The orientation of the \( \eta \) Carinae binary system

Amit Kashi* and Noam Soker*

Department of Physics, Technion-Israel Institute of Technology, Haifa 32000, Israel

Accepted 2008 August 24. Received 2008 August 15; in original form 2008 March 30

ABSTRACT

We examine a variety of observations that shed light on the orientation of the semimajor axis of the \( \eta \) Carinae massive binary system. Under several assumptions, we study the following observations: the Doppler shifts of some \( \text{He}\,\text{ii}\ \lambda\,P\ \text{Cygni} \) lines that are attributed to the secondary’s wind, of one \( \text{Fe}\,\text{ii} \) line that is attributed to the primary’s wind and of the Paschen emission lines which are attributed to the shocked primary’s wind, are computed in our model and are compared with observations. We compute the hydrogen column density towards the binary system in our model, and find a good agreement with that deduced from X-ray observations. We calculate the ionization of the surrounding gas blobs by the radiation of the hotter secondary star, and compare with observations of a highly excited \([\text{Ar}\,\text{iii}]\) narrow line. We find that all of these support an orientation where for most of the time the secondary – the hotter, less-massive star – is behind the primary star. The secondary comes closer to the observer only for a short time near periastron passage in its highly eccentric (\( e \approx 0.9 \)) orbit. Further supporting arguments are also listed, followed by a discussion on some open and complicated issues.

Key words: binaries: general – stars: individual: \( \eta \) Car – stars: mass loss – stars: winds, outflows.

1 INTRODUCTION

The \( P = 5.54 \) yr \((P = 222.7 \pm 1.3 \) d; Damineli et al. 2008a\)) periodicity of the massive binary system \( \eta \) Car is observed in the radio (Duncan & White 2003), infrared (IR) (Whitelock et al. 2004), visible (e.g. van Genderen et al. 2006) X-ray (Corcoran 2005) and in many emission and absorption lines (e.g. Damineli et al. 2008a,b). According to most of the models, the periodicity, for example of the spectroscopic event and the X-ray minimum, follows the 5.54 yr periodic change in the orbital separation in this highly eccentric, \( e \approx 0.9 \), binary system (e.g. Hillier et al. 2006). The spectroscopic event is defined by the fading, or even the disappearance, of high-ionization emission lines (e.g. Zanella, Wolf & Stahl 1984; Damineli 1996; Damineli et al. 1998; Damineli et al. 2000, 2008a,b). The rapid changes in the continuum, lines and X-ray properties (e.g. Corcoran 2005; Davidson et al. 2005; Martin et al. 2006a,b; van Genderen et al. 2006; Nielsen et al. 2007a; Damineli et al. 2008b) are assumed to occur near periastron passages, although less rapid variations occur along the entire orbit.

It is generally agreed that the orbital plane lies in the equatorial plane of the bipolar structure – the Homunculus (Davidson et al. 2001). The inclination angle (the angle between a line perpendicular to the orbital plane and the line of sight) is \( i \approx 45^\circ \), with \( i = 41^\circ–43^\circ \) being a popular value (Davidson et al. 2001; Smith 2002). However, there is a disagreement about the orientation of the semimajor axis in the orbital plane – the periastron longitude. We will use the commonly used periastron longitude angle \( \omega; \omega = 0^\circ \) for a case when the secondary is towards the observer at an orbital angle of \( 90^\circ \) after periastron, \( \omega = 90^\circ \) for a case when the secondary is towards the observer at periastron, \( \omega = 180^\circ \) for a case when the secondary is towards the observer at an orbital angle \( 90^\circ \) before periastron and \( \omega = 270^\circ \) for a case when the secondary is towards the observer at apastron and so on.

While some groups argue that the secondary (less massive) star is away from us during periastron passages, \( \omega = 270^\circ \) (e.g. Nielsen et al. 2007a; Damineli et al. 2008b), others argue that the secondary is towards us during periastron passages, \( \omega = 90^\circ \) [Abraham et al. 2005; Falceta-Gonçalves, Jatenco-Pereira & Abraham 2005; Kashi & Soker 2007b (hereafter KS07), who use the angle \( \gamma = 90^\circ – \omega \)]. Other semimajor axis orientations have also been proposed (Davidson 1997; Smith et al. 2004; Dorland 2007; Henley et al. 2008; Okazaki et al. 2008a). Abraham & Falceta-Gonçalves (2007) have obtained the orientation angle in the range \( \omega = 60^\circ–90^\circ \), independent of the orbital inclination. They did not assume that the binary and the Homunculus orbital plane coincide. This contradicts the binary interacting model that we support. In addition, Falceta-Gonçalves et al. (2005) suggested a model for the X-ray emission with the usage of a problematic expression for the X-ray emission (see Akashi, Soker & Behar 2006). We therefore will not use the arguments of these two papers to support our claim for \( \omega \approx 90^\circ \).

In the present paper, we consider all observations that can shed light on the periastron longitude. As will be shown, all of them
support a value of $\omega \simeq 90^\circ$. Namely, the secondary is behind the primary most of the time and passes in front of the primary for a short time near periastrom passage. In Section 2 we discuss the Doppler shifts of several lines; in Section 3 we discuss the hydrogen column density as deduced from X-ray observations and in Section 4 we discuss the variation in the intensity of high-excitation narrow lines. Our discussion and predictions are given in Section 5.

Our assumptions concerning the location where some lines are produced are different from those assumed by other research groups. We therefore present in Fig. 1 a map of the origins of different lines according to our assumptions. This figure should be consulted when reading the following sections.

2 THE DOPPLER SHIFTS AS ARISING FROM ORBITAL MOTION

2.1 The binary system

In applying the orbital motion, we will use two sets of binary parameters. In both, the orbital period is taken to be 2024 d (in a recent paper the orbital period was found to be 2021–2024 d; Damineli et al. 2008a), and the eccentricity is very high, $e \sim 0.9$. The terminal winds’ velocities are taken to be $v_1 = 500 \text{ km s}^{-1}$ and $v_2 = 3000 \text{ km s}^{-1}$, where subscripts 1 and 2 stand for the primary and the secondary, respectively. The first binary model is based on the commonly used binary parameters as taken from many papers (e.g. Ishibashi et al. 1999; Damineli et al. 2000; Corcoran et al. 2001, 2004; Hillier et al. 2001; Pittard & Corcoran 2002; Smith et al. 2004; Verner, Bruhweiler & Gull 2005; Dorland 2007; Damineli et al. 2008a,b). This is termed by us (Kashi & Soker 2009, hereafter KS09) as the ‘small-masses model’. The assumed stellar masses are $M_1 = 120 \text{ M}_\odot$ and $M_2 = 30 \text{ M}_\odot$. The calculated semi-major axis is $a = 16.64 \text{ au}$ and the orbital separation at periastrom is $r = 1.66 - 1.16 \text{ au}$ for $e = 0.9$–0.93, respectively.

The masses of the small-masses model are fixed more or less by the assumption that each star is at its Eddington luminosity limit. However, fitting to evolution of main-sequence stars gives much higher masses. This model is termed as the ‘big-masses model’. Following KS09, we take masses that come from evolutionary considerations (see also Damineli et al. 2000), and are in the ranges $M_1 \simeq 150$–$180 \text{ M}_\odot$, as the primary has lost several tens of solar masses and $M_2 \simeq 60$–$70 \text{ M}_\odot$.

