

Université de Montréal

**Étude observationnelle de la collision des vents dans les systèmes  
Wolf-Rayet+O**

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## RÉSUMÉ

Nous présentons les résultats de trois campagnes d'observation d'un mois chacune dans le cadre de l'étude de la collision des vents dans les systèmes binaires Wolf-Rayet + OB. Ce travail se concentre sur l'étude des objets de l'hémisphère sud n'ayant jamais encore fait l'objet d'études poussées dans ce contexte. À cela, nous avons ajouté l'objet archétype pour ce type de systèmes : WR 140 (WC7pd + O5.5fc) qui a effectué son dernier passage périastre en janvier 2009. Les deux premières campagnes (spectroscopiques), ont permis une mise à jour des éléments orbitaux ainsi qu'une estimation de la géométrie de la zone de collision des vents et d'autres paramètres fondamentaux des étoiles pour 6 systèmes binaires : WR 12 (WN8h), 21 (WN5o+O7V), 30 (WC6+O7.5V), 31 (WN4o+O8), 47 (WN6o+O5) et 140. Une période non-orbitale courte (probablement reliée à la rotation) a également été mesurée pour un des objets : WR 69 (WC9d+OB), avec une période orbitale bien plus grande. La troisième campagne (photométrique) a révélé une variabilité étonnamment faible dans un échantillon de 20 étoiles WC8/9. Cela supporte l'idée que les pulsations ne sont pas courantes dans ce type d'étoiles et qu'il est peu probable que celles-ci soient le mécanisme dominant de formation de poussière, suggérant, par défaut, le rôle prédominant de la collision des vents.

**Mots clés :** (étoiles :) binaires : général, étoiles : paramètres fondamentaux (classification, couleurs, luminosités, masses, rayons, températures, etc.), étoiles : Wolf-Rayet, étoiles : vents, éjections de masse, étoiles : Wolf-Rayet – matière circumstellaire, étoiles : carbone



## ABSTRACT

We present the results from three month-long observational campaigns related to WR+O colliding-wind binaries. This work focuses on southern-hemisphere systems that have not yet been previously studied in this context. To this we add the northern archetype colliding-wind system WR 140 (WC7pd + O5.5fc), which recently suffered another active periastron passage in 2009 January. The first two (spectroscopic) campaigns allowed us to update the orbital elements and to constrain the geometry of the colliding wind region and other fundamental parameters, for 6 binary systems: WR 12 (WN8h), WR 21 (WN5o+O7V), WR 30 (WC6+O7.5V), WR 31 (WN4o+O8), WR 47 (WN6o+O5) and WR 140. A short non-orbital (probably rotational) period was also detected in the WR component of one of our objects, WR 69 (WC9d+OB), with a much longer orbital period. The third (photometric) campaign revealed a surprisingly low level of variability in our sample of 20 WC8/9 stars. This supports the idea that pulsations are not very frequent in this type of star and are unlikely to be the main mechanism for dust formation, suggesting, by default, a predominant role of wind-wind collision.

**Keywords :** (stars:) binaries: general, stars: fundamental parameters (classification, colors, luminosities, masses, radii, temperatures, etc.), stars: Wolf-Rayet , stars: winds, outflows, stars: Wolf-Rayet – circumstellar matter, stars: carbon



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# CHAPITRE 1

## INTRODUCTION

Les étoiles Wolf-Rayet sont le stade évolué des étoiles les plus massives de notre univers (Lamers et al., 1991). Ces étoiles, dont les masses des progéniteurs sur la séquence principale (les étoiles OB) sont plus grandes que  $25 M_{\odot}$  finiront toutes leur vie par une explosion de type supernova<sup>1</sup> (Crowther, 2007), laissant derrière elles une étoile à neutrons ou un trou noir. Ces étoiles existent sous trois types spectraux distincts pour la population I (Crowther, 2007) : celles dont le spectre est riche en raies d'hélium et d'azote provenant du cycle CNO (brûlage de H) : type WN. Et celles dont le spectre est riche en raies d'hélium, d'azote et d'oxygène, provenant du cycle triple-alpha (brûlage de He) : types WC et WO. Certaines WN présentent en plus des raies d'hydrogène et sont alors notées WNh. La particularité observationnelle la plus frappante de ces étoiles est la présence dans leur spectre de raies d'émission extrêmement fortes et élargies d'un millier de  $\text{km s}^{-1}$ . Ces raies sont dues à la présence d'un vent très dense s'échappant de l'étoile à des vitesses de l'ordre de  $1000 \text{ km s}^{-1}$ . Ce vent est tellement dense qu'il bloque complètement la lumière provenant de la surface (hydrostatique) de l'étoile. Il est également responsable d'une perte de masse totale tout au cours de la phase post- séquence principale (environ 500000 ans) de plus de la moitié de sa masse d'origine (Maeder and Meynet, 2010). Il est donc impensable d'essayer de comprendre l'évolution des étoiles massives sans considérer le paramètre clé qu'est leur taux de perte de masse  $\dot{M}$ . Or, même s'il existe plusieurs méthodes pour mesurer ce paramètre (Lamers and Cassinelli, 1999), ces mesures ne sont certainement pas triviales et sont souvent indirectes, elles sont notamment rendues difficiles par le caractère très inhomogène des vents (clumping). La connaissance

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<sup>1</sup>Les étoiles de masse situées entre 8 et  $25 M_{\odot}$  passeront plutôt par le stade de supergéante rouge (RSG) avant d'exploser en supernova de type II (i.e. riche en H), sans passer par le stade WR. La limite de masse pour former une WR dépend en réalité de la métallicité du milieu mais est estimée à  $25 M_{\odot}$  dans le voisinage solaire (Crowther, 2007). Pour les masses entre 25 et  $30 M_{\odot}$ , l'étoile passera par une phase RSG suivi d'une phase WN.

simultanée de la masse et du taux de perte de masse d’une Wolf-Rayet, essentielle à l’élaboration de modèles d’évolution, n’est, pour cette raison, acquise que pour un nombre limité d’étoiles. L’évolution du taux de perte de masse au cours de la vie d’une étoile est encore moins contrainte (voir section 1.1).

Lorsque deux étoiles massives interagissent gravitationnellement dans un système binaire, cela donne lieu à des objets astronomiques totalement inédits : binaires à collision de vents, binaires produisant de la poussière (“dustars”, Williams et al., 1987a), binaires à transfert de masse, binaires X massives, blue stragglers, objets de Thorne-Zytkow, etc. Nous nous intéressons ici aux binaires à collision de vents, et plus particulièrement aux binaires constituées d’une étoile Wolf-Rayet et d’une étoile OB, possédant un vent plus faible que celui de la WR. L’intérêt de ces binaires, dans le cadre de la compréhension des étoiles massives (en dehors de l’aspect purement hydrodynamique), est qu’elles donnent accès à une mesure simultanée des masses des deux composantes (à l’aide des lois de Kepler) et du taux de perte de masse des étoiles ne dépendant pas du clumping des vents, comme nous le verrons au cours de cette introduction. Ces objets sont également très lumineux, et ce dans une gamme de longueurs d’onde allant du domaine radio aux rayons X. Ils mettent en jeu des phénomènes physiques très variés, rendant leur étude passionnante.

