balmer absorption components.

Figure 3.15: Hβ and Hγ. Note the similarities in the variations of Hβ (a) and Hγ (b) spectroscopic profiles. The ratio between the bluer and redder emission peaks increased within 18 days in both Balmer profiles. Hα feature, shown in Fig. 3.12, exhibited a similar effect. In Fig. 3.15a no offset has been applied to the red Hβ spectrum so that the change in profile shape can be seen more clearly.
The LT spectra are the only spectroscopic dataset of this research project, whose wavelength coverage includes Hα, the most prominent optical spectral feature of Sh2-71’s central star. The measured basal width and peak separation of Hα emission are ~26 and 4.7 Å, respectively. This is in close agreement with the values, reported by Cuesta et al. (1993) (see subsection 1.7.2 and Fig. 1.13c). However, the most important outcome of the LT spectroscopy is the observed spectral variability. The variations of Balmer line profiles (see Figs. 3.12 and 3.15) coincide with the same profile changes of other spectral lines (see Fig. 3.16). The similarities in the spectral variability suggest that all variations result from the same mechanism. Possible reasons of the observed variations are variable blueshifted stellar wind absorption and redshifted accretion absorption component that could affect the profiles of various elements, including all the Balmer lines and FeII. An inflowing accreting material could also be responsible for the observed inversed P Cygni profile – enhanced blueshifted emission and redshifted absorption. Interestingly, the LT spectra have shown that P Cygni profile of the Hα feature is not inversed, which in contrast to other spectral features implies outflowing (decreting) material. On the other hand, periodic spectral variations could result from various reasons, e.g. stellar rotation with starspots, pulsations, rotating binary system, etc. A detailed discussion on the possible models for the observed properties is provided in Chapter 4.
The primary goal, the accurate analysis of the spectroscopic correlation with the brightness phase, had to be abandoned due to the fact that less than half of the LT spectra were usable. Despite this significant reduction in the number of spectra, the LT spectroscopy provided valuable spectroscopic information. In contrast to the INT spectroscopic dataset, the LT spectra covered the first two Balmer lines, which exhibit the most pronounced spectral variability. Hα and Hβ, as well as two strong FeII lines, covered by the LT wavelength range, provide an excellent platform for constructing a possible progenitor of the observed properties. However, it has to be reiterated that the spectral variations, observed in the LT spectra, are based on only three red and two blue spectra of the central star.

### 3.4 Spectroscopic Analysis

#### 3.4.1 Radial Velocity

The goal of radial velocity (RV) analysis is to find any potential periodicity in RV measurements and foremost to verify for this object if the 68 days brightness variations are accompanied by a corresponding Doppler effect. The amplitude and the period of RV variations could provide evidence for the central star’s binarity. In addition, the determination of the orbital period could also yield several useful estimated constraints for other binary parameters, e.g. the mass ratio and separation between the two stars.

Stellar RVs are measured by the wavelength positions of the stellar absorption lines. The difference between the observed wavelength and the rest wavelength of the same transition, calculated theoretically or observed in a laboratory, yields the object’s RV, relative to the observer. The relation between wavelength shift and RV is given by equation (7) (Einstein 1905):

\[ \frac{\lambda_{\text{obs}} - \lambda_{\text{lab}}}{\lambda_{\text{lab}}} = \sqrt{1 + \frac{RV}{c}} - 1 \approx \frac{RV}{c}, \]

where \( \lambda_{\text{obs}} \) denotes the observed wavelength, \( \lambda_{\text{lab}} \) is the rest wavelength and \( c \) is the speed of light. RV is positive, when the source is receding. The simplification is valid if the object’s radial velocity is small compared to the speed of light.
The impact of the Earth’s orbital and rotational velocity on the measured RVs of the object can be removed with the heliocentric RV correction. For this reason, I used Starlink application RV, which yields the observer’s radial velocity in the object’s direction for the specified observing location and observing time. Heliocentric correction is then applied by the subtraction of the observer’s velocity from the measured radial velocity of the object.

The RV analysis was performed for the largest spectroscopic dataset provided by the INT. Many of the spectral lines of the Sh2-71’s central star were affected by the profile’s shape variability that could reduce the reliability of the measured RVs (e.g. see the Balmer lines in Fig. 3.5). Furthermore, certain metallic spectral lines exhibited occasional pronounced decrease of the line strength and in some cases even complete absence of a spectral line (FeII 4233 Å, TiII 4395 Å). Additional constraint in choosing a suitable spectral line was a reduced wavelength overlap due to the different wavelength ranges (CaII 3934 Å). I also avoided the HeI line at 4471.5 Å, whose profile might be influenced by the unresolved TiII line at 4468.5 Å. The line, identified as the most suitable candidate for the RV measurements, was 4481 Å MgII.

The wavelengths of the spectral lines were measured by Starlink spectral analysis tool SPLAT with the Gaussian fit of the lines’ absorption profiles. The peak position of the Gaussian fit provided the wavelength of the stellar absorption line, whereas the uncertainty was calculated from the Gaussian fit error and the standard deviation of the wavelength calibration. The wavelength accuracy of the INT spectra and the reliability of the RV measuring technique were verified by the two RV standard stars, HD136202 and HD187691. The standard stars’ RVs, tabulated by Simbad, lay within the ±5 km/s error bars of the measured RVs.

Fig. 3.17 presents the heliocentric corrected MgII RV measurements for the INT spectroscopic dataset. For comparison, the RVs have also been measured for Hγ. The RVs versus 68 day brightness variation phase is plotted in Fig. 3.18.
Figure 3.17: RV of the central star, derived from the observed redshift of MgII and Hγ. The deviation of the RVs between the two lines is mostly a result of the variable profiles of the Balmer lines.

Figure 3.18: RV of the central star versus the brightness variation phase. The brightness phases correspond to the predicted lightcurve, parameterized by Mikulašek et al. (2007) (see equation (5)).