In KS07, we fixed the stellar masses and varied the inclination angle $i$ and the periastrom longitude $\omega$ (using the angle $\gamma = 90^\circ - \omega$). In the present paper, we fix $i = 45^\circ$, $e = 0.9$ or 0.93 and $\omega = 90^\circ$ (secondary towards us at periastrom), and use two sets of masses: $(M_1, M_2) = (120, 30)$ and $(160, 60) \text{ M}_\odot$. We will not play at all with the parameters, hence the fitting will not be perfect. Small variations of the parameters $i$, $\omega$, $M_1$, $M_2$ and $e$ can easily give better fits. But until we know better the values of these parameters by other means, it is pointless to play with these five parameters. What we do stress is that the orbital motion can explain the observed Doppler shifts.

2.2 Observations supporting orbital motion interpretation

There are several arguments supporting the interpretation of the periodic variations in the Doppler shifts as arising from the orbital motion. However, the interpretation is not straightforward. (also see attributed the Doppler shift of the hydrogen emission Paschen lines to the orbital motion of the primary star. Using this interpretation, they found the eccentricity to be $e = 0.75$, much lower than the now commonly assumed value of $e \simeq 0.9$; Davidson (1997) found $e \simeq 0.7$–0.8. In addition, this amplitude of the Doppler shift if originating in one of the stars is not compatible with other Doppler shifts as we discuss in this paper. Alternatively, we will attribute the Paschen line emission to the shocked primary’s wind near the stagnation point of the colliding winds.

By attributing some of the He I P Cygni lines to the secondary – the less massive star – we were able to explain the Doppler periodicity with the commonly used binary parameters (KS07). We consider this as a strong support to the orbital motion interpretation and devote Section 2.3 to study this in more detail. There we also show that the Fe II 6455 might be coming from the primary’s wind. If our interpretation is correct, $\eta$ Car is a double-lined spectroscopic.

Figure 1. A schematic map of the $\eta$ Car system at apastron. The origins of the different lines are marked on the map. According to our suggested model ($\omega = 90^\circ$), the observer is on the left-hand side of the system.
binary. We note that our view that some of the He i lines originate in the secondary’s wind is in dispute, as most researchers in the field attribute the He i P Cygni lines to other sources, either the primary or the winds collision region (Falca et al., 2007; Humphreys, Davidson & Koppelman, 2008). Our reply to some of the criticisms of Humphreys et al. (2008) is given in section 5 of KS07, where we also discuss why attributing the He i lines to the primary’s wind is problematic. In particular, if the change in the Doppler shift of the absorbed part was due to the change in the distance of the wind from the primary surface where the lines are formed, hence a change in the wind velocity, we would expect the lines emission width to change accordingly, and its centre to stay at the stellar primary velocity. This is contrary to observations. Take the He i λ7067 line for example, as reported by Nielsen et al. (2007a), the centre of the emission part follows the minimum in the line profile, and its width does not change much. This is what is expected if the entire source of the line changes its velocity, as in the orbital interpretation of the Doppler shifts.

In our model, the emission of the λ7065, λ5876 and λ4471 He i lines is formed in the secondary’s wind. The secondary luminosity of ~20 per cent of the primary luminosity seems to be enough for forming these lines (see KS07). We attribute the absorption in these lines – the P Cygni profile – to the secondary wind. The absorption is from the emission in this line. Indeed, the minimum of the absorption in the λ7065 and λ5876 lines is clearly above the continuum of the η Car spectrum (Hillier et al. 2001). In the case of the He i λ4471 line, it is less clear. However, we note that the entire region from 4400 to 4500 Å is full of lines (e.g. Pereira et al. 2008) that form a region above the real continuum. Some of these lines are formed in the secondary star. It is therefore possible that the continuum is actually 10 per cent lower than the value used by Hillier et al. (2001), and that the He i λ4471 line has its absorption above the continuum as well. We conclude that even though the helium lines show strong absorption, it is still above the continuum, and therefore it is possible that the lines are emitted and absorbed in the secondary’s wind, as we suggest here.

Another support to the orbital motion interpretation might come from the silicon and sulphur X-ray lines. If the lines are taken to be formed where the fast secondary’s wind is shocked, then their Doppler shift variations with orbital phase can be explained by using the same orbital parameters as used to explain the Doppler shifts of the He i lines (Behar et al. 2007). Henely et al. (2008) studied the possibility that the Doppler shifts of these X-ray lines result mainly from the outflow velocity of the shocked secondary’s wind. As can be seen from their fig. 18, this model fails. Although the outflow velocity of the shocked secondary is likely to contribute to the Doppler shift, the main variations in the centre of the lines must be explained by other means, e.g. the orbital motion as we suggest.

The Fe ii λ6455 line also shows variations in the Doppler shift towards the central source (Damineli et al. 2008b), but not in the polar-direction reflected light (Stahl et al. 2005; this line will be studied in Section 2.3).

2.3 Fitting the Doppler shift with orbital motion

We start by presenting the main results of KS07 and extending them to the big-masses model. The assumption is that lines with a difference between minimum and maximum of more than 200 km s\(^{-1}\) arise from the secondary star or the acceleration zone of its wind. We attribute the He i line formation to the acceleration zone of the secondary’s wind. Our model is based on a simple Doppler shift. The parameters are (i) inclination, (ii) masses of two stars, (iii) eccentricity, (iv) orbital period. All these are taken from other works in the literature; we only changed the inclination and eccentricity by several per cent. (v) The secondary’s wind region where the lines are formed. This is the only new parameter in our model. By fitting the line formation region to \(v = -340\) km s\(^{-1}\), we got the best fit (KS07). Notable are some of the broad He i lines. These lines are formed in the dense part of the wind, namely where the velocity is far from its terminal speed. Typically, the velocity there is in the range 400–500 km s\(^{-1}\), while the secondary’s wind terminal velocity is ~3000 km s\(^{-1}\). In the first panel of Fig. 2, we present two cases of the small-masses model; more examples for this model are in KS07. As the data are noisy, a fit should be used. We did not add the ~8 km s\(^{-1}\) system velocity of η Car as it is negligible relative to other uncertainties at this stage. Our calculations are compared with the data from Nielsen et al. (2007a). Nielsen et al. (2007a) do not present a physical model, but rather a pure mathematical fitting.
to the data points. Being so close to Nielsen et al. (2007a) model is considered favourable to our model and the suggested orientation.

The orbital motion alone cannot explain in full the Doppler variations, and the other effects related to the intrinsic properties of the winds must be included in a full study. The ionizing radiation of the secondary’s wind influences the primary’s wind acceleration zone (e.g. Nielsen et al. 2007a; Damineli et al. 2008b), and the radiation of the colliding winds might influence the acceleration zone of the secondary’s wind (Soker & Behar 2006). The influence of these types of radiation increases with decreasing orbital separation, and is not equal at similar phases before and after phase zero (periastron). These are the main effects related to the variations in line intensity. However, compared to the Doppler shift, these are second-order effects. This is supported by the observation of intensity variations toward the polar direction, but without Doppler variations (see Sections 2.2). For that, our fit, in particular near periastron, cannot be perfect as these effects are not included in the present study. Considering this, and the fact that we did not try to play with the parameters too much, we consider the results presented in Fig. 2 as a strong support to the orbital motion interpretation of the Doppler shifts and for the values of $\omega = 90^\circ$. We note that Nielsen et al. (2007a) and Damineli et al. (2008b) attributed the P Cygni lines to the primary’s wind and their Doppler shift variations to the influence of the secondary’s ionizing radiation.