Nous décrirons d’abord dans une première section quelques aspects théoriques et observationnels des vents d’étoiles massives, puis, dans une deuxième section, nous décrirons la physique impliquée dans les systèmes binaires à collision de vents. Le travail réalisé dans cette thèse concernant essentiellement les longueurs d’ondes visibles, nous limiterons notre introduction à ce domaine.

## 1.1 Les vents d’étoiles massives

Les étoiles massives possèdent un vent qui va leur enlever au cours de leur vie ( $\simeq 5 \text{ Man}$ ) plus de la moitié de leur masse d’origine, et ce, au cours de différentes phases. Lors de la phase OB, le vent est ténu et rapide avec un taux de perte de masse  $\dot{M} \simeq 10^{-7} \text{ M}_{\odot} \text{ an}^{-1}$  et des vitesses terminales de l’ordre de plusieurs milliers

de  $\text{km s}^{-1}$  (Prinja et al., 1990). Puis, lors de la phase Wolf-Rayet ( $\simeq 0.5 \text{ Man}$ ), il devient beaucoup plus dense ( $\dot{M} \simeq 10^{-5} \text{ M}_{\odot} \text{ an}^{-1}$ ,  $v_{\infty} \simeq 2000 \text{ km s}^{-1}$ ) et est alors optiquement épais (Crowther, 2007). Lors des éruptions LBV, le vent peut atteindre des états extrêmes avec des taux de perte de masse de  $1 \text{ M}_{\odot} \text{ an}^{-1}$  et des vitesses terminales très lentes de l'ordre de quelques centaines de  $\text{km s}^{-1}$ . C'est ce type d'éruption qu'a connu  $\eta$  Carinae en 1843, multipliant alors sa luminosité par plus d'un facteur 10 et devenant ainsi la deuxième étoile la plus brillante du ciel (Smith, 2008). Les étoiles entre 25 et  $30 \text{ M}_{\odot}$  passent, elles, plutôt par une phase supergéante rouge (RSG) (voir note de bas de page précédente), au cours de laquelle l'étoile se dilate jusqu'à atteindre plusieurs centaines de rayons solaires, avant d'atteindre le stade WN.

Le scénario d'évolution des étoiles massives accepté par le plus grand nombre des astronomes est actuellement le suivant (Crowther, 2007) :

**$M > 75 \text{ M}_{\odot}$  :** O  $\rightarrow$  WN (H-rich)  $\rightarrow$  LBV  $\rightarrow$  WN (H-poor)  $\rightarrow$  WC  $\rightarrow$  SNIc

**$40 < M < 75 \text{ M}_{\odot}$  :** O  $\rightarrow$  LBV  $\rightarrow$  WN (H-poor)  $\rightarrow$  WC  $\rightarrow$  SNIc

**$25 < M < 40 \text{ M}_{\odot}$  :** O  $\rightarrow$  LBV/RSG  $\rightarrow$  WN (H-poor)  $\rightarrow$  SNIb

Contrairement au soleil, où le vent est poussé par la pression thermique du gaz très chaud de la couronne, créée par des ondes acoustiques provenant de la convection à la surface du soleil, le vent des étoiles massives est poussé par la radiation. En effet, comme Saha l'a initialement suggéré en 1919, chaque photon de fréquence  $\nu$  transporte une quantité de mouvement donnée par (Lamers and Cassinelli, 1999) :

$$p = \frac{h\nu}{c} .$$

Cette pression de radiation est particulièrement efficace dans les raies des métaux présents dans le vent, principalement les raies de Si IV, C IV, N V et S VI. En effet, ces raies sont particulièrement abondantes dans le domaine UV, là où les étoiles massives, avec des températures effectives de surface (hydrostatique) allant de 10000-50000 K (étoiles OB) à 30000-100000 K (étoiles WR) ont leur maximum de flux (Crowther, 2007). La théorie décrivant les vents poussés par la radiation

(radiation-driven winds) a été initialement développée par Milne (1924) et Johnson (1925) puis, suite aux premières observations UV d'étoiles chaudes (Morton, 1967), améliorée par Lucy and Solomon (1970) et Castor et al. (1975). Pour les régions du vent appartenant au plus haut régime de températures, lorsque le gaz est complètement ionisé, c'est l'opacité due à la diffusion Thompson sur les électrons libres qui domine l'opacité totale (voir Fig. 1.1). Dans le régime non-relativiste, la section efficace de diffusion Thompson ne dépend pas de la longueur d'onde des photons et s'écrit (Lamers and Cassinelli, 1999) :

$$\sigma_T = \frac{1}{6\pi\epsilon_0^2} \left( \frac{e^2}{m_e c^2} \right)^2 = 6.65 \times 10^{-29} \text{ m}^2.$$

Cette opacité est d'autant plus dominante que la métallicité de l'étoile est faible, un cas extrême étant celui des premières populations d'étoiles, avec une métallicité nulle. Les travaux de Owocki and Gayley (1997) ont montré que l'opacité des électrons libres était le moteur dominant dans les régions internes, optiquement épaisses, du vent des étoiles Wolf-Rayet.

### 1.1.1 Instabilités dans les vents poussés par la radiation

Les vents poussés par la radiation que nous venons de décrire sont fondamentalement instables à des perturbations de vitesse (Fig. 1.2, Lamers and Cassinelli, 1999). En effet, une petite augmentation de vitesse d'une zone du vent va déplacer la fréquence d'absorption des raies hors de l'"ombre" de la matière sous-jacente, provoquant ainsi une augmentation de l'accélération car la matière est exposée à une plus grande quantité de flux photosphérique. Il a été démontré (Owocki and Rybicki, 1984) que cette instabilité se produisait lorsque la perturbation avait lieu sur une échelle de longueur plus petite que la longueur de Sobolev :

$$L_{\text{sobolev}} = \frac{v_{\text{thermal}}}{|dv/dr|}$$

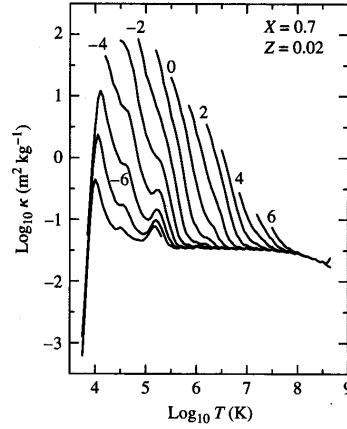


Figure 1.1 – Opacité de Rosseland pour un gaz composé de 70% d’hydrogène, de 28% d’hélium et de 2% de métaux en masse. Les courbes sont indexées par la valeur du logarithme de la densité ( $\log_{10}\rho$  en  $\text{kg}/\text{m}^3$ ). Pour les hautes températures, toutes les courbes convergent vers une valeur unique de l’opacité du fait de la prédominance croissante de la diffusion électronique. Le déclin des courbes d’opacité avant d’atteindre une valeur constante suit à peu près une loi de Kramers ( $\kappa \propto T^{-3.5}$ ), l’opacité est alors due principalement à l’absorption lié-libre et libre-libre des photons. (Tiré de Carroll and Ostlie 2006)

On pense que ce mécanisme est responsable de la création de “clumps” de gaz choqué à travers le volume du vent dans les étoiles OB et WR. L’émission en rayons X observée pour les étoiles OB montre des températures de  $\simeq 10^6$  K, représentatif de chocs peu forts (Lamers and Cassinelli, 1999).