Figs. 3.17 and 3.18 demonstrate that the RVs, derived from the redshift measurements of Hγ, deviate from MgII RV appreciably more than expected by the error bars. The major contribution to the observed differences can be attributed to the variable Hγ spectral profile. On the other hand, the RVs of both spectral lines exhibit the same
scatter in Fig. 3.18, which does not provide any correlation with the brightness phase for any of the two lines.

In order to increase the accuracy and to verify the RV results, derived from the two individual spectral lines, I also performed a cross correlation among the normalised INT spectra. In contrast to the redshift measurements of the individual spectral lines, the cross correlation compares the shapes of the entire spectra, including the profiles and wavelength positions of all the spectral features on the two cross correlated spectra. The cross correlation can therefore increase reliability of the determined RVs (Nordström et al. 1994). FIGARO command hcross finds the wavelength shifts to obtain the best possible overlaps between the spectra. The command converts the wavelength shifts into radial velocities and also reports the correlation errors. The reference spectrum was chosen to be the first INT spectrum, obtained on 10th July 1996. Therefore all the RVs provided by hcross, were given relative to the RV of the first INT spectrum. Adopting the 2.1 km/s absolute heliocentric RV measurement of the MgII line for the first spectrum, transformed the relative RVs of the other spectra into absolute RVs. RVs obtained by the cross correlation are shown in Figs. 3.19 and 3.20.

![INT RV - Cross Correlated](image)

Figure 3.19: Cross correlated RV of the central star. The average value of the measured RVs is 26 km/s.
Figure 3.20: Cross correlated RV of the central star versus the brightness phase. Note that the scatter of cross correlated RVs is smaller than in Fig. 3.18.

The RVs obtained by the cross correlation, exhibit a smaller scatter as the values obtained from the individual spectral lines. Nevertheless, the scatter, present in Fig. 3.20, proves that the measured RVs are not correlated to the 68 day brightness variation phase. Any other periods of the potential RV periodicity have been examined with different types of periodograms. Fig. 3.21 shows the Lomb-Scargle periodogram of the cross correlation RV measurements.

Figure 3.21: Lomb-Scargle periodogram of the cross correlated RVs. The absence of any frequency features suggests that the variability of the RVs, observed by the INT, is random. The other possible reason for the absence of frequency features are the undersampled RV measurements.
The periodogram, shown in Fig. 3.21, did not detect any periodicities in the analysed sample of RVs. The simplest explanation for this result is that the measured RVs actually are random. A contributing factor to the demonstrated absence of periodicity is an uncertainty of the RV measurements due to the variable profiles of the spectral lines. It is also possible that a potential periodicity is not detected because of the undersampled INT data.

As briefly mentioned in subsection 1.7.2 and further discussed in subsection 4.5, some researchers have suggested that the star observed and analysed in this research project, actually is not the right central star of the PN. However, a useful result, derived from the INT RV analysis, is that the average RV of the observed star is ~26 km/s (see Fig. 3.19). This is in near agreement with the reported average RV of the PN, 24.7 km/s (Schneider et al. 1983). The small difference between the measured stellar RV and the RV of the PN supports the theory that the observed star actually is the central star of the PN.

Another important result was provided by the AAT spectroscopic dataset. Despite the unreliable and offsetted wavelength calibration spectrum, the spectra of the central star did not exhibit any noticeable RV variations within the AAT observing timespan of 4.5 hours (see Fig. 3.7).

### 3.4.2 Synthetic Spectra

Fitting the observed stellar spectra with theoretically computed synthetic spectra can provide a list of potential stellar atmospheric parameters of the observed star. For this task, I used a library of normalised 1 Å/pixel synthetic spectra, compiled by Munari et al. (2005). The synthetic spectra were computed with the Kurucz’s code, based on a variety of atmospheric parameters, including the effective temperature, surface gravity, rotational velocity and metallicity. The parameters, used for the computation of the best fitted synthetic spectrum, have been adopted as the estimated atmospheric parameterization of the observed central star.

Because the Sh2-71’s central star is likely to be a binary (De Marco 2006), the measured spectra are assumed to be composite spectra of two stars of different spectral types. However, if the ionising component of the binary system is assumed to be a white dwarf or even if it is an O sub dwarf, its contribution to the composite spectrum with a hot main sequence star is negligible in optical wavelengths (Smalley 1997). Therefore, even if the central star is a binary, the parameters of the synthetic spectrum, which
provide the best fit to the observed spectrum, can be used for the atmospheric characterisation of the optically dominating main sequence (or giant) star alone.

3.4.2.1 Rotational Velocity

Among all the fitted atmospheric parameters, the determination of the stellar rotational velocity is the most straightforward and was therefore chosen to be fitted first. First of all, the best-fitted value of the rotational velocity is not dependant on any other atmospheric parameter, whereas the inverse is not true. Secondly, the widths of the stellar lines can provide the estimated initial guess for the fitting process, as well as the verification of the final fitted value of the rotational velocity. Stellar rotation is the strongest broadening mechanism for the majority of the stellar absorption lines. However, some lines, especially hydrogen, also exhibit a pronounced gravitational broadening effect, which can be orders of magnitude stronger than the stellar rotation broadening (compare the basal widths of the Balmer series with other lines in Fig. 3.3 or 3.4). As a consequence, changing the stellar rotational velocities in the fitting process has very little effect on the widths of the Balmer lines. It is therefore necessary to fit the spectral profiles of other spectral lines to obtain a reliable value of the rotational velocity. In the spectroscopic survey of 116 Be type stars, Chauville et al. (2001) identified the 4471 Å HeI line as the best choice for the rotational velocity fitting purposes due to the small impact of the circumstellar material on the line’s profile. However, in contrast to Chauville’s suggestion, I used the 4481 Å MgII line, because of the concern that the width of the HeI line might be affected by the 4468.5 Å TiII line. The comparison between different synthetic spectra of different rotational velocities and the measured spectrum of the central star is shown in Fig. 3.22.