Some lines show much lower periodic velocity variations and are likely to be formed in other regions of the binary system and its winds. The first candidate is the primary and its wind. We attribute the Fe II $\lambda 6455$ line to the primary’s wind. Using the same binary parameters as for the lines arising from the secondary and, in particular, the same orientation with $\omega = 90^\circ$, namely the primary is on the far side at periastron and towards us at apastron, we calculated the Doppler shift of this line. Our results compared with the observations of Damineli et al. (2008b) near periastron passage are presented in Fig. 3. As discussed above, it is unlikely to perfectly fit the shifts near periastron. However, our interpretation correctly gives the amplitude of the velocity variation and suggests that this line can arise in the primary’s wind.

The colliding wind region can also be a source of lines showing orbital Doppler shifts. Behar, Nordon & Soker (2007) have already considered the Doppler shift of X-ray emission lines from the shocked secondary’s wind region. This region is closer to the centre of mass than the secondary is, hence the amplitude of the Doppler shift is lower. Still closer to the centre of mass is the shocked primary’s wind along the line joining the two stars. This region is only a few $\times 0.01r_2$ closer to the centre of mass than the stagnation point, where $r_2$ is the distance of the secondary star from the stagnation point (Kashi & Soker 2007a). We therefore take this region to be at a distance $0.02r_2$ closer to the centre of mass (closer to the primary) than the stagnation point is. We will try to attribute the Doppler shift of the hydrogen Paschen emission lines to this region. Damineli et al. (1997), 2000) and Davidson (1997) already noted that the Doppler shifts of these lines might follow an eccentric orbit behaviour, but they attributed its source to the primary star.

In Fig. 4, we plot the observations of Damineli et al. (2000) as well as the expected Doppler shift of lines emitted from a region near the stagnation point in our binary model. Our calculation fits very nicely with the Doppler shift. Only just before and after the phase zero, our calculated Doppler shift overestimates the observed value. In any case, because of the ionization by the two stars (see the above discussion) and the expected collapse of the winds-interaction region, we would be surprised if the simple model fits the observations very close to phase zero. Near periastron passages, the colliding-winds region might collapse on to the secondary (Soker 2005; KS09). Namely, this region ceases to exist in its simple form, and other regions become the dominant source of the Paschen lines. Although the fit is not perfect, it is good enough considering we did not change any parameter to fit the Doppler shifts of the hydrogen lines but rather use the same parameters as we have done in explaining the He I and Fe II lines. The figure shows our results for the small-masses model $(M_1, M_2) = (120, 30) M_\odot$. Our results for the big-masses model, $(M_1, M_2) = (160, 60) M_\odot$, which are not presented, are very similar to those of the small-masses model.

Let us elaborate more on the nature of the wind interaction. The location of the stagnation point is determined by the equilibrium between the ram pressures of the two winds. The orbital motion influences the ram pressure, mainly that of the slower primary’s wind. Therefore, when the two stars approach each other, the stagnation point moves closer to the secondary star. Namely, its distance from the centre of mass increases. The opposite behaviour occurs as the distance between the two stars increases after periastron passages. This effect can be seen in fig. 2 of Akashi et al. (2006). The effect is significant near periastron. We neglect this asymmetrical behaviour before and after periastron passage for the following reasons. (1)
Figure 4. The Doppler shift of the Paschen lines. In our model, the source of the Paschen lines is the shocked primary’s wind which is 0.02 $r_2$, closer to the centre of mass (closer to the primary) than the stagnation point is, where $r_2$ is the distance of the secondary star from the stagnation point. The inclination angle is $i = 45\degree$ for both panels. Diamonds: Paschen lines data for 5 keV gas, because we know that this region $kT \approx 10$ keV. To calculate the column density for our preferred orientation of the Paschen lines, we do not know where exactly they are formed, and how this region changes its location as the shocked primary’s wind density increases with decreasing distance. (2) Near periastron many other effects are likely to occur. The radiations from the two stars influence the colliding winds; gravity, e.g. tidal interaction, can also influence the colliding wind process. (3) An accretion process might occur for ~10 weeks (KS09). (4) Due to the orbital motion of the stars, the apex of the colliding winds – the stagnation point – will not be exactly on the line joining the centres of the two stars (Soker 2005). The angular offset is too small to be considered in our analytical study, and we must neglect it. For all these effects that add to the uncertainties near periastron passages, and are not considered by us, there is no justification to add the asymmetry in the stagnation point location to the present first-order calculations. We neglect all these effects in the present study, and in addition do not really try to fit the parameters to the Paschen lines.

Even if the Paschen lines formed indeed in the shocked primary’s wind, we do not know where exactly they are formed, and how this region changes its location as the shocked primary’s wind density increases with decreasing distance. (2) Near periastron many other effects are likely to occur. The radiations from the two stars influence the colliding winds; gravity, e.g. tidal interaction, can also influence the colliding wind process. (3) An accretion process might occur for ~10 weeks (KS09). (4) Due to the orbital motion of the stars, the apex of the colliding winds – the stagnation point – will not be exactly on the line joining the centres of the two stars (Soker 2005). The angular offset is too small to be considered in our analytical study, and we must neglect it. For all these effects that add to the uncertainties near periastron passages, and are not considered by us, there is no justification to add the asymmetry in the stagnation point location to the present first-order calculations. We neglect all these effects in the present study, and in addition do not really try to fit the parameters to the Paschen lines.

3 THE COLUMN DENSITY

In this section, we discuss the hydrogen column density as deduced from X-ray observations. Hamaguchi et al. (2007) calculated the hydrogen column density towards the X-ray emitting gas in the centre of $\eta$ Car. They differentiated between the column density towards gas having a temperature of $kT > 5$ keV and a gas of $kT \sim 1$ keV. We will study only the emission from the $kT > 5$ keV gas, because we know that this region comes from the shocked secondary’s wind before it suffers any adiabatic cooling. Therefore, we can safely estimate its location to be close to the stagnation region, but somewhat closer to the secondary.

The source and location of the $kT \sim 1$ keV gas, on the other hand, are less secure. It might come from a fast polar outflow from the primary star (Smith et al. 2004 argue for a fast polar wind from the primary star). Alternatively, it might come from an oblique shock of the secondary’s wind away from the secondary.

Yet, another possibility is a post-shocked secondary’s wind that had cooled adiabatically as it streamed away from the stagnation-point region. The very important thing that we do learn from Hamaguchi et al. (2007) analysis of the X-ray emission by the $kT \sim 1$ keV gas is that the contribution of gas around the interaction region to the column density is $N_H \sim 5 \times 10^{22}$ cm$^{-2}$. This implies that any column density of $N_H \gtrsim 5 \times 10^{22}$ cm$^{-2}$ must come from material close to the binary system, at the most ~100 au.

Indeed, the column density towards the $kT > 5$ keV gas at phase 0.47 is $N_H = 17 \times 10^{22}$ cm$^{-2}$. This is hard to explain in a model where the secondary is towards as during this phase, i.e. $\omega \sim 270\degree$. This is because then our line of sight towards the shocked secondary’s wind will go through the undisturbed secondary’s wind, namely through the opening of the conical shock formed by the colliding wind. Instead, we suggest that the secondary star resides on the far side during apastron, and the column density includes the undisturbed primary’s wind as well as the shocked primary’s wind.