## 1.2 Physique et observations des vents d’étoiles massives en collision

Il est maintenant établi qu’au moins 40% des étoiles massives appartiennent à des systèmes binaires (Foellmi, 2003). Lorsque les deux étoiles membres du système sont toutes les deux des étoiles massives dotées d’un vent, les vents vont entrer en collision et thermaliser ainsi leur énergie cinétique ( $v_\infty > 1000 \text{ km s}^{-1}$ )<sup>2</sup>, les températures atteintes par le gaz choqué sont de l’ordre de quelques millions de K, produisant un fort rayonnement X (thermique). La région de collision de vent, dont la forme est à peu près celle d’un cône d’ouverture  $2\theta$ , va également être

<sup>2</sup>Pour un vent de  $1000 \text{ km s}^{-1}$ , l’énergie de thermalisation est  $(1/2)m_p v^2 \simeq 5 \text{ keV}$

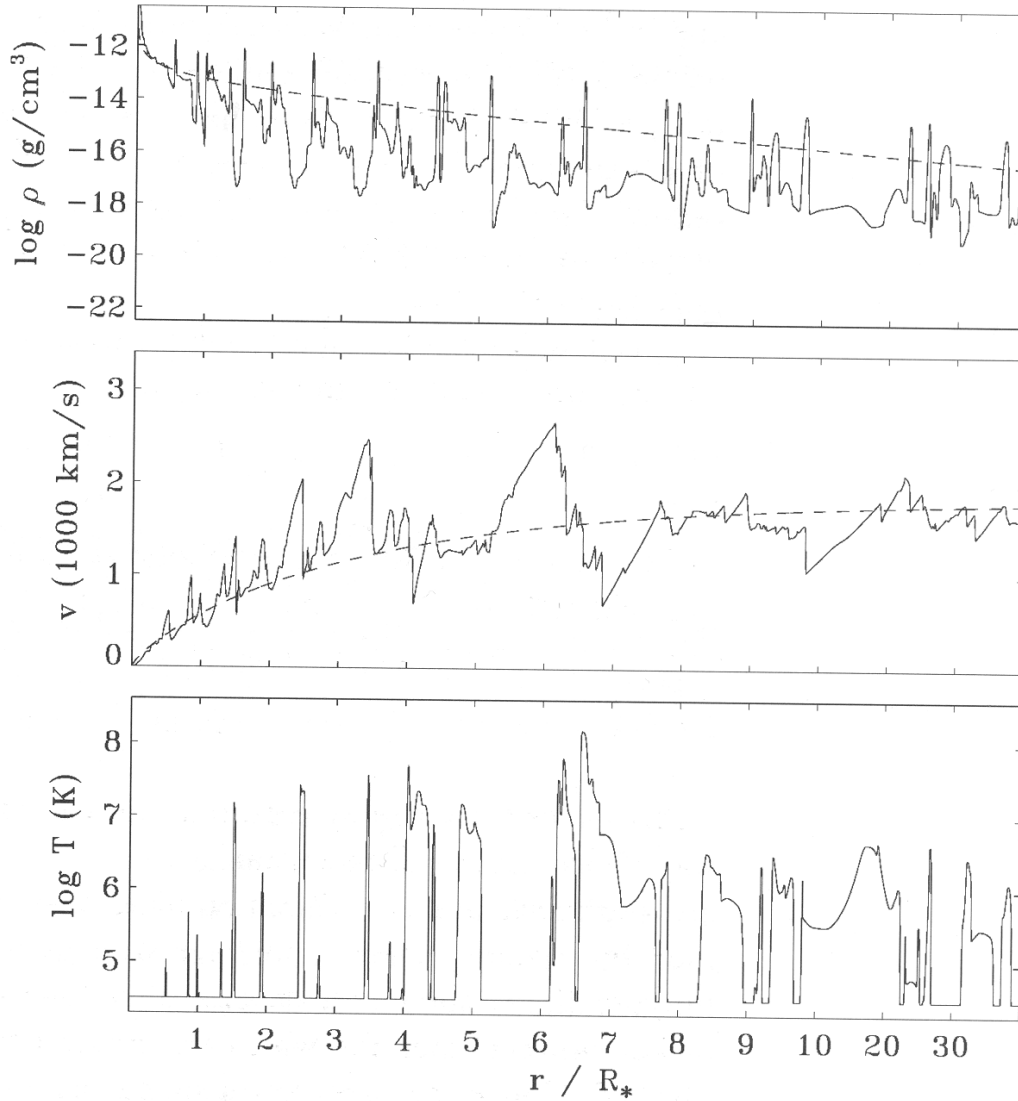


Figure 1.2 – Instabilités dues à une perturbation de vitesse dans un vent. Les zones plus accélérées vont sortir de l’“ombre” de la matière sous-jacente et absorber plus de flux, subissant alors un excès d’accélération radiative. Ce phénomène est à l’origine de chocs et de surdensités dans les vents d’étoiles chaudes. Les lignes en pointillés indiquent la structure du vent non perturbé. (Tiré de Lamers and Cassinelli 1999).



responsable d’une émission non-thermique. D’abord dans le domaine radio – les électrons relativistes accélérés par le mécanisme de Fermi vont, en présence d’un champ magnétique, émettre un rayonnement synchrotron. Puis dans le domaine X – cette même population d’électrons relativistes peut, par diffusion Compton inverse sur les photons UV, produire des rayons X non-thermiques et même des rayons gamma.

### 1.2.1 Géométrie d’un système à collision de vents

Nous décrivons ici une dérivation analytique de la forme de la surface de contact des deux vents, dans le cas purement radiatif, en suivant la démarche décrite dans Cantó et al. (1996). Cette démarche inclut, en plus de la simple égalité des pressions dynamiques des deux vents ( $\rho_A \cdot v_A^\infty = \rho_B \cdot v_B^\infty$  – approche de Dyson et al. 1993), les équations de continuité de la masse et de conservation de la quantité de mouvement et du moment cinétique. La prise en compte de ces équations introduit, comme nous le verrons, un facteur de “pression centrifuge”, qui va avoir tendance à augmenter l’angle d’ouverture du cône de choc. Le cas purement adiabatique est décrit dans la sous-section suivante en suivant l’approche de Gayley (2009).