The width of the spectral lines was best reproduced by the synthetic spectrum with projected rotational velocity 200 km/s (see Fig. 3.22). This is in agreement with the estimated upper value of 270 km/s, derived from the MgII basal width of ~8 Å. The finding indicates that the central star is a rapid rotator.
Figure 3.22: Determination of the rotational velocity. The first INT spectrum, obtained on 10th July 1996 (black line), is compared to synthetic spectra with stellar rotational velocities ranging from 100 km/s (green line) to 300 km/s (blue line). The width of the absorption lines was best reproduced by the rotational velocity 200 km/s (red line). Other atmospheric parameters are the same for all synthetic spectra: effective temperature 11000 K, \( \log(g) \) 4.0 cm/s\(^2\) and solar metallicity.

### 3.4.2.2 Surface Gravity

The stellar surface gravitational acceleration has the greatest velocity effect on the lightest element, hydrogen. Because the gravitational broadening primarily broadens the hydrogen lines, changes in the gravity parameter are most easily studied for the hydrogen profiles. As indicated in Fig. 3.5, even the higher order Balmer lines also exhibit appreciable emission components. It was therefore necessary to avoid fitting the entire hydrogen profiles, and instead to fit only the Balmer absorption wings, which are unaffected by the emission.

The determination of the best-fitted surface gravitational acceleration depends on the selected effective temperature and vice versa. In contrast to the rotational velocity, the mutual dependence between gravity and temperature requires that the two parameters are determined iteratively. Higher reliability of the iterated parameters is attained by determining the luminosity class before the final temperature classification is made (Garrison et al. 1994). Fig. 3.23 shows the spectroscopic effect of choosing different gravity values for a 11000 K star.
3.4.2.3 Effective Temperature

Different choices of the effective temperature affect the shape of the entire synthetic spectrum. However, the most pronounced synthetic spectral changes in the examined temperature range have resulted in the variable strength of the hydrogen and helium absorption lines. Higher effective temperature of the synthetic spectra resulted in the reduction of the Balmer line strengths, whereas the strengths of the helium lines increased. On the other hand, changes in temperature only slightly influenced the depth of the synthesised MgII line. Therefore, various choices of effective temperature have resulted in different strength ratios between the 4471 Å HeI and 4481 Å MgII line. The ratio between the two lines can therefore be used as a tool for determining the effective temperature (see Table 3.3).
Table 3.3: HeI (4471 Å) to MgII (4481 Å) ratio dependence on the effective temperature. The provided ratios are calculated for synthetic spectra with parameters log($g$)=4.0 cm/s$^2$, [M/H]=0 and $v_{\text{rot}}$=200 km/s.

Fig. 3.24 compares the measured spectrum in the wavelength region of HeI and MgII with synthetic spectra of various effective temperatures. The same synthetic spectra are compared to the Hδ profile in Fig. 3.25.

Figure 3.24: The effect of the effective temperature on the HeI/MgII ratio for the temperatures given in Table 3.3. A higher effective temperature increases the strength of the 4471 Å HeI line, whereas the strength of the 4481 Å MgII line is slightly reduced. The measured HeI/MgII ratio is best reproduced by the synthetic spectrum with effective temperature 11000 K (red line). Note how the HeI line becomes affected by the 4468.5 Å TiII line at lower effective temperatures.
Figure 3.25: The impact of the effective temperature on the Balmer lines. The absorption wings of the Hδ profile are well fitted by the 11000 K synthetic spectrum (red line). This confirms the choice of 11000 K, which was made based on the HeI/MgII ratio (see Fig. 3.24). The same as for the gravity fitting process, the central region of the Balmer profiles have not been fitted, because of the presence of the emission and self-absorption components.

The observed ratio between the HeI and MgII lines in the first INT spectrum was best reproduced by a synthetic spectrum with effective temperature 11000 K (see Fig. 3.24). The measured Balmer profiles were again best fitted with the synthetic spectrum with the effective temperature 11000 K (see Fig. 3.25). Hence fitting the HeI/MgII ratio and fitting the Hδ profile gives a consistent value for effective temperature.

Among other variable spectral features, the observed ratio between the HeI and MgII line has one of the most pronounced spectral variabilities (see Fig. 3.5) and therefore is an indicator of potentially variable temperature. A shortlist of measured line ratios and the corresponding estimated effective temperature are given in Table 3.4.

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</table>

Table 3.4: Estimated effective temperature. The measured ratio between HeI (4471 Å) and MgII (4481 Å) and the corresponding effective temperature (see Table 3.3) is provided for various INT spectra. The INT spectrum, obtained on 27th October 1996, exhibits the smallest line ratio, while the largest ratio can be seen in the spectrum, taken on 16th March 1998.
Although the fitted effective temperatures are provided only as the estimated values, the temperature analysis yielded reliable constraints for the temperature range of the central star. Temperatures lower than 10500 K would result in the highly distorted HeI line due to the occurrence of the unresolved TiII line. On the other hand, the exponential line ratio dependence on temperature excludes temperatures above 13000 K. The measured minimum and maximum effective temperature constrain the spectral type of the central star to B9-B7. The measured average effective temperature 12000 K suggests spectral type B8 (Kenneth 1992).

3.4.2.4 Metallicity
The discrepancy between the synthetic and measured spectral line strengths (see Figs. 3.22, 3.24 and 3.26) can be attributed to the high metallicity of the central star. The high metallicity could result from dumped metal-rich material in a close binary system. Higher metallicity intensifies the metal absorption lines and has no effect on the hydrogen lines. The highest metallicity available in the applied library of synthetic spectra, was 0.5. Because the observed spectra indicate even higher values, the metallicity could not be fitted. This diminishes the reliability of the other determined atmospheric parameters. Nevertheless, the expected extrapolated impact of the higher metallicity on the selection of the other three parameters is relatively small. The atmospheric characterisation of the central star is expected to remain reliable within the provided error bars.

3.4.2.5 Combined Atmospheric Parameters
Table 3.5 lists the stellar atmospheric parameters of the synthetic spectrum, which yielded the best fit of the measured INT spectra. The comparison between the two spectra is given in Fig. 3.26.