In principle, if the line of sight goes through the shocked primary’s wind, i.e. $\omega = 270\degree$, then absorption can be very high. However, if this was the case near apastron, then at a later phase the line of sight would have gone through the primary’s wind before the wind was shocked. In this case, the calculations of Akashi et al. (2006) in their equation (6) show that the column density becomes $N_H = 17 \times 10^{22}$ cm$^{-2}$ only at an orbital separation of $a = 10$ au at phase ~0.06 (or 0.94). Therefore, orientation with $\omega = 270\degree$ will not explain the value of $N_H = 17 \times 10^{22}$ cm$^{-2}$ at phase 0.92 as observed by Hamaguchi et al. (2007). A very specific orientation and wind properties would be required to account for a large $N_H$ both at phases 0.5 and 0.92 in the case the secondary is towards us at apastron ($\omega = 270\degree$). Till date, there is no specific model showing this. On the contrary, a new paper by Okazaki et al. (2008b) shows that near apastron the expected opening angle of the conical shell is larger than the angle of the line of sight from the equatorial plane ($90\degree - i$), and therefore we observe the X-ray emitting region through the low-density secondary wind, as we assume.

To calculate the column density for our preferred orientation of $\omega = 90\degree$; we use the geometry drawn schematically in Fig. 5 for the colliding winds (for more details see Kashi & Soker 2007a, 2008). The contact discontinuity shape is approximated by a hyperbola at

Figure 5. Schematic drawing of the collision region of the two stellar winds and the definition of several quantities used in the paper. The point marked ‘X-ray source’ in the post-shocked secondary wind is where we take the $kT > 5$ keV gas to reside. From there we calculate the value of $N_H$.© 2008 The Authors. Journal compilation © 2008 RAS, MNRAS 390, 1751–1761

Downloaded from https://academic.oup.com/mnras/article-abstract/390/4/1751/980918 on 29 July 2018
A. Kashi and N. Soker

a distance $D_2 \simeq 0.3 \, r$ from the secondary at the stagnation point of the colliding winds, where the two wind momenta balance each other along the symmetry axis and with an asymptotic angle of $\phi_a = 60^\circ$.

The density of the post-shocked primary’s wind introduces large uncertainties because it is strongly influenced by the magnetic field (Kashi & Soker 2007a). We will therefore limit the ram-pressure compression factor to the smaller magnetic pressure compression factor ($f_m$). The thickness of the post-shocked primary’s wind shell (from the contact discontinuity to the primary’s wind shockwave) is calculated accordingly. The hydrogen number density is

$$n_H = \frac{0.43 \, f_m \, M_1}{4 \pi r_s^2 \, v_1 \, \mu m_H} \quad (1)$$

where $v_{wind}$ is the primary’s wind velocity. The compression factor together with a few more assumptions allows us to calculate the velocity of the post-shock primary’s wind out from the shock region and the width of the conical shock. From the width we can calculate the values of $d_s$ and $l_m$ defined in Fig. 5.

The source of the hard X-ray is the post-shocked secondary’s wind (Corcoran 2005; Akashi et al. 2006), taken here at the point marked in Fig. 5 by ‘X-ray source’; of course, the X-ray emitting region is more extended and our treatment is crude. This point is located at a distance of $(1 - u)D_2$ from the stagnation point or a distance of $uD_2$ from the secondary. We assume $u = 0.7$. As is evident from the figure, the column density has two main components: the post-shocked primary’s wind component $N_{H,\, shock}$ (the conical shell) and the undisturbed, free-expanding, primary’s wind component ($N_{H,1}$). We calculate the contribution of each component to the total column density ($N_{H,\, tot}$) as a function of orbital angle $\theta$.

As before, we take the inclination angle to be $i = 45^\circ$, and assume that the secondary is away from us during an apastron passage ($\omega = 90^\circ$). This geometry explicitly determines the direction from which the system is observed (i.e. line of sight) at each orbital phase. For every orbital angle $\theta$, we calculate the relevant direction angle $\xi$ to the observer. Considering the orientation of the conical shell at that orbital angle, we calculate the thickness of the conical shell in that direction and integrate $n_H$ over the width to find the column density of the first component

$$N_{H,\, shock} = \int_{l_{\text{out}}}^{l_{\text{in}}} n_H \, dl, \quad (2)$$

where $l = l_{\text{out}} + l_{\text{in}}$. The second component contributing to the column density is calculated from the point on the line of sight where the shock terminates to infinity (contribution decreases fast with distance),

$$N_{H,1} = \int_{l}^{\infty} n_H \, dl. \quad (3)$$

To calculate the total column density, we added a constant value of $4 \times 10^{22}$ cm$^{-2}$ for the material residing in the outer regions, e.g. in the Homunculus and interstellar medium.

The three column density components and the total column density are plotted in Fig. 6. The $N_{H} [\geq 5 \, \text{keV}]$ component from Hamaguchi et al. (2007) is also plotted.

At phase $-0.015$, the relative orbital velocity of the two stars is large, $\sim 200$ km s$^{-1}$, and the winds of the conical shell (Nielsen et al. 2007a; Okazaki et al. 2008a), as well as the collapse of the colliding wind region (KS09), cannot be neglected until after the recovery at phase $\sim 0.03$ (60 d after periastron passage). For that, from phase $-0.02$ to 0.03, the total column density does not include a contribution from the conical shell, and the undisturbed primary’s wind is assumed to be extended to the secondary, instead of towards the stagnation point. This is the reason for the change in behaviour of the total column density near periastron. In any case, close to periastron our treatment is very crude, but still gives the observed values.

We do not take into account the absorption through the unshocked secondary’s wind close to periastron. There are two reasons for that: first is that there is no conical shell near periastron. Instead, the secondary’s wind covers a short distance in the radial direction away from the primary because the conical shell (winds-interaction region) is wrapped around. The second reason is that we follow the accretion model, where near periastron the secondary accretes mass and the secondary’s wind is very weak.

Our model reproduces to within a factor of 2 the results of Hamaguchi et al. (2007) and to a good degree its qualitative behaviour. This is done without any parameter fitting. For example, if the primary’s wind is such that its mass-loss rate is $2^{1/2}$ smaller and its wind velocity $2^{1/2}$ larger than the values we use, then the column density will be very close to the observed value. The primary’s wind momentum flux would stay the same, such that the X-ray proper-motion of the shocked secondary’s wind would not change. Our results clearly show that from our preferred line of sight ($\omega = 90^\circ$, $i = 45^\circ$), the column density hardly changes during most of the orbital cycle.
and can supply the required high column density in accordance with the observations. When the system approaches periastron passage, there is a fast increase of $N_{\text{HI, shock}}$ and $N_{\text{HI,1}}$, followed by a decrease after periastron passage.

We note that the observations in phase 0.47 mentioned by Hamaguchi et al. (2007), $N_{\text{HI}} = 17 \times 10^{22}$ cm$^{-2}$, occurred when the 2–10 keV emission was 10–15 per cent above its average value during that time (Corcoran 2005). This could result from a higher density of the primary wind. According to the model of Akashi et al. (2006), the dependence of the X-ray emission on the primary mass loss is $L_x \sim M_p^{1/2}$. Namely, it is quite possible that the primary mass-loss rate, and hence $N_{\text{HI}}$, was higher than the average value near apastron by a factor of $\sim 1.25$, and this is the reason Hamaguchi et al. (2007) did not find the column density to increase between phases 0.47 ($-0.53$) and 0.92 ($-0.08$).