#### 1.2.1.1 Le paramètre $\eta$

Considérons un système binaire formé de deux étoiles A et B séparées par une distance  $D(t)$ , le paramètre fondamental qui va définir la géométrie de la région de collision des vents (RCV) est le rapport des quantités de mouvement  $\eta$  :

$$\eta = \frac{\dot{M}_A \cdot v_A^\infty}{\dot{M}_B \cdot v_B^\infty}. \quad (1.1)$$

- Si  $\eta > 1$  → Le cône de choc sera concave autour de l'étoile B  
 Si  $\eta < 1$  → Le cône de choc sera concave autour de l'étoile A  
 Si  $\eta = 1$  → Dans le cas idéal (pas de forces de Coriolis),  
 le choc formera un plan médian entre A et B

La distance du centre de l'étoile A au point de stagnation le long de l'axe reliant A et B est donnée par :

$$R_0 = D \frac{\sqrt{\eta}}{1 + \sqrt{\eta}} . \quad (1.2)$$

### 1.2.2 Équation algébrique de la surface de contact

On considère deux écoulements supersoniques non-accélérés entrant en collision le long d'une coquille fine d'équation  $R(\theta)$ . On suppose que le problème est de symétrie cylindrique sans composante azimutale de la vitesse. La vitesse le long de la surface de contact, supposée unique, s'écrit donc :

$$\bar{\mathbf{v}} = v_r \hat{\mathbf{r}} + v_z \hat{\mathbf{z}} , \quad (1.3)$$

où  $\hat{\mathbf{r}}$  et  $\hat{\mathbf{z}}$  sont les vecteurs unitaires perpendiculaire et parallèle à l'axe AB (Fig. 1.3).

Considérons maintenant une ligne de contrôle partant de A et faisant un angle  $\theta$  avec l'axe AB (Fig. 1.3), on définit les flux suivants à travers la ligne de contrôle :

$$\dot{M}(\theta) \quad \text{--} \quad \text{Flux de masse} \quad (1.4)$$

$$\dot{\Pi}_r(\theta) \quad \text{--} \quad \text{Flux de la composante r de la qté de mvt (= } \dot{M}v_r) \quad (1.5)$$

$$\dot{\Pi}_z(\theta) \quad \text{--} \quad \text{Flux de la composante z de la qté de mvt (= } \dot{M}v_z) \quad (1.6)$$

$$\dot{J}(\theta) \quad \text{--} \quad \text{Flux de moment cinétique (= } \dot{M}v_\theta R(\theta)) \quad (1.7)$$

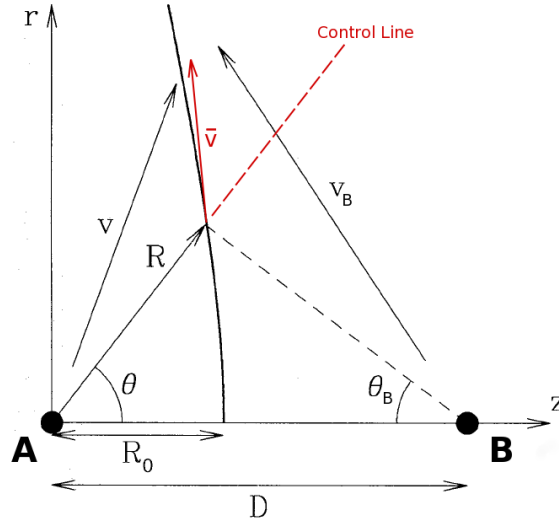


Figure 1.3 – Paramétrisation utilisée pour le problème de l'interaction de deux vents sphériques.  $R_0$  est le lieu d'égalité des pressions dynamiques et est donné par  $R_0 = (\sqrt{\eta}D)/(1 + \sqrt{\eta})$ . (Tiré de Cantó et al. 1996)

Ces quantités sont reliées entre elles par les relations :

$$\dot{\Pi}_r(\theta)\hat{\mathbf{r}} + \dot{\Pi}_z(\theta)\hat{\mathbf{z}} = \dot{M}(\theta)[v_r(\theta)\hat{\mathbf{r}} + v_z(\theta)\hat{\mathbf{z}}] \quad (1.8)$$

$$J(\theta) = \dot{M}(\theta)v_\theta(\theta)R(\theta) , \quad (1.9)$$

où  $v_\theta$  est donné par :

$$v_\theta = v_r \cos \theta - v_z \sin \theta . \quad (1.10)$$

Définissons aussi les taux de variation de ces mêmes quantités, injectés par les vents A et B dans la portion de la surface de contact de 0 à  $\theta$ . Soit :

$$\begin{aligned} \dot{M}_A(\theta), \dot{\Pi}_{rA}(\theta), \dot{\Pi}_{zA}(\theta) \text{ et } J_A(\theta) & \quad \text{pour le vent A,} \\ \dot{M}_B(\theta), \dot{\Pi}_{rB}(\theta), \dot{\Pi}_{zB}(\theta) \text{ et } J_B(\theta) & \quad \text{pour le vent B.} \end{aligned}$$

La condition de stationnarité du système, exigeant que ce qui traverse la ligne de contrôle par seconde égale ce qui est injecté par les vents entre 0 et  $\theta$ , nous

amène au système d'équations suivant :

$$\dot{\Pi}_r(\boldsymbol{\theta})\hat{\mathbf{r}} + \dot{\Pi}_z(\boldsymbol{\theta})\hat{\mathbf{z}} = [\dot{\Pi}_{rA}(\boldsymbol{\theta}) + \dot{\Pi}_{rB}(\boldsymbol{\theta})]\hat{\mathbf{r}} + [\dot{\Pi}_{zA}(\boldsymbol{\theta}) + \dot{\Pi}_{zB}(\boldsymbol{\theta})]\hat{\mathbf{z}} \quad (1.11)$$

$$\dot{J}(\boldsymbol{\theta}) = \dot{J}_A(\boldsymbol{\theta}) + \dot{J}_B(\boldsymbol{\theta}) \quad (1.12)$$

$$\dot{M}(\boldsymbol{\theta}) = \dot{M}_A(\boldsymbol{\theta}) + \dot{M}_B(\boldsymbol{\theta}) . \quad (1.13)$$

À l'aide de 1.8 et 1.9, les relations 1.11 et 1.12 deviennent :

$$\dot{M}(v_r\hat{\mathbf{r}} + v_z\hat{\mathbf{z}}) = (\dot{\Pi}_{rA} + \dot{\Pi}_{rB})\hat{\mathbf{r}} + (\dot{\Pi}_{zA} + \dot{\Pi}_{zB})\hat{\mathbf{z}} \quad (1.14)$$

$$\dot{M}v_\theta R = \dot{J}_A + \dot{J}_B . \quad (1.15)$$

Ceci, en utilisant 1.10, nous amène à la relation recherchée, à savoir une équation reliant la position du choc  $R$  aux paramètres des vents :

$$R = \frac{\dot{J}_A + \dot{J}_B}{(\dot{\Pi}_{rA} + \dot{\Pi}_{rB})\cos \boldsymbol{\theta} + (\dot{\Pi}_{zA} + \dot{\Pi}_{zB})\sin \boldsymbol{\theta}} z . \quad (1.16)$$

Une fois que la position du choc  $R(\boldsymbol{\theta})$  est déterminée, le flux de masse  $\dot{M}$  est donné par l'équation 1.13, la vitesse de l'écoulement par 1.14 et 1.15 et la densité surfacique de masse  $\boldsymbol{\sigma}$  dans la couche par :

$$\dot{M} = 2\pi(R \sin \boldsymbol{\theta})\boldsymbol{\sigma}v . \quad (1.17)$$

### Interaction vent sphérique/vent sphérique

On considère maintenant le cas à deux vents sphériques. On note  $\dot{M}_A^0$  et  $\dot{M}_B^0$  les taux de perte de masse respectifs des étoiles A et B, et  $\boldsymbol{\theta}_B$ , l'angle mesurant la position d'un point de la surface de contact (le même que celui repéré par  $\boldsymbol{\theta}$ ) à partir de l'étoile B (voir Fig. 1.3).