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>( T_{\text{eff}} ) [K]</td>
<td>11000 – 13000</td>
</tr>
<tr>
<td>( \log(g) ) [cm/s(^2)]</td>
<td>4.0 ± 0.25</td>
</tr>
<tr>
<td>( v_{\text{rot}} \cdot \sin(i) ) [km/s]</td>
<td>200(^{+50}_{-25})</td>
</tr>
<tr>
<td>([\text{M/H}])</td>
<td>(\gg 0)</td>
</tr>
</tbody>
</table>

Table 3.5: Estimated atmospheric parameters of the central star. The effective temperature, stellar surface gravity and projected rotational velocity have been fitted with the synthetic spectra. Metallicity could not be fitted, because it exceeded the values, provided in the library of synthetic spectra.
The comparison between the measured and synthetic spectra indicates that the central star is a rapidly rotating metal-rich hot main sequence star and supports the contention that its spectral type is ~B8V, as originally proposed by Kohoutek (1979).

### 3.4.3 Standard Spectra

The measured spectra of the central star have been compared to a number of standard stellar spectra of various spectral types. The comparison was used to verify the provisionally determined B8V spectral type of the central star (see subsection 3.4.2). The standard spectra, i.e. spectra with no peculiarities that could have resulted in spectral misclassification, have been chosen from the spectral library, provided by Bagnulo et al. (2003). Figs. 3.27, 3.28 and 3.29 show the comparison between the measured INT spectrum of the central star and the standard spectra of spectral types A7V, B8V and B6V, respectively.
Figure 3.27: Comparison with a standard A7V spectrum. The comparison with the standard spectrum of Altair reveals many discrepancies. Note the absence of SiII lines (3856 and 3863 Å), HeI (4026 and 4471 Å) and presence of MnI (4033 Å), CaI (4227 Å) and TiII (4468 Å). In addition, the Balmer profiles of a main sequence A7 star exhibit too strong gravitation broadening.

Figure 3.28: Comparison with a standard B8V spectrum. In contrast to the standard A7 spectrum, B8 spectrum exhibits SiII and HeI lines, whereas TiII (4468 Å) is no longer present at higher temperatures of the B8 star. The measured depth ratio between the HeI (4471 Å) and MgII line (4481 Å) is perfectly matched with the standard B8V spectrum. Also note the comparable amount of gravitational broadening in Balmer absorption lines. Disregard the narrowness of other lines, due to the smaller projected rotational velocity of HD162586.
The measured spectra of the central star have been compared to many standard spectra, including the spectra of giants. The shown shortlisted spectral comparisons demonstrate that the spectral type of the central star is likely to be B8V. The average measured stellar parameters of a B8V star are:

- effective temperature \((11600 \pm 600) \text{ K}\) (Wu et al. 2011)
- \(\log(g)\) \((4.0 \pm 0.1) \text{ cm/s}^2\) (Wu et al. 2011)
- mass \(~2.6 \, M_\odot\) (Belikov 1995)
- radius \((2.9 \pm 0.8) \, R_\odot\) (Pasinetti-Fracassini et al. 2001)
- equatorial rotational velocity \(~250 \, \text{km/s}\) (Glebocki et al. 2005)

Each of these values was derived from the referenced catalogues, all of which can be accessed through VizieR. The similarity between the average stellar parameters and the parameters obtained by the synthetic spectra (see Table 3.5), justify the B8V classification of the central star.
The B8V classification refutes the spectral type A7V-F0V, which was proposed by Cuesta et al. (1993). Their spectral classification was based on the calculated effective temperatures of the binary components, required to reproduce the measured spectral flux in the wavelength range 6000-7400 Å. For the computation of the effective temperature, they were using an assumed value of the core extinction coefficient. In addition to likely underestimated extinction coefficient, Cuesta et al. (1993) were also using the slightly underestimated distance of 0.7 kpc. More recent and more accurate measurements place Sh2-71 at a distance of 1.0 kpc (Stanghellini et al. 2008, also see Table 1.1). Cuesta noted, however, the presence of SiII lines (6347 and 6371 Å) in their spectra, which correspond to stars with temperatures higher than A7V. The spectral wavelength range, covered by this project’s INT spectra, reveals many more spectral features that cannot be explained by an A7V star. The synthetic spectra fitting and comparison with the standard spectra provide strong evidence that the spectral type of the central star is ~B8V.
4 Possible Models

This chapter presents possible astrophysical causes of the photometric and spectroscopic properties of the observed star in an attempt to produce a realistic model. Since the star is believed to be the central star of the PN, the model should also provide plausible conditions for the formation of the bipolar PN Sh2-71.

4.1 Cataclysmic Binary

A high level of nebular ionisation led astronomers to a suggestion that the central star of the PN Sh2-71 is a binary star between a hot ionising star and optically more luminous main sequence companion star (Kohoutek 1979). A binary suggestion was further developed by Cuesta et al. (1993), who proposed a detailed cataclysmic binary model to explain the observed kinematic properties of the nebular core. According to their model, the main-sequence star of a semidetached binary overflows the Lagrangian points $L_1$ and $L_2$. The resulting accretion disc around the ionising star was suggested to be responsible for the observed Balmer emission in the spectra of the central star, while the $L_2$ overflown material could represent the collimation mechanism. As it is unlikely that Keplerian rotation of the accretion disc could produce the distinct double peaked profiles and the observed widths of the Balmer lines (>1000 km/s), Cuesta et al. (1993) proposed that the Balmer emission originates from the two collimated high-velocity outflowing jets (see Fig. 4.1).

The model, presented by Cuesta et al. (1993), provides a plausible explanation for the observed spectral line’s profiles as well as for the bipolar shape of the PN. However, a small separation between the stars, required for mass transfer to establish between a B8V and an ionising star, would correspond to an orbital period of less than 1.0 day. If the mass of a B8V star is assumed to be 2.6 $M_\odot$ (see subsection 3.4.3) and the mass of a stellar remnant of a recently more massive star is 0.68 $M_\odot$ (see equations (1) and (2)), then the consequent RV variations of stellar absorption lines would be $\pm 65 \cdot \sin(i)$ km/s, whereas the RV of Balmer emission components would vary with $\pm 250 \cdot \sin(i)$ km/s. The overall shape of the PN implies that the inclination of the binary
Figure 4.1: Cataclysmic binary model, proposed by Cuesta et al. (1993). A main-sequence companion star overflows L\textsubscript{1} and establishes a mass transfer towards the ionising star. The inflowing material with electron scattering opacity becomes ejected by the ionising star's radiation pressure. The ejected high-velocity outflow is then collimated by the slowly expanding collimating disk, which results from the L\textsubscript{2} overflow. Credit: Cuesta et al. (1993), page 391.