Another effect not considered by us, and that introduces more variations both in the time variation and in the absorption by the conical shell along different directions, is the corrugated structure of the shocked primary wind. The model of the shock and the wind interaction region, as observed at $\beta = 45^\circ$, is the one built for the radio emission (Corcoran 2005). This could result from a higher density of the primary wind. According to the model of Akashi et al. (2006), the dependence of the X-ray emission on the primary mass loss is $L_x \sim M_p^{1/2}$. Namely, it is quite possible that the primary mass-loss rate, and hence $N_{\text{HI}}$, was higher than the average value near apastron by a factor of $\sim 1.25$, and this is the reason Hamaguchi et al. (2007) did not find the column density to increase between phases 0.47 ($-0.53$) and 0.92 ($-0.08$).

As mentioned earlier, behind the secondary shock, it is not possible to account for $N_{\text{HI}} = 17 \times 10^{22}$ cm$^{-2}$ near apastron. It cannot come from the nebula, as the nebula can supply $5 \times 10^{22}$ cm$^{-2}$ at the most. We therefore conclude that the $N_{\text{HI}}$ observations support a value of $\omega \sim 90^\circ$.

4 HIGH-EXCITATION NARROW LINES

The narrow lines originate mainly in the Weigelt Blobs (WBs) located around the main source (e.g. Smith et al. 2004; Gull, Vieira Kober & Nielsen 2006; Nielsen, Ivarsson & Gull 2007b; Damineli et al. 2008b). A representative high-excitation narrow line is the $[\text{Ar} \text{III}] \lambda 7135$ line (Damineli et al. 2008b). As noted by Damineli et al. (2008b), this line has many similarities to the radio light curve at 3 cm (Duncan & White 2003), mainly by showing a continuous variability along the cycle. Although the main source of the line is the WBs, contributions from other regions are expected (Groh, Damineli & Jablonski 2007). This can be deduced from the presence of gas between the WBs and to other directions at about the same distance (e.g. fig. 1 in Gull et al. 2006).

In an attempt to model this line, we build a simple model based on the periastron longitude $\omega = 90^\circ$ proposed by us. We examine high-ionization lines, such that the ionization energy must be supplied by the secondary and not by the primary. We examine the rate of secondary ionizing photons that reach Weigelt Blob-D (WB-D), as a function of the orbital phase. The location and angular dimensions of WB-D were measured from fig. 1 of Gull et al. (2006), taking the distance to $\eta$ Car to be 2.3 kpc. Blob D is located 0.25 arcsec from the central source (Gull et al. 2006; Nielsen et al. 2007b). We assume that its location lies exactly on the continuation of the semimajor axis of the elliptic orbit (See Fig. 1). Taking the inclination angle to be $45^\circ$, we find Blob D to be $r_\text{B} \sim 850$ au (5 light-days) from the primary. With the secondary at the apex, Blob D forms a base of a cone with a full opening angle of $\sim 30^\circ$ (Gull et al. 2006). We therefore examine the ionizing radiation emitted within a cone with a half-opening angle of $15^\circ$ and towards Blob D. Weigelt Blob B (WB–B) seems to be in the same direction as Blob D, and this model applies to emission from Blob B as well. However, there is a contribution to the emission-line intensity from other regions, such as Blob C. Therefore, the model will only show that the periastron longitude that we propose can reproduce the general observed behaviour. Detailed modelling and parameter fitting are postponed until after the next event.

The model is based on the one we built for the radio emission (Kashi & Soker 2007a), where the gas responsible for the radio free-free emission was assumed to be ionized mainly by the secondary. Whereas in the model for the radio emission, the secondary ionizing radiation to all directions was considered, presently we examine only the radiation towards the WBs and within a cone with a half opening angle of $15^\circ$. For more details on the model, the reader should consult Kashi & Soker (2007a).

In its way from the secondary to the blob, the secondary radiation suffers from absorption by the post-shocked primary’s wind gas in the conical shell and by the undisturbed primary’s wind. The total absorbing rate per steradian is marked by $N_{\text{abs}}$. We take the rate of secondary’s ionizing photons entering the $15^\circ$ cone towards the blob and subtract from it the absorption rate. Our results very close to periastron cannot be considered accurate for the same arguments given in Section 3. The primary’s wind mass-loss rate is $M_1 = \frac{3 \times 10^{-4}}{\eta_1}$ M$_\odot$ yr$^{-1}$ and its terminal speed $v_1 = 500$ km s$^{-1}$. From these we can calculate the absorption by the undisturbed primary’s wind. The absorption by the post-shock gas in the conical shell depends strongly on its density there because the recombination rate of the gas depends on the density squared. The density of the post-shock primary gas depends on the compression of the post-shock gas by the ram pressure of the primary’s wind. The compression is determined by the balance between the ram pressure of the wind and the internal pressure of the post-shock gas, including thermal and magnetic pressure. Therefore, the compression strongly depends on the magnetic field strength and geometry in the pre-shock primary’s wind. The magnetic field strength is characterized by the pre-shock magnetic to ram pressure ratio $\eta_B$ (see Kashi & Soker 2007a for more technical details). This compression factor of the shocked wind introduces large uncertainties.

It is difficult to know the exact recombination rate close to the stagnation point ($\alpha \sim 140^\circ$–$180^\circ$; see Kashi & Soker 2007a). In that region, self-ionization may change the recombination rate. In order to partially compensate for this effect, we have made a cut-off to the conical shell’s increased recombination rate at $\alpha = 140^\circ$. Namely, we assumed that the recombination rate per steradian in the range $\alpha = 140^\circ$–$180^\circ$ is equal to the one calculated at $\alpha = 140^\circ$. This procedure causes a small change in behaviour of the calculated rate of ionizing photons reaching the blobs, like the ones seen in the upper line of Fig. 7, at $\phi = \pm 0.0264$ ($\theta = \pm 125^\circ$). The rate of ionizing photons at Blob D is at minimum close to periastron because the dominant absorber is the primary’s undisturbed wind. The column density between the secondary and the blob becomes larger as the secondary dives into the primary’s wind acceleration zone close to the periastron passage (recall that the conical shell does not exist near periastron).

In the spirit of this paper, we minimize the parameter fitting and consider only the most influential processes. For example, we do not consider the very likely possibility that the primary wind is clumpy and neglect the instabilities in the wind S interaction region (Pittard et al. 1998; Pittard & Corcoran 2002; Okazaki et al. 2008a).

In addition, we do not consider the ionization of the conical shell by the radiation of the shocked winds, which might be significant near the stagnation point. We stay with the same parameters that we used in our model for the radio emission. These are the pre-shock magnetic to ram pressures ratio, $\eta_B = 0.001$–0.1, and the rate of emission of ionizing photons per steradian by the secondary, $N_{\text{abs}} = 2 - 3 \times 10^{48}$ s$^{-1}$ sr$^{-1}$. In a recent paper, Teodoro et al. (2008) obtained $N_{\text{abs}} = 2 \times 10^{48}$ s$^{-1}$ sr$^{-1}$. The rate of ionizing photons
A. Kashi and N. Soker

\[ \eta_0 = 0.214 \text{ sr} \] is the solid angle covered by the blob. In Fig. 7, \[ \eta = 0.4 \] and on the pre-shock primary’s wind and the rate of emission of ionizing photons by the secondary, \( N_{t1} \) per steradian.

reaching Blob D is given by

\[ \dot{N}_{ID} = [\dot{N}_{t1} - \dot{N}_{abs}(%0) \Omega], \]  

where \( \Omega = 0.214 \text{ sr} \) is the solid angle covered by the blob. In Fig. 7, we show the dependence of \( \dot{N}_{ID} \) on \( N_{t1} \) and on \( \eta_B \).