La solution s'écrit alors (Cantó et al., 1996) :

$$\boldsymbol{\theta}_B \cot \boldsymbol{\theta}_B = 1 + \eta(\boldsymbol{\theta} \cot \boldsymbol{\theta} - 1) , \quad (1.18)$$

où  $\eta$  est défini par l'équation 1.1. Cette équation peut être résolue numériquement pour une valeur donnée de  $\theta$ , la valeur de  $R$  correspondante est alors donnée par la relation géométrique :

$$R = D \sin \theta_B \csc (\theta + \theta_B) \quad (1.19)$$

Une solution explicite approchée peut être obtenue pour  $\theta_B$  petit :

$$\theta_B \simeq \left\{ \frac{15}{2} \left[ -1 + \sqrt{1 + \frac{4}{5} \eta (1 - \theta \cot \theta)} \right] \right\}^{1/2} \quad (1.20)$$

La forme des solutions implicite et explicite est tracée à la Fig. 1.4 pour plusieurs valeurs de  $\eta$ .

### Angle d'ouverture asymptotique du cône de choc (cas purement radiatif)

L'angle d'ouverture du cône de choc  $\theta_\infty$  est un paramètre essentiel pour contraindre la valeur de  $\eta$  par les observations. Ce paramètre est obtenu à partir de l'équation 1.18 en appliquant la condition  $\theta_\infty + \theta_{B_\infty} = \pi$ , ce qui donne :

$$\theta_\infty - \tan \theta_\infty = \frac{\pi}{1 - \eta} . \quad (1.21)$$

Le graphique de  $\theta_\infty$  en fonction de  $\eta$  est donné à la Fig. 1.5. L'angle d'ouverture obtenu grâce à cette démarche est systématiquement plus grand que lorsqu'il est calculé simplement avec des considérations d'égalité des pressions dynamiques (Dyson et al. 1993, voir Fig.1.6) . Cet effet est dû à la prise en compte d'un effet de pression centrifuge occasionnée par l'écoulement de gaz le long de la surface de contact. Cet effet, également décrit dans Cantó et al. (1998), peut être vu comme la tendance du fluide à maintenir sa trajectoire rectiligne et uniforme, du fait de son inertie.

L'angle d'ouverture du cône peut également être augmenté du fait du freinage

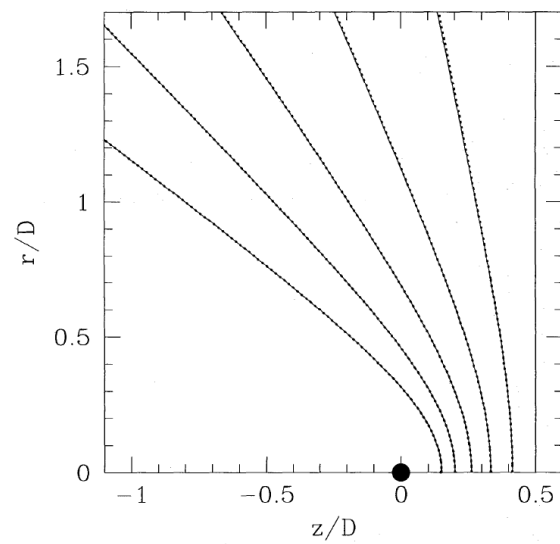


Figure 1.4 – Solution exacte implicite sans tenir compte de la rotation du système (ligne pleine), et solution explicite approchée (ligne pointillée) pour le problème d'interaction de 2 vents sphériques pour plusieurs valeurs du paramètre  $\eta$  : 1 (verticale), 0.5, 0.25, 0.125, 0.0625, et 0.03125. Les positions des étoiles sont à  $z/D = 0$  et  $z/D = 1$ . (Tiré de Cantó et al. 1996)

radiatif du vent d'une des étoiles par le rayonnement de son compagnon (Gayley et al., 1997, Owocki and Gayley, 1995). Ce freinage peut aller jusqu'à provoquer le quasi-arrêt d'un des vents et se manifeste observationnellement par un déficit de l'émission X par rapport aux valeurs attendues.

**(cas purement adiabatique)**

Lorsque le refroidissement de la région de collision des vents se fait de manière purement adiabatique, on a conservation de l'énergie (énergie thermique du gaz + énergie cinétique de l'écoulement). Une fois que le gaz est complètement refroidit, l'énergie cinétique, et donc la vitesse de l'écoulement, revient alors à sa valeur d'origine. On a également une augmentation de la température et donc de la pression du fait de l'absence de refroidissement radiatif significatif, ce qui a pour effet une expansion explosive du cône de choc, i.e. une augmentation de son angle d'ouverture. Gayley (2009) distingue deux situations pour ce type de choc : celle dans laquelle il n'y a pas de mélange des écoulements dans la RCV, chaque écoulement retrouve alors sa vitesse d'écoulement d'origine de part et d'autre de la surface de contact et on obtient alors la relation :

$$\eta = \tan^4 \left( \frac{\theta}{2} \right). \quad (1.22)$$

Et celle dans laquelle il y a un mélange homogène du gaz dans la RCV. On a alors une vitesse d'écoulement unique et la relation suivante entre  $\eta$  et  $\theta$  :

$$\eta = \frac{\cos^2 \theta}{8u(3 \cos \theta - 1)} \sec^6 \left( \frac{\theta}{2} \right) [2(1 + u^2) \cos^2 \theta - 2(1 + u^2) - u \sin^2 \theta \tan^2 \theta + \sqrt{2} \sqrt{1 + u + 4u^2 + u^3 + u^4 + (1 - u)^2(1 + u + u^2) \cos(2\theta) \sin \theta \tan \theta}] \quad (1.23)$$

où  $u = v_A/v_B$ .

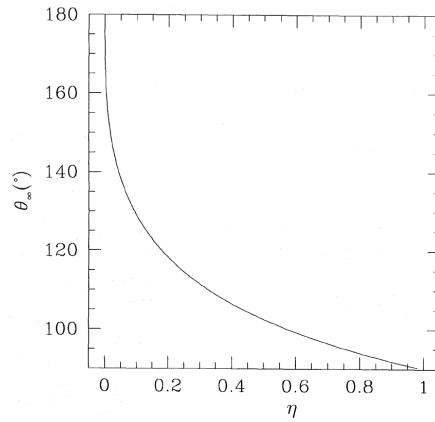


Figure 1.5 – Angle d’ouverture asymptotique du cône de choc en fonction du paramètre  $\eta$  pour la solution de Cantó et al. (1996). Pour  $\eta$  tendant vers 0, l’angle d’ouverture du cône tend vers  $180^\circ$ , ce qui correspond à un cône parallèle à l’axe des étoiles. (Tiré de Cantó et al. 1996)