The system is large (Cuesta et al. 1993, also see Fig. 1.10). Therefore, according to the cataclysmic binary model and the identified B8V spectral type of a companion star, the detected ±40 km/s RV variations (see Figs. 3.18 and 3.20) might be explained with cataclysmic binary model. However, the AAT spectra do not reveal any noticeable spectral line profile or RV variations of Balmer emission lines in the 4.5 hour timespan of spectroscopic monitoring (see Fig. 3.7), which implies that the short orbital period is unlikely. Even if the mass of an ionising star is expected to be much larger, say 1 M\textsubscript{⊙}, we should still see a significant portion of ±230 \cdot \sin(i) km/s RV shifts of Balmer emission lines and the corresponding variations of the profile shapes. In addition, the short orbital period, required for the cataclysmic model, could not explain the 68 day brightness variation period. On the other hand, the eventual 68 day periodic heating and cooling of the accreting star due to the variable mass transfer (Gänsicke 1999) could not reproduce the observed stability of the lightcurve period and amplitude. Even if the ionising star is assumed to possess a strong magnetic field (10\textsuperscript{3} T) and that the accretion stream is diverted into a pair of magnetically-driven high-velocity inflowing polar streams (Cropper 1990), which might reproduce a double peaked Balmer profile, the above mentioned discrepancies remain.
The cataclysmic model, introduced by Cuesta et al. (1993), contradicts some of the observed spectroscopic properties and does not provide a plausible explanation for the 68 day variation period. The new spectroscopic results, presented by this thesis, prove that the cataclysmic model is highly unlikely.

4.2 Be Binary

A Be type star is “a non-supergiant B-type star whose spectrum has, or had at some time, one or more Balmer lines in emission” (Collins 1987). Because Be stars rotate rapidly, the gravity becomes balanced by the centrifugal force in the equatorial regions. Then mild prominence activity or other minor disturbances lead to an equatorial ejection of matter, which drifts into an extended circumstellar disc (Stoeckley 1968). The excited decretion disc gives rise to Balmer and singly-ionised metal emission lines, observed in the spectra of Be stars (Kogure et al. 1982). The mechanism responsible for the spectral emission components is illustrated in Fig. 4.2.

![Figure 4.2: Schematic structure of spectral profiles of Be type stars. Figure on the left shows a pole-on Be star and its circumstellar disc. The corresponding spectral profile components are plotted on the right. If the Be system is observed close to edge-on, the disc occults the star and introduces a narrow absorption component in the spectral line’s profile (2). On the other hand, the receding and approaching part of the rotating disc result in the red (3) and blueshifted (1) emission components. The shape of the spectral profile depends on the inclination of the disc. The broad absorption feature is gravitationally broadened stellar absorption component. \( \Delta \lambda_{\text{basal}} \), \( \Delta \lambda_{\text{FWHM}} \), and \( \Delta \lambda_{\text{peak}} \) are basal width, FWHM and peak separation of the spectral line, respectively. Adapted from Kogure et al. (1982), page 284 and Slettebak (1976), page 131.](image-url)
The characteristic widths of the spectral lines, introduced in Fig. 4.2, are related to the rotational velocity of the Be star. Hanuschik (1989) analysed the correlation between circumstellar emission lines and projected rotational velocities, obtained from the widths of photospheric absorption lines for 115 Be stars. He found that the correlation was most reliable for emission lines’ FWHMs ($\Delta\lambda_{\text{FWHM}}$). On the other hand, the correlation with basal widths ($\Delta\lambda_{\text{basal}}$) hardly produced any correlation due to electron scattering, whereas the peak separation ($\Delta\lambda_{\text{peak}}$) yielded similar results as $\Delta\lambda_{\text{FWHM}}$ but also required the measurements of the line’s equivalent widths. The Hα FWHM dependence on the projected rotational velocity is given in equation (8), whereas equation (9) is provided for Hβ and FeII emission lines (Hanuschik 1989):

$$\frac{\Delta\lambda_{\text{FWHM}}(\text{H}\alpha) \cdot c}{\lambda_{\text{H}\alpha}} = 1.4 \cdot v_{\text{rot}} \sin(i) + 50 \text{km/s}$$

(8)

$$\frac{\Delta\lambda_{\text{FWHM}}(\text{H}\beta, \text{FeII}) \cdot c}{\lambda_{\text{H}\beta, \text{FeII}}} = 1.2 \cdot v_{\text{rot}} \sin(i) + 30 \text{km/s}$$

(9)

where the expression on the left represents the FWHM of the emission line in velocity units.

The average FWHM of the Hα and Hβ profiles, measured in the spectra of LT and AAT spectroscopic datasets, are 9.3 and 5.7 Å, respectively. According to equations (8) and (9), the corresponding projected rotational velocity is $\sim$270 km/s. Because this value does not differ significantly from the fitted (200$^{+50}_{-25}$ km/s) and estimated projected rotational velocity ($\sim$270 km/s) (see subsection 3.4.2.1), the circumstellar Be disc could be the possible progenitor of the observed emission lines in the spectra of the Sh2-71’s central star.

It has been reported that a significant fraction of Be stars exhibit a certain degree of spectral line profile variability due to many potential contributing factors, e.g. variable wind absorption, nonradial stellar pulsations, starspots, magnetically modulated circumstellar rotation, etc. (Porter et al. 2003). Beside the spectral variability, typical Be stars can also exhibit brightness variability. However, the largest reported brightness variations are irregular and of the order of 0.1$^m$ or less (Feinstein 1968). The observed periodic 0.8$^m$ brightness variations of the central star could thus be explained only with an additional mechanism. The following discussion is firstly provided for a warped circumstellar disc due to the inclined magnetic field (see Figs. 4.3 and 4.4). The second
possibility is a precession of a disc due to the misalignment between binary and disc rotation axes (the example is V725 Tau by Haigh et al. (2004), general discussion by Martin et al. (2011)).