As expected, the ionizing flux reaching the blobs increases with the increase in the secondary photon ionizing rate. Because the value of \( \dot{N}_{ID} \) is the difference of two numbers, it might become sensitive to the value of \( N_{t1} \) when the absorption rate \( \dot{N}_{abs} \) is non-negligible. This is clearly the case for our model. As the strength of the pre-shock magnetic field, parametrized by \( \eta_B \), increases, compression decreases. The lower density implies a lower photon absorption rate \( \dot{N}_{abs} \) and the value of \( \dot{N}_{ID} \) increases.

To reach a better fitting, we now consider two effects that were neglected in the simple calculation presented in Fig. 7. They are the likely variation of the magnetic field with time, and the presence of gas around Blobs D and B. In Kashi & Soker (2007a), we showed the magnetic field of the primary’s wind to be an important factor in the behaviour of \( \eta \) Car. The problem with the magnetic fields is its stochastic nature, as is well known from the Sun. We have no knowledge of its behaviour. But knowing that it is very important, we try to deduce its behaviour in our model. So, basically, in our model the behaviour of ionizing photons can provide us some information about the magnetic field in the primary wind. From the behaviour of the Sun we know that the magnetic field can have semiperiodic, as well as stochastic, variations. In addition, the magnetic field has spatial variations, neglected by us. We consider the simplest variation possible where the magnetic strength of the primary’s wind decreases from a maximum value during its active phase, taken to be at phase \( \phi = -0.5 \) in the time period we consider, to a quiescent value. The functional form we use is

\[ \eta_B = \eta_{B-act} \exp(-t/\tau) + \eta_{B-q}, \]  

where \( \eta_{B-q} \) is the quiescent value of \( \eta_B \). We have no prior knowledge of the values of these parameters, so we looked for a simple fitting. We found that \( \tau = 1/6 \times \text{cycle} \approx 340 \text{ d} \), \( \eta_{B-q} = 0.001 \) and \( \eta_{B-act} = 0.004 \) give good fitting. The maximum value of \( \eta_B \) occurs at \( \phi = -0.5 \) (\( \tau = 1012 \text{ d} \)) and it is \( \eta_B(\text{max}) = 0.024 \). In Fig. 8, we plot the total number of ionizing photons available to form narrow lines in Blobs B and D when the value of \( \eta_B \) from equation (5) and \( \dot{N}_{t1} = 3.5 \times 10^{48} \text{ s}^{-1} \text{sr}^{-1} \). The fit is also drawn with 80 d recombination time delay (thick dash–dotted line).

The matter residing in the vicinity of the blobs, as well as in other regions, has two effects. First, these regions are continued to be ionized even after photons do not reach the WBs any more. Secondly, their lower density implies that their recombination time is longer, and their decline can lag behind on a typical time-scale of weeks, e.g. six weeks for an electron density of \( n_e = 10^{-6} \text{ cm}^{-3} \), which is nearly \( \sim 10 \times \) times lower than the density in the blobs (Damineli et al. 2008b).

These effects, i.e. recombination time of lower density regions, are important in the decline, when the contribution of the WBs to the emission lines rapidly declines. We find that a shift of the \( \dot{N}_{ID} \) plot by 80 d to a later time gives better fit to the observation just before the event. This is presented by the dash–dotted line in Fig. 8. Our explanation to the better fit with this 80 d displacement is that the contribution to this line emission just before the event, when the emission is already weak, comes mainly from the regions with a recombination time of \( \sim 80 \text{ d} \). It is evident that the fitting during most of the cycle, when emission is high, does not change much but the fitting during the decline phase is much better.

We can summarize this part by emphasizing the success of the model based on our proposed periastron longitude. By a simple model and fitting, the behaviour of the high-excitation line emission can be reproduced if the absorption of the secondary ionizing radiation by the shocked primary’s wind in the conical shell is considered. This can only be the case for a periastron longitude of \( \omega \sim 90^\circ \), namely the secondary is away from us during apastron and during most of the orbital motion (see Fig. 1). The main ingredients of the model are the same as that used by us in explaining the radio emission from \( \eta \) Car (Kashi & Soker 2007a), and no new model had to be invented here. We only had to use the parts relevant to the WBs. The consideration of the variation of the magnetic field, for example, was also used in the radio model.
5 DISCUSSION AND SUMMARY

5.1 Main results

Our goal was to learn about the orientation of the semimajor axis of the η Carinae binary system from several different observations. We found that an orientation where the hotter secondary star is closer to us for a short time at periastron, i.e. a periastron longitude of $\omega = 90^\circ$, fits these observations. During most of the time, the primary is closer to us as depicted in Fig. 1.

In Section 2, we attributed the high-excitation He II lines to the secondary wind, and by that could reasonably fit the variation of the Doppler shift with the orbital motion (for more details see KS07). The Doppler shift of the low-excitation Fe II $\lambda 6455$ line could be fitted by attributing its origin to the primary stellar wind. The orbital motion explanation for the Doppler shift accounts also for the absence of the Doppler shift variation towards the polar directions. The Doppler shift of the Paschen lines could be accounted for if they are assumed to be formed in the shocked primary’s wind near the stagnation point (see Fig. 1). We also note that the variation in the Doppler shift of the X-ray lines might also be caused by the orbital motion (Behar et al. 2007).

Trying to explain the lines’ behaviour, we used two sets of primary and secondary masses which we consider possible (see KS09) and obtain nice fits without overplaying with other parameters. We emphasize that claims for exaggerated values for those masses are no problem for the qualitative model we present. Using lower masses, we were still able to fit all the lines in the paper by slightly adjusting other parameters (eccentricity, inclination, magnetic compression, cone-opening angle, mass-loss rates, etc.) well within their acceptable range.

In Section 3, we examined the hydrogen column density towards the hard X-ray emitting gas, $N_H(>5\,\text{keV})$. The column density at several times along the orbit is given by Hamaguchi et al. (2007). The column density is sensitive to the mass-loss rate and velocity of the primary’s wind and to the nature of the wind interaction process, e.g. where exactly the gas emitting the hard X-ray resides. We cannot reproduce the exact variation of $N_H$ with orbital phase, but could reproduce the approximate value at each phase (Fig. 6).

In an opposite binary orientation, where the secondary is towards us during most of the time (near apastron), the expected value of $N_H$ near apastron is much smaller than the observed value. It is also expected to rise towards periastron by a much larger factor than what is observed. Overall, although our fit is not perfect and requires further work, it has less severe problems than what a model based on an opposite orbital orientation would have.

We did not deal with the value of $N_H(1\,\text{keV})$ towards the gas emitting the soft X-ray ($\sim 1\,\text{keV}$), as we are not sure where this gas is located. However, during most of the time $N_H(1\,\text{keV}) \simeq 0.3 N_H(>5\,\text{keV})$. This suggests that if it was not for the dense primary’s wind around the secondary, the hot X-ray emitting gas would have had lower $N_H$ than the observed value. It is not easy to account for the high $N_H$ value in a model where during most of the time we observed the shocked secondary’s wind through the tenuous secondary wind ($\omega \sim 270^\circ$).