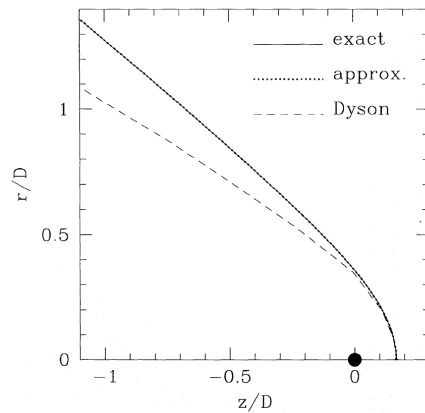


Figure 1.6 – Comparaison, pour  $\eta = 0.04$ , des solutions pour la surface de contact de Dyson et al. (1993) (trait en tirets), basée uniquement sur l’égalité des pressions dynamiques des deux vents, et de Cantó et al. (1996) (solutions exacte, trait plein, et approchée, pointillés), incluant l’effet de la pression centrifuge due à l’écoulement le long de la surface. (Tiré de Cantó et al. 1996)



### 1.2.3 Physique des chocs de la zone de collision des vents

Dans cette section, nous suivons l'approche de Eichler and Usov (1993) qui proposent une description de la physique de la région de collision des vents. Deux cônes de choc  $\tilde{S}_A$  et  $\tilde{S}_B$  correspondant aux vents de l'étoile A et de l'étoile B vont se former de part et d'autre de la surface de contact, dont nous venons de calculer une solution. Ces deux cônes de choc possèdent chacun un angle d'ouverture  $\theta_A$  et  $\theta_B$ <sup>3</sup> comme cela est illustré à la Fig. 1.7 où  $\Delta\theta = |\theta_A - \theta_B|$ .

Les conditions physiques de part et d'autre de la discontinuité sont données par les relations de Rankine-Hugoniot :

$$\rho_1 v_{1\perp} = \rho_2 v_{2\perp} \quad v_{1\parallel} = v_{2\parallel} \quad (1.24)$$

$$P_1 + \rho_1 v_{1\perp}^2 = P_2 + \rho_2 v_{2\perp}^2 \quad (1.25)$$

$$H_1 + \frac{1}{2} v_{1\perp}^2 = H_2 + \frac{1}{2} v_{2\perp}^2 \quad (1.26)$$

où  $H$  est l'enthalpie spécifique définie par :

$$H = \frac{\gamma}{\gamma - 1} \frac{P}{\rho},$$

avec  $\gamma$  l'indice polytropique du gaz ( $P \propto \rho^\gamma$ ). Le milieu amont (pré-choc) est noté 1, et  $v_\perp$  et  $v_\parallel$  sont les composantes perpendiculaire et parallèle au choc de la vitesse. La force du choc est caractérisée par le rapport de compression :

$$\frac{\rho_2}{\rho_1} = r \quad (1.27)$$

et le rapport des vitesses s'exprime en fonction de ce rapport :

$$\frac{v_2}{v_1} = \frac{1}{r}. \quad (1.28)$$

L'expression de ce rapport en fonction de la vitesse amont peut être obtenue

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<sup>3</sup>Notez bien qu'ici,  $\theta_B$  ne désigne plus la même quantité qu'à la section précédente.

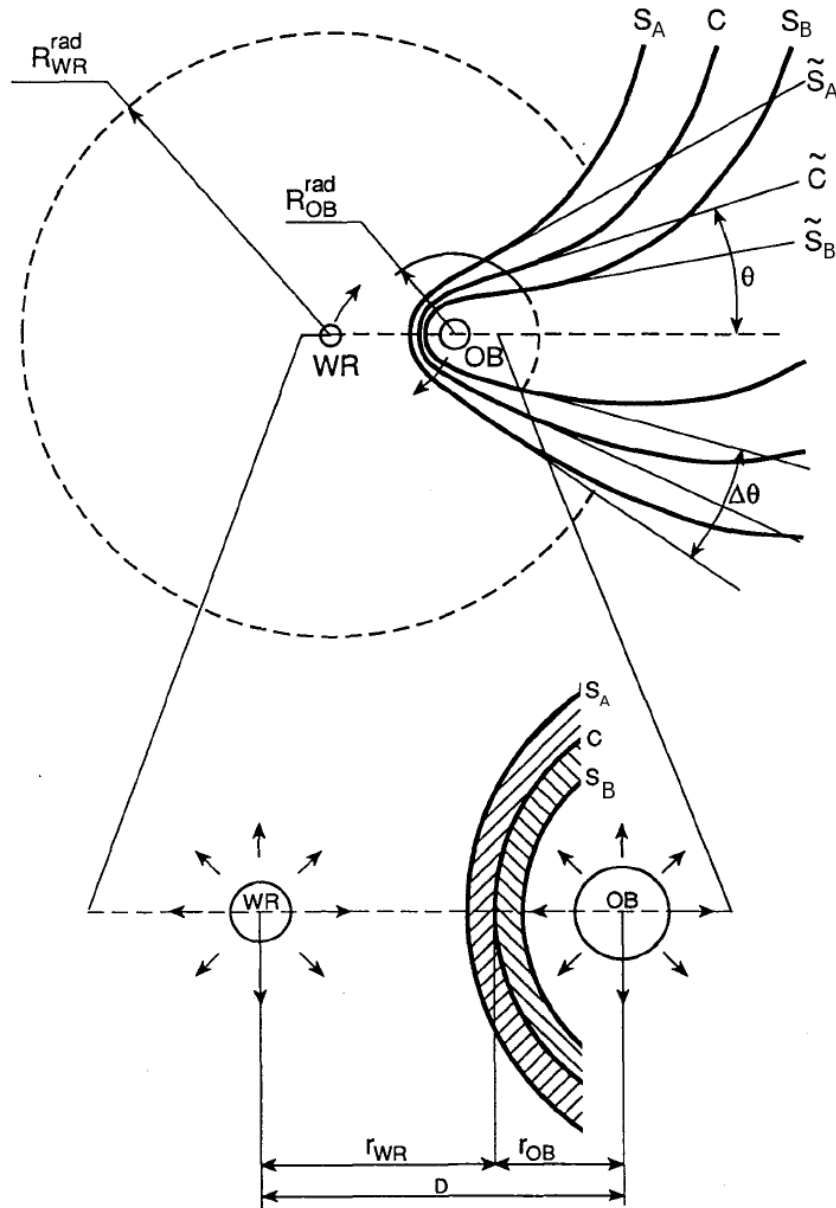


Figure 1.7 – Géométrie et paramétrisation des surfaces de chocs dans un système WR+OB. Les surfaces  $\tilde{C}$ ,  $\tilde{S}_A$  et  $\tilde{S}_B$  sont les surfaces de contact et de choc lorsque l'on néglige les forces de Coriolis.  $C$ ,  $S_A$  et  $S_B$  sont ces mêmes surfaces lorsque l'on prend en compte les forces de Coriolis.  $R_{WR}^{rad}$  et  $R_{OB}^{rad}$  sont les rayons des photosphères des étoiles dans le domaine radio. (Tiré de Eichler and Usov 1993)

grâce à un remaniement des relations de Rankine-Hugoniot :

$$r = \frac{\rho_2}{\rho_1} = \frac{(\gamma+1)M_1^2}{2 + (\gamma-1)M_1^2} \quad (1.29)$$

$$\frac{P_2}{P_1} = \frac{2\gamma M_1^2 - (\gamma-1)}{\gamma+1} \quad (1.30)$$

où  $M_1$  est le nombre de Mach de l'écoulement en amont.