The magnetically warped circumstellar disc brightness variation mechanism can be illustrated with the example of AA Tau (a T Tauri type star), whose lightcurve is shown in Fig. 4.3.

![Image](image_url)

Figure 4.3: V lightcurve of AA Tau. The lightcurve is a result of a rotating magnetically warped circumstellar disc. Credit: Bouvier et al. (2013), page 4.

The suggested mechanism for the observed photometric and spectroscopic properties of AA Tau is a periodic occultation of the star by the corotating magnetically warped disc (see Fig. 4.4). The disc is warped due to the misalignment between the magnetic field and stellar rotation (O'Sullivan et al. 2005).

![Image](image_url)

Figure 4.4: The periodical occultation of the star by the corotating magnetically warped circumstellar disc. The images were produced with a computer simulation. Credit: O'Sullivan et al. (2005), page 637.

Although this mechanism could explain the amplitude and the lightcurve shape of the Sh2-71’s central star, it remains unclear how could a rapidly rotating B8Ve star induce a much longer 68 day disc corotation period. In addition, in the case of AA Tau the
calculated strength of the magnetic field was 5.2 kG (O’Sullivan et al. 2005), whereas the suggested upper limit for Be stars is \( \sim 1 \) kG (Porter et al. 2003). Therefore, the magnetically warped Be star disc is not a possible explanation of the observed central star’s lightcurve.

The second, more likely explanation for the observed photometric properties is a binary system between a B8Ve and a companion star. The reported fraction of B8e among all B8 stars is 10\%, vast majority of them being main sequence stars (Kogure et al. 1982), whereas the binary fraction of Be stars is \( \sim 1/3 \) (Porter et al. 2003). Combined fractions suggest that about 3\% B stars are B8Ve binaries, which yields a reasonably high probability that the central star of Sh2-71 could be a B8Ve binary.

An example of a Be binary is \( \phi \) Per, a binary system between a B0.5IVe and an O type subdwarf (sdO6) (Božić et al. 1995). In the past, the system comprised a normal B type secondary star and a more massive and more luminous primary star. As the primary star exhausted the hydrogen fuel in its core, it expanded into a red giant. Since the stars formed a close binary system, the red giant star filled its Roche lobe and the mass transfer to the B type star established. Eventually, the entire envelope of the red giant star was transferred to the B star and exposed the core of the red giant. The exposed He-shell burning core of the former red giant star has thus transformed into a subdwarf secondary star, whereas the accreted red giant’s excess mass transformed the normal B star into a Be type primary star (see Fig. 4.5).

Figure 4.5: The formation mechanism for the close Be binary system \( \phi \) Per. Credit: Schombert (2013).
As the Be binary system rotates, the profiles of spectral lines undergo pronounced variations, which are correlated with the orbital phase (see Fig. 4.6).

Figure 4.6: Spectral variations of the Be binary φ Per. The spectral profiles of Hα are shown for two different orbital phases on the left, while the spectra on the right show the Hβ profile (Be Star Spectra Database).

The Be binary mechanism, demonstrated with the example of φ Per, could explain the pronounced spectral variations, observed in the spectra of Sh2-71’s central star (see Figs. 3.12 and 3.15). A close Be binary system could also result in brightness variations due to various contributing factors, e.g. irradiation, tidal distortions, reflection, or even eclipses. However, according to Fig. 3.6, the spectral variations are not correlated with the 68 days brightness variation phase. Therefore, a plausible mechanism should also provide an explanation for the uncorrelated periodic 0.8° brightness variability.

Martin et al. (2011) showed that a misalignment between a binary orbital plane and plane of a Be disc can invoke tidal torques, which cause the precession of the Be disc. The intensity of the torques depends on the stellar masses, separation between the stars and the eccentricity of the binary system (Martin et al. 2011). An example of the binary-invoked Be disc precession is V725 Tau. In this example, the Be disc precession is manifested in periodic quasi-sinusoidal 0.4 V magnitude variations with a period of 1400 days, whereas the orbital period is 110 days (Haigh et al. 2004). Due to a different binary system of Sh2-71, possibly large eccentricity and much shorter orbital period, the
precessing disc might be responsible for the observed 68 days brightness variations. If the system is seen close to edge-on, then a doubled, 136 days, precession period could also theoretically result in a sinusoidal 68 day lightcurve.

Supporting evidence for the precessing disc is the detected variation of the reddening effect, which coincidences with the period and phase of the central star’s lightcurve (see Fig. 2.19 and equation (6)). The absorption effect, corresponding to the measured variable component of the reddening, is estimated by equation (10):

\[
A_V = R_V \cdot \frac{E(B-V)}{E(U-V)} \cdot E(U-V) = 0.72
\]

where \(A_V\) is the variable component of extinction for V band, \(E(B-V)/E(U-V) = 0.61\) is the published mean colour excess ratio for conversion between U and B filters (Rieke et al. 1985), \(E(U-V) = 0.382\) is the measured reddening (see equation (6)) and \(R_V = 3.1\) is the typical ratio \(A_V/E(B-V)\) for the Galactic extinction (Rieke et al. 1985). The estimated extinction, derived from the reddening, provides roughly the same amplitude of brightness variability, as measured in the V lightcurve of the central star (0.8\(m\)). Thus, the gas and dust extinction due to the precessing Be disc could be a realistic explanation for the central star’s photometric behaviour.