In Section 4, we calculated the hard ionizing radiation that is emitted by the secondary and reaches the WBs. We found that when the absorption by the undisturbed and shocked primary’s wind is considered, the qualitative behaviour of the high-excitation [Ar III] $\lambda 7135$ line that is assumed to be emitted mainly by the WBs (Damineli et al. 2008b) can be reproduced.

The absorption by the shocked primary’s wind – in the conical shell – depends on the compression of the post-shock gas (Fig. 7). In our model (Kashi & Soker 2007a), the compression is constrained by the post-shock magnetic pressure. The post-shock magnetic pressure is determined by the ratio of the pre-shock magnetic pressure to the wind’s ram pressure $p_w$. To fit the observed behaviour (Fig. 8), we had to assume that the magnetic field in the primary’s wind evolves according to equation (5). In general, it is expected that the magnetic field in stellar wind will change over time-scales of years to tens of years, as it is very well known for our Sun.

To summarize, using the model that has been proposed in our study of the radio emission from η Car (Kashi & Soker 2007a), we reproduced the basic properties of the ionizing radiation that are required to form the high-excitation narrow lines.

5.2 Further considerations

We list four observational results that our proposed $\omega = 90^\circ$ periastron longitude can account for.

5.2.1 The evolution of the radio emission

The evolution of the radio emission also supports the orientation proposed here. The radio image contains some bright knots, with fainter emission in areas between and around these knots (Duncan & White 2003). In particular, we note a bright radio knot to the same direction relative to the centre as the WBs are. If the orientation was opposite to what we claim here, then the ionizing radiation of the secondary towards this bright region would be constant until $\lesssim 3$ months before periastron. The reason is that the radiation from the secondary to the knot would propagate through the tenuous secondary’s wind. Namely, in the model where the secondary is towards us near apastron, the bright radio region would stay more or less at the same maximum brightness until phase $\sim -0.05$, while region to the sides would decline slowly, as the secondary dives into the denser part of the primary’s wind. However, a careful examination of the radio movie (White et al. 2005) shows that the radio knot fades after apastron passage towards the 1998 and 2003 minima, as the rest of the nebula does. This is expected in the periastron longitude $\omega = 90^\circ$ proposed by us, when the knot is being irradiated through the primary’s wind, which continuously becomes denser as the system approaches periastron. This fading of the radio emission as a result of the secondary ‘diving’ into the dense primary’s wind has been discussed before by Duncan & White (2003) and Abraham & Falceta-Gonçalves (2007).

5.2.2 The purple haze

Smith et al. (2004) presented the ultraviolet (UV) images of the Homunculus at six epochs around the 2003 minimum. They found that just before periastron, the bright UV region extended to the south-east along the symmetry axis of the Homunculus. Smith et al. (2004) assumed that the UV emitting-reflecting regions are in the equatorial plane, and from that they deduced that the semimajor axis is perpendicular to our line of sight, namely a periastron longitude of $\omega = 0^\circ$.

Abraham & Falceta-Gonçalves (2007) also attributed the excess UV radiation to the secondary star. Close to periastron, this radiation, according to Abraham & Falceta-Gonçalves (2007), is confined by dust absorption to the interior region of the colliding-winds cone; the cone points eastwards before and westwards after the event.

Orientation of the $\eta$ Carinae binary system
We suggest a different interpretation of that behaviour. We note that the south-east lobe is towards us and attribute the UV-bright region in the south-east direction to the polar direction, rather than to the equatorial region. For example, a transient polar outflow (Behar et al. 2007) that opens a cone for the ionization and UV illumination might be the cause of this illumination rather than the orbital orientation at this phase.

We also note the following interesting result: the region of the WBs became UV bright for a short time about a month before periastron passage (Smith et al. 2004). This is at the same time as the peak in the narrow He I line intensity; the He I peak is seen a month before the 2003.5 minimum (event I), but not in the previous two minima (Damineli et al. 2008b). According to the orientation proposed in our model, the secondary is towards this side (closer to us) near that time, and it is possible that for a short time the ionizing radiation from the secondary towards the blobs actually increased; e.g. due to clumpy primary’s wind or instabilities, an opening in the dense conical shell was formed.

5.2.3 The He I λ10830 line

The He I λ10830 high-excitation line has a P Cygni profile with absorption changing from −640 to −450 km s⁻¹ (Damineli et al. 2008b). Just before periastron, a wide wing in absorption appears, reaching − 1000 km s⁻¹ a month before periastron and −1400 km s⁻¹ at periastron. This shows that there is a flow of a high-velocity gas towards us at periastron. The primary star is not expected to blow such a fast wind towards us, and if it does, why it cannot be seen in other phases as well. We attribute this behaviour to the shocked secondary’s wind or polar outflow. In our proposed orientation, the secondary is closer to us at periastron. The shocked secondary’s wind will flow towards us near periastron passage. This is why the wide wing of the He I λ10830 line is seen only close to periastron passage. In any case, as suggested by the recent study of Teodoro et al. (2008), the interpretation of the He I λ10830 requires careful attention and will be the subject of a forthcoming paper.

5.2.4 The location of the WBs

Chesneau et al. (2005) observed the subarcsecond butterfly-shaped dusty environment surrounding η Car with the Very Large Telescope Interferometer (VLTI), using the mid-IR instrument mid-infrared instrument for VLTI (MIDI) and the adaptive Optics system National Aeronautical Charting Office. As shown in fig. 7, the lower density ‘SE filament’ (named so by Smith et al. 2003a) is apparently aligned in the same direction as the NW WBs complex.

The existence and the large mass of the WBs might be related to their location close to the periastron of the system orbit, where the secondary’s wind is strongly disturbed by the primary’s dense wind. The counterpart (the SE direction) of this high-density region is a very low density region, but has some filaments inside it. The NE direction even has a lower density. The higher density in the NW might be the consequence of the eccentric orbital motion of the secondary. We predict that 3D numerical simulations using our proposed orientation will reproduce this density asymmetry.

5.3 Unclosed issues

The proposed orbital orientation and our study might have some weak points.

1) Coincidence. The coincidence that the secondary’s wind velocity where the lines are formed, \( v_{\text{line}} \approx -430 \text{ km s}^{-1} \), is practically the same as the primary wind terminal velocity. Although this is a somewhat weak point in our model, our answer to that coincident (KS07) is that the same can be said about the model where the He I lines originate in the primary’s wind. How come the changes in the velocity of the regions where the lines are formed in the primary’s wind exactly mimic the secondary velocity around the centre of mass? Another coincidence is that the best orientation is exactly for the secondary to be towards us at periastron and not even several degrees from that direction. However, we found (KS07) that several degrees deviation from \( \omega = 90^\circ \) is possible. Some other weak points regarding the orbital motion interpretation for the Doppler shifts are listed by Davidson (1997). However, like Damineli, Conti & Lopes (1997), he attributed the He I lines to the primary, while we attribute them to the secondary.

2) Magnetic field. The absorption of ionizing radiation depends strongly on the magnetic field in the primary’s wind. Our need for magnetic field adds a parameter to the model. As it stands now, it is a weak point. However, if magnetic fields are detected in the primary wind, this becomes a strong point. In particular, the magnetic field becomes strong in the shocked primary’s wind. We encourage a search for magnetic fields in the shocked primary region, e.g. by looking for its influence on some lines. The strength of the magnetic field near the stagnation point is expected to be \( B_{\text{shock}} \approx 100 (r/1 \text{ au})^{-1} \text{ G} \) if we assume that the post-shock magnetic pressure is about equal to the wind ram pressure. This field can be detected, e.g., in the Paschen lines very close to the periastron passage.