Pour un choc fort, c'est à dire pour  $v_1$  beaucoup plus grand que la vitesse du son ( $c_s = \sqrt{\gamma P/\rho}$ ), ce rapport tend vers la valeur limite :

$$r = \frac{\gamma+1}{\gamma-1}, \quad (1.31)$$

soit  $r=4$  pour un choc adiabatique et un gaz parfait mono-atomique ( $\gamma = 5/3$ ) et  $r = \infty$  pour un choc isotherme ( $\gamma = 1$ ), représentant assez bien un choc radiatif, pour lequel le milieu choqué se refroidit très vite par rayonnement. Dans ce dernier cas, la matière post-choc est comprimée dans une couche très dense de vitesse nulle par rapport au choc. Dans tout ce qui suit, nous considérerons toujours l'hypothèse du choc fort, qui est largement vérifiée pour les écoulements très supersoniques que nous avons à considérer.

### 1.2.3.1 Propriétés générales

Les conditions pré-choc sont typiquement (Eichler and Usov, 1993) :

$$\begin{aligned} T &\simeq (2-4) \times 10^4 \text{ K} \\ v &\simeq (2-3) \times 10^3 \text{ km s}^{-1}, \end{aligned}$$

ce qui donne dans le cas d'un choc fort adiabatique des conditions post-chocs données par 1.27 et 1.28 avec  $r=4$  et la loi des gaz parfaits :

$$P = \frac{\rho k T}{\mu m_p}, \quad (1.32)$$

où  $\mu$  est la masse moléculaire moyenne. On obtient notamment pour la température :

$$T = \frac{3\mu m_p v_1^2}{16k},$$

soit

$$T \simeq 1.36 \times 10^7 \left( \frac{v_1}{10^8 \text{ cm s}^{-1}} \right)^2 \text{ K}$$

pour des abondances solaires. La température sera plus élevée pour des abondances supérieures aux abondances solaires puisque  $\mu$  sera plus grand. Après le choc, l'écoulement est contrôlé par la pression thermique et par la quantité de mouvement restante. L'importance de cette dernière composante augmente lorsque l'on s'éloigne du point de stagnation, du fait de la plus grande obliquité du choc, et donc des vitesses transverses plus grandes. Pour des binaires très rapprochées, il n'existera pas de position d'équilibre pour le choc et le vent le plus fort (celui de la Wolf-Rayet) va s'écraser directement sur la photosphère de son compagnon (l'étoile OB).

### 1.2.3.2 Le refroidissement radiatif

À des températures de l'ordre de  $10^7$  K, le gaz est presque totalement ionisé, le mécanisme de radiation est alors dominé par le rayonnement bremsstrahlung (libre-libre) des électrons libres déviés par les champs électriques des noyaux. Le taux de perte d'énergie par unité de volume par rayonnement libre-libre  $Q$  a été calculé par Usov (1992) :

$$Q = N_i N_e \Lambda(T) = \int_0^\infty Q_\nu d\nu = C_1 N_i N_e Z^2 T^{1/2} \text{ erg s}^{-1} \text{ cm}^{-3}, \quad (1.33)$$

où  $C_1 = 1.4 \times 10^{-27} g_0(T)$ , et  $g_0(T) \simeq 1.2$  est le facteur de gaunt moyenné en temps qui varie peu avec la température.  $Z$  est le nombre de charge des atomes,  $N$  ( $=N_i + N_e$ ),  $N_i$  et  $N_e$  sont en  $\text{cm}^{-3}$ . Ce type de refroidissement est donc très sensible à la densité du gaz ( $Q \propto \rho^2 \sqrt{T}$ ) et relativement peu sensible à la température. De plus, si l'on considère le taux de refroidissement total, tous mécanismes confondus, pour

une même densité, il est intéressant de noter que les températures aux alentours de  $3 \times 10^7$  K constituent un minimum, comme le montre la figure 1.8 pour 3 abondances différentes.

### 1.2.3.3 Le paramètre $\chi$

Pour quantifier l'importance du refroidissement dans la zone de collision des vents, il est pratique de définir une échelle caractéristique de temps, dont une définition est donnée par Stevens et al. (1992) :

$$t_{\text{cool}} = \frac{kT_s}{4N_w\Lambda(T_s)} , \quad (1.34)$$

où  $T_s$  est la température du gaz choqué,  $N_w$  est la densité en nombre du vent avant le choc et  $\Lambda(T_s)$  est le taux de refroidissement défini par 1.33. Ce temps caractéristique est à comparer avec le temps caractéristique d'échappement du système, également estimé dans Stevens et al. (1992) pour une portion de gaz située près de l'axe reliant les deux étoiles :

$$t_{\text{esc}} = \frac{d}{c_s} . \quad (1.35)$$

On a fait ici l'hypothèse (légitimée par les modèles) que la vitesse de l'écoulement du gaz choqué près de l'axe AB est proche de la vitesse du son ( $c_s$ ).  $d$  est la distance de l'étoile à la surface de contact.

On peut alors définir un paramètre  $\chi$  quantifiant l'importance du refroidissement radiatif sur la dynamique du système :

$$\chi = \frac{t_{\text{cool}}}{t_{\text{esc}}} . \quad (1.36)$$

Pour  $\chi > 1$ , le choc sera adiabatique tandis que pour  $\chi \ll 1$ , il sera isotherme. Une estimation utile dans le cas d'un système WR+OB peut être obtenue en

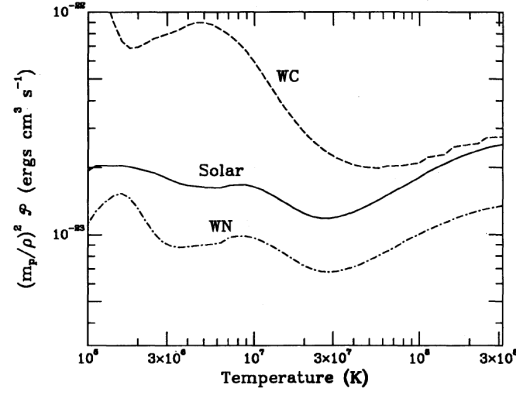


Figure 1.8 – Taux de refroidissement total par émission X ( $h\nu$  compris entre 0.1 et 30 keV) en fonction de la température pour trois abondances différentes : Solaire, WN et WC. (Tiré de Stevens et al. 1992)

négligeant la dépendance en température de  $\Lambda(T)$ , ce qui donne :

$$\chi \simeq \left( \frac{v}{1000 \text{ km s}^{-1}} \right) \left( \frac{d}{10^{12} \text{ cm}} \right) \left( \frac{\dot{M}}{10^{-7} M_{\odot} \text{ an}^{-1}} \right)^{-1}, \quad (1.37)$$

où  $v$  est la vitesse du vent et  $d$  est la distance à la surface de contact. Cette relation n'est valide que pour une gamme restreinte de valeurs de  $v$  autour de  $(1-2) \times 10^3 \text{ km s}^{-1}$  (typiques des systèmes considérés) du fait que l'on ait négligé la dépendance en température (et donc en vitesse) de  $\Lambda$ . Néanmoins, dans son domaine de validité, cette relation montre que  $\chi$  est proportionnel à  $d$  (et donc à  $P^{2/3}$  grâce à la 3e loi de Kepler), soit : plus la période de la binaire est grande, plus le choc sera adiabatique. Ce comportement est dû à l'effet dominant de la densité sur l'émissivité. En fait, pour des périodes supérieures à quelques jours, le choc sera largement adiabatique, même si, dans le détail, la limite exacte entre adiabatique et isotherme dépend très fortement des paramètres du vent ( $\dot{M}, v_w$ ). Notamment, les binaires comportant des WR vont avoir des chocs radiatifs pour des périodes plus longues que les binaires comportant seulement des étoiles OB.