Spectrophotometric properties similar to Sh2-71 have not been reported for any other PN. However, close binary central stars which have undergone mass transfer have been suggested as progenitors of many other PNe. A binary formation mechanism similar to \(\phi\) Per (see Fig. 4.5) has been suggested for the central star of bipolar PN NGC 2346. The former mass interaction between the close binary components resulted in the equatorial expulsion of the red giant’s envelope, which provided the collimating mechanism for the binary shape of the PN (HubbleSite 1999) (see Fig. 4.7). The current central star of NGC 2346 is a binary system between A5V and O type subdwarf with an orbital period of 16 days (Smalley 1997). If the same binary formation mechanism is applied to the central star of Sh2-71, the expected maximum orbital period would be \(~20\) days. This is the estimated current orbital period, which would enable the mass transfer in the pre-PN phase between a red giant star and a B8V star. The proposed Be binary formation mechanism similar to \(\phi\) Per therefore explains the presence of a Be disc, it provides the required collimation for the bipolar shape of Sh2-71, and also explains the observed high metallicities in the spectra of the central star (see subsection 3.4.2.4). The metal-rich material, recently dumped onto the B8V star in the pre-PN
phase would not have had enough time to be fully mixed into the primary star (Smalley 1997).

Figure 4.7: Bipolar PN NGC 2346. The bipolar shape originates from the fast stellar wind, which is collimated by the equatorially expelled common envelope material. The central star is believed to be a close binary A5V-sdO with the orbital period of 16 days (Smalley 1997). Credit: NASA and The Hubble Heritage Team (AURA/STScI).

The proposed precessing Be disc model is highly tentative without further spectroscopic data. However, it provides the explanation for the observed brightness variations and uncorrelated pronounced spectral variability. In addition, the mechanism also provides a potential collimation mechanism, required for the bipolar shape of Sh2-71. An extended high-cadence sample of Hα and Hβ profiles could verify any potential periodicity of the spectral variations and thus confirm or reject the Be binary hypothesis.

### 4.3 Herbig Be

Herbig Be stars are young non-supergiant B type stars, which are still embedded in embryonic envelopes and often surrounded by accretion discs (Mannings 1995). Circumstellar material, excited by the central B star, is responsible for the observed
emission spectral lines in the spectra of Herbig Be stars. Circumstellar emission and absorption from the Herbig accretion discs can result in very similar spectral line profile shapes and variability as Be decretion discs (Harrington et al. 2009) (see Fig. 4.2).

The similarities between Be and Herbig Be stars lead to reconsideration of magnetically warped circumstellar disc model, which could explain 68 days brightness variations. B type stars do not posses deep outer convection zones, which are required for magnetohydrodynamic dynamo to establish magnetic fields. The discovered presence of magnetic fields in these hot stars led scientists to the assumption of fossil magnetic fields – passively decaying magnetic fields, which were formed or swept up in the stellar formation processes (Wade et al. 2007). Herbig Be stars are among the youngest B type stellar objects, with ages up to 10 Myr (Waters et al. 1998). Therefore, their magnetic fields are expected to be stronger than for Be stars, and are hypothesised to reach up to few $10^4$ G (Wade et al. 2007).

In contrast to Be stars, the stronger magnetic fields of Herbig Be stars could result in the corotating magnetically warped disc, analogue to the model presented by O’Sullivan et al. (2005) (see subsection 4.2). This mechanism could explain the observed sinusoidal 0.8th brightness variations and the uncorrelated pronounced spectral variations. However, because stellar rotational periods of Herbig Be stars are considerably shorter than 68 days (Waters et al. 1998), the effect of a corotating disc does not provide a plausible explanation for the observed 68 days photometric variations. In addition, Herbig Be model is not consistent with the assumption that the observed star is the central star of the PN because a young stellar object cannot form a binary system with a much more evolved progenitor of the PN.

**4.4 Other Models**

Various other models have been considered in trying to explain the observed spectrophotometric properties of the Sh2-71’s central star.

Radial stellar pulsations might provide the required lightcurve (De Cat 2002), but would also result in the correlated RV and spectral profile shape variations. In addition, pulsations alone cannot explain the emission lines in the spectra of the central star.

Stellar rotation and the presence of starspots can be ruled out as a progenitor of the observed brightness variations, since the 68 period greatly exceeds the estimated
orbital period of a rapidly rotating B8V star. In addition, rotating starspots could not reproduce the amplitude of 0.8 m and its constancy.

Irradiation, reflection, tidal distortion and possibly even eclipses in a close binary system may all contribute to the brightness variability. Although the shape and the amplitude of the lightcurve can be reproduced for the shortest orbital periods with the various binary effects, it is impossible to reproduce the same lightcurve for a 68 day period. A more likely close binary brightness variation mechanism is variable optical depth of the circumstellar material due to the orbital motion of a binary central star with the orbital period of 68 days (Mikulašek et al. 2007). However, the orbiting close binary is expected to exhibit noticeable spectral variations, which should be correlated with the brightness phase. This contradicts to the finding represented in Fig. 3.6, which rejects the 68 days spectral variability period. An orbital period of 68 days is further refuted by the RV measurements, which do not exhibit any correlation with the brightness phase (see Figs. 3.18 and 3.20).

A triple star system is also a potential contributing factor. But without further data, the triple star system remains only a hypothesized possibility.

None of the models, discussed in this and previous subsections provides a reliable explanation for the spectrophotometric behaviour of the central star. However, the observations are best reproduced with a close Be binary system and a misaligned precessing disc (see subsection 4.2).

Further high-cadence (~1 day) spectroscopic data for the spectral region Hα and Hβ could reveal any potential periodicity of the spectral features and RV variations, implied by several presented mechanisms. In addition, high-cadence and higher S/N photometry is required to unambiguously confirm the presence of possible 17.2 days eclipsing events. Until more data are obtained, the suggested model of a precessing disc in a Be binary system should be regarded as tentative.

### 4.5 Different Central Star

If the observed star is the central star of the PN Sh2-71, then a model for the observed spectrophotometric properties would have to be constrained by the presence of the progenitor’s stellar remnant. However, due to the other field stars near the centre of the PN, it is necessary to consider the possibility that the observed star is actually not the central star of Sh2-71. Frew et al. (2007) suggested a much dimmer 19th star to be the central star (see Fig. 4.8). They claim that the measured brightness of this star
corresponds to the expected luminosity of the progenitor’s remnant at the nebula’s distance of 1 kpc (Michaud 2012). In addition, their suggested central star is also closer to the apparent centre of the nebula, whereas the observed and currently accepted central star exhibits a slight offset (Frew et al. 2007).