We note the following. Our need for magnetic field comes from the following consideration. If the post-shock primary gas has no magnetic field, it is compressed to a very high density in the conical shell, such that it absorbs too much of the ionizing radiation. Instabilities in the conical shell can reduce this absorption even without magnetic fields because most of the gas might be concentrated in dense clumps and the ionizing radiation escapes between these clumps. The study of this process requires 3D numerical simulations.

3) Helium lines from the secondary’s wind. The secondary luminosity is only \( \sim 20 \) per cent of the total luminosity. This might cause some problems in attributing the He I lines to its wind. However, we note that most of the radiation from η Car is in the IR anyhow. A detailed study of the line formation in the secondary is required, similar to the one Hillier et al. (2001, 2006) conducted for the primary star. For the time being, we note that many stars similar to the secondary are known to have P Cygni He I lines with velocity much smaller than their terminal velocity (see discussion in KS07).

ACKNOWLEDGMENTS

We thank Otmar Stahl, Arnout van Genderen and Diego Falceta-Gonçalves and an anonymous referee for enlightening comments. This research was supported by the Asher Space Research Institute in the Technion, P. and E. Nathan research fund and the Israel Science Foundation (grant No. 89/08).

REFERENCES

Abraham Z., Falceta-Gonçalves D., 2007, MNRAS, 378, 309
Abraham Z., Falceta-Gonçalves D., Dominici T. P., NYman L.-A. D. P., McAluliffe F., Capronni A., Jatenco-Pereira V., 2005, A&A, 437, 977
Akashi M., Soker N., Behar E., 2006, ApJ, 644, 451
Behar E., Nordon R., Soker N., 2007, ApJ, 666, L97
Corcoran M. F., 2005, AJ, 129, 2018
Corcoran M. F., Ishibashi K., Swank J. H., Petre R., 2001, ApJ, 547, 1034
Corcoran M. F., Pittard J. M., Stevens I. R., Henley D. B., Pollock A. M. T., 2004, Proceedings of X-Ray and Radio Connections. Santa Fe, NM, 2004 February 3–6 (astro-ph/0406294)

Chesneau O., Min M., Herbst T., Waters L. B. F. M., Hillier D. J., Leinert C., de Koter A., Pascucci I., 2005, A&A, 435, 1043

Damineli A., 1996, ApJ, 460, L49
Damineli A., Conti P. S., Lopes D. F., 1997, New Astron., 2, 107
Damineli A., Stahl O., Kaufer A., Wolf B., Quast G., Lopes D. F., 1998, A&AS, 133, 299
Damineli A., Kaufer A., Wolf B., Stahl O., Lopes D. F., de Araujo F. X., 2000, ApJ, 528, L101
Damineli A. et al., 2008a, MNRAS, 384, 1649
Damineli A., Hillier D. J., Corcoran M. F., Stahl O., Groh J. H., Arias J., Teodoro M., Morrell N., 2008b, MNRAS, 386, 2330
Davidson K., 1997, New Astron., 2, 387
Davidson K. et al., 2005, AJ, 129, 900
Dorland B. N., 2007, PhD Thesis, University of Maryland
Duncan R. A., White S. M., 2003, MNRAS, 338, 425

Falceta-Gonçalves D., Jatenco-Pereira V., Abraham Z., 2005, MNRAS, 357, 895
Falceta-Gonçalves D., Abraham Z., Jatenco-Pereira V., 2007, IAUS, 240, 198

Groh J. H., Damineli A., Jablonski F., 2007, A&A, 465, 993
Gull T. R., Vieira Kober G., Nielsen K. E., 2006, ApJ, 163, 173

Hamaguchi K. et al., 2007, ApJ, 663, 522
Henley D. B., Corcoran M. F., Pittard J. M., Stevens I. R., Hamaguchi K., Gull T. R., 2008, ApJ, 680, 705
Hillier D. J., Davidson K., Ishibashi K., Gull T., 2001, ApJ, 553, 837
Hillier D. J., Lanz T., Heap S. R., Hubeny I., Smith L. J., Evans C. J., Lennon D. J., Bourret J. C., 2003, ApJ, 588, 1039
Hillier D. I. et al., 2006, ApJ, 642, 1098

Humphreys R. M., Davidson K., Koppelman M., 2008, AJ, 135, 1249
Ishibashi K., Corcoran M. F., Davidson K., Swank J. H., Petre R., Drake S. A., Damineli A., White S., 1999, ApJ, 524, 983
Kashi A., Soker N., 2007a, MNRAS, 378, 1609
Kashi A., Soker N., 2007b, New Astron., 12, 590 (KS07)
Kashi A., Soker N., 2008, New Astron., 13, 569

Kashi A., Soker N., 2009, New Astron., 14, 11
Martin J. C., Davidson K., Hamann F., Stahl O., Weis K., 2006a, PASP, 118, 697
Martin J. C., Davidson K., Humphreys R. M., Hillier D. J., Ishibashi K., 2006b, ApJ, 640, 474
Nielsen K. E., Corcoran M. F., Gull T. R., Hillier D. J., Hamaguchi K., Ivarsson S., Lindler D. J., 2007a, ApJ, 660, 669
Nielsen K. E., Ivarsson S., Gull T. R., 2007b, ApJ, 168, 289

Okazaki A. T., Owoki S. P., Russell C. M. P., Corcoran M. F., 2008a, in Bresolin F., Crowther P. A., Puls J., eds, IAU Symp. 250, Massive Stars as Cosmic Engines. Cambridge Univ. Press, Cambridge, p. 1330
Okazaki A. T., Owoki S. P., Russell C. M., Corcoran M. F., 2008b, MNRAS, 390, 1730

Pereira C. B., Marcelino W. L. F., Machado M., de Araujo F. X., 2008, A&A, 477, 877
Pittard J. M., Corcoran M. F., 2002, A&A, 383, 636
Pittard J. M., Stevens I. R., Corcoran M. F., Ishibashi K., 1998, MNRAS, 299, L5
Smith N., 2002, MNRAS, 337, 1252
Smith N., Morse J. A., 2003, AAS, 202, 3218

Smith N., Morse J. A., Collins N. R., Gull T. R., 2004, ApJ, 610, L105
Soker N., 2005, ApJ, 635, 540
Soker N., Behar E., 2006, ApJ, 652, 1563
Stahl O., Weis K., Bomans D. J., Davidson K., Gull T. R., Humphreys R. M., 2005, A&A, 435, 303
Teodoro M., Damineli A., Sharp R. G., Groh J. H., Barbosa C. L., 2008, MNRAS, 387, 564

van Genderen A. M., Sterken C., Allen W. H., Walker W. S. G., 2006, JAD, 12, 3

Verner E., Bruhweiler F., Gull T., 2005, ApJ, 624, 973
White S. M., Duncan R. A., Chapman J. M., Koribalski B., 2005, in Humphreys R., Stanek K., eds, ASP Conf. Ser. Vol. 332, The Fate of the Most Massive Stars. Astron. Soc. Pac., San Francisco, p. 129
Whitelock P. A., Feast M. W., Marang F., Breedt E., 2004, MNRAS, 352, 447
Zanella R., Wolf B., Stahl O., 1984, A&A, 137, 79

This paper has been typeset from a \TeX/\LaTeX file prepared by the author.