Figure 4.8: An alternative suggestion for the central star of Sh2-71. The marked 19th magnitude star is suggested to be the central star (Frew et al. 2007). The currently accepted central star of Sh2-71 is a much brighter star, 7" SE from the dim star. Image is a zoomed version of Fig. 1.10.

The central position of the dim star favours its identification as the central star of the PN. However, the somewhat symmetrical nebular structure about a line eastwards from the bright star (see Fig. 4.8) favours this star to continue to be considered as the central star. The photometric claim, provided by Frew et al., could be challenged by the fact that the most likely collimating mechanism for the bipolar shape of the PN is a binary central star. Therefore, it is reasonable to expect that the binary companion would provide additional brightness contribution. An additional indirect supporting evidence for the bright central star is the near agreement between the reported average RV of the PN and measured mean stellar RV (see subsection 3.4.1 and Fig. 3.19).

Nevertheless, even if the observed star is not the central star of the PN, its unusual spectrophotometric behaviour implies that the star itself is an intriguing object, which urges for further observations to be fully understood.
5 Conclusion

This thesis presented photometric and spectroscopic observations of the unusual central star of planetary nebula (PN) Sh2-71. The aim was to verify the presence of assumed eclipsing events and to find any potential correlation between the spectral variations and the brightness phase of the central star. The observational results were applied in the discussion on the different possible models for the origin of the observed properties.

The photometric images were obtained by Byrne Observatory at Sedgwick (BOS), Faulkes Telescope South (FTS) and Liverpool Telescope (LT). Each of the three different datasets was reduced manually with Starlink data reduction software. The differential aperture photometry yielded a combined lightcurve, which was in the agreement with the reported ephemeris for the sinusoidal 0.8 V magnitude brightness variations with a period of 68 days. In addition, the manual data reduction confirmed the presence of two brightness dips, previously seen in preliminary BOS lightcurve obtained by the automatic data reduction process. Furthermore, the presence of three additional brightness dips in the LT lightcurve implied that the dips are occurring periodically with the period of 17.2 days. The comparison between U and V LT lightcurves revealed that the 68 days brightness variations are accompanied by the variable reddening effect.

The LT spectroscopic dataset, purposely obtained for this research project, was expanded by 25 archived Isaac Newton Telescope (INT) spectra and with an archived set of high-cadence Anglo-Australian Telescope (AAT) spectra of the central star. Manual data reduction of INT and AAT spectra and the automatically reduced LT spectra exhibited pronounced spectral variations, particularly the Balmer line profiles. The comparison between the spectra, corresponding to different brightness phase, revealed that spectral variations are not correlated with the variable brightness. On the other hand, eight high-resolution AAT spectra with 30 minutes cadence did not show any noticeable spectral changes in the 4.5 hour timespan. Combined spectroscopic results suggested that the spectral variations must occur on the timescale of few days. The measured radial velocity (RV) span of ±40 km/s was not correlated with the 68 day brightness variation period. The agreement between the measured average stellar RV of ~26 km/s and the reported mean RV of the PN 24.7 km/s, provided supporting evidence that the observed star is indeed the central star of the PN. A comparison between the
measured and synthetic spectra yielded stellar atmospheric parameters $T_{\text{eff}}$ 12000 K, log($g$) 4.0 cm/s$^2$, $v_{\text{rot}} \cdot \sin(i)$ 200 km/s and indicated a high value of metallicity. The comparison with standard stellar spectra suggested that the spectral type of the central star is ~B8V, which is supported by the fitted atmospheric parameters.

Spectrophotometric properties were applied in the discussion on the possible models for the central star. The previously proposed cataclysmic binary model was revisited. The required $\lesssim$1 day orbital period for the cataclysmic binary is expected to exhibit pronounced spectral shape and RV variations on the similar, $\lesssim$1 day, timescale. The short timescale periodicity of INT RV measurements could not be verified due to undersampled data. However, the detected stability of Balmer profiles in AAT spectra rejected the cataclysmic binary model. From the various potential models, discussed in this thesis, the observed properties of the central star were best reproduced with a precessing Be disc in a misaligned close binary system. Periodic disc occultations were indirectly supported by the variable reddening effect. The model also provides the necessary collimation for the formation of a bipolar shape of the PN. However, due to the undersampled spectroscopic data, the anticipated periodicity of the RV variations could not be reliably verified. In addition, the lack of spectra with H$\alpha$ and H$\beta$ coverage prevented an accurate analysis of the Balmer emission variability. Therefore, the suggested model for the central star should be regarded as tentative.

Future photometric observations should aim to confirm the indicated 17.2 days eclipsing period. An additional LT V photometric dataset, targeted on the extrapolated positions of the brightness dips and possibly with a higher signal to noise ratio, could provide a reliable verification of the proposed periodicity. However, in order to model the potential eclipses, a much higher cadence would be required, in particular for the purposes of the accurate subtraction of the quasi-sinusoidal brightness variations. For this reason, SuperWASP observations will hopefully be obtained shortly. Alternatively, the accurate determination of the lightcurve shape could also be provided with 50 observing nights of at least every second day cadence with a 30 s exposure in V filter with the LT or equivalent. Three times larger signal to noise ratio with a 2 day cadence, coupled with the hourly cadence at the expected eclipsing events could enable the definitive eclipsing analysis.

The spectroscopic dataset presented in this thesis was relatively large. However, to determine any potential periodicity of the spectral profile and RV variations, a cadence of one day would be desirable. In addition, future spectroscopic observations
should aim to cover the H\textalpha{} and H\beta lines, since the variability of the emission components can be most easily examined for the lowest Balmer orders.

Despite relatively modest photometric dataset, incompatibilities between filters and spectral wavelength ranges, it was possible to extract a number of useful parameters for the central star and to discuss the possible models. Above all, the thesis has shown that the central star of the PN Sh2-71 is an intriguing object that deserves further observations to be fully understood. Improving our knowledge of every central star from a handful set of analysed progenitors of bipolar PNe, can help us to better understand the shaping mechanisms.
References