Le refroidissement radiatif va aussi avoir une importance prédominante dans les régions obliques du choc (éloignées de l'axe AB) pour lesquelles la vitesse normale au choc est très inférieure à  $(1-2) \times 10^3 \text{ km s}^{-1}$ . Cet effet est dû à

la rapide augmentation du taux de refroidissement total lorsque  $T$  diminue. La relation 1.34 n'est alors plus dans son domaine de validité et sous-estime largement le taux de refroidissement dans ce cas. Un cas critique de choc oblique se produit notamment lorsque le vent de la WR s'écrase directement sur la photosphère de l'étoile OB.

L'effet du paramètre  $\chi$  sur la géométrie du choc est montré à la Fig. 1.9 où on voit que, comme attendu, plus le choc est radiatif, plus la couche de gaz choqué devient fine.

#### 1.2.3.4 Instabilités

La zone de collision des vents est sujette à plusieurs instabilités dynamiques, et ce, d'autant plus que les chocs sont radiatifs. Des simulations ont été effectuées par Stevens et al. (1992) pour 3 configurations "choc 1/choc 2" : adiabatique/adiabatique, adiabatique/radiatif, et radiatif/radiatif. Les résultats de ces simulations sont montrés à la figure 1.10.

Trois types d'instabilités ont pu être identifiées dans les simulations :

**L'instabilité de Kelvin-Helmoltz.** Cette instabilité a lieu à la surface de contact de deux écoulements parallèles de vitesses différentes  $u_1$  et  $u_2$ . Cette instabilité est caractérisée par son temps caractéristique de croissance théorique. Ce temps est, dans le cas linéaire :

$$t_{\text{KH}} = \frac{\lambda}{u_1 - u_2} \left[ \frac{\rho_1 + \rho_2}{2\pi(\rho_1\rho_2)^{1/2}} \right], \quad (1.38)$$

où  $\lambda$  est la longueur d'onde de la perturbation. Lorsque l'instabilité devient non-linéaire, ce temps caractéristique doit être remplacé par le temps caractéristique de l'écoulement  $D/\nu$ . Un exemple d'instabilité de Kelvin-Helmoltz est montré à la Fig. 1.10(a). Cette instabilité est aussi présente dans les simulations (b) et (c) mais est mélangée à l'instabilité "thin-shell" décrite plus loin.

**L'instabilité thermique.** Ce type d'instabilité est potentiellement à attendre dans tout fluide astrophysique impliquant un phénomène de refroidissement dé-

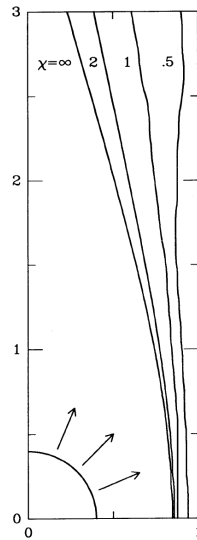


Figure 1.9 – Position du choc pour deux vents de moments identiques pour différentes valeurs du paramètre de refroidissement  $\chi$ . (Tiré de Stevens et al. 1992)

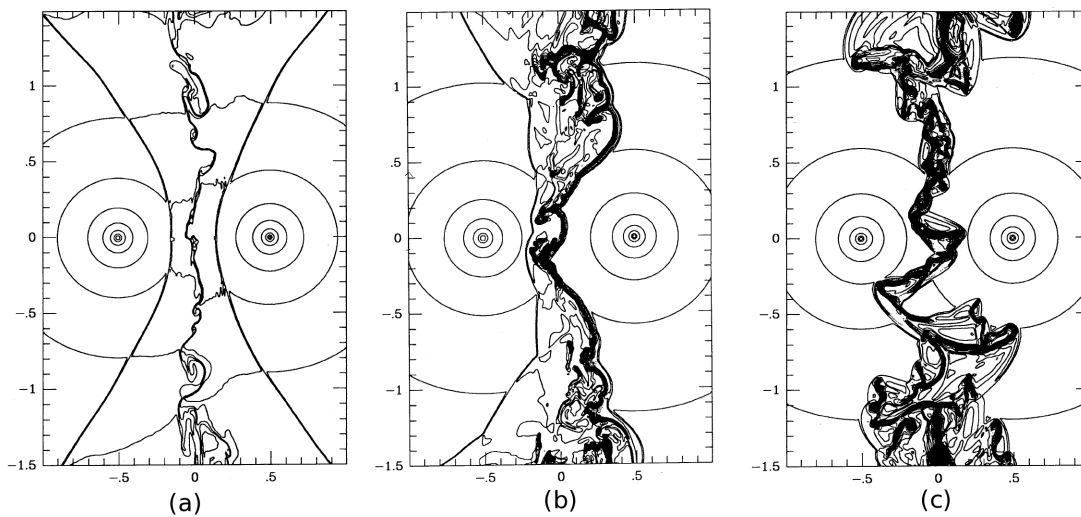


Figure 1.10 – Simulations de chocs pour des vents sphériques de mêmes moments  $\dot{M}v_\infty$  mais ayant des vitesses terminales différentes. Trois configurations “choc 1/choc 2” sont simulées : (a) adiabatique/adiabatique, (b) adiabatique/radiatif, et (c) radiatif/radiatif. Le cas (a) présente seulement l’instabilité de Kelvin-Helmholtz du fait de la différence de vitesse de part et d’autre de la surface de contact, tandis que les cas (b) et (c) présentent également l’instabilité “thin-shell”. (Tiré de Stevens et al. 1992)



pendant de la densité et de la température. Cette instabilité est généralement responsable de condensations locales où d’oscillations temporelles. C’est ce dernier phénomène qui est mis en évidence dans les travaux de Chevalier and Imamura (1982) et de Imamura et al. (1984) sur les chocs radiatifs dans lesquels on montre que l’utilisation d’une fonction de refroidissement  $\Lambda \propto T^\alpha$  engendre une instabilité pour  $\alpha$  plus petit qu’une valeur critique d’environ 0.4.

**L’instabilité “thin-shell”.** Les simulations montrent que la coquille de gaz dense et froid formée dans un choc radiatif est dynamiquement instable et est très rapidement détruite dans l’écoulement (Fig. 1.10 (b) et (c)). L’amplitude de cette instabilité est encore plus importante dans le cas radiatif/radiatif.

Les conséquences attendues des instabilités de K-H et T-S sont une augmentation du volume de gaz émettant en X, ainsi qu’une variabilité temporelle sur un temps caractéristique d’environ  $D/v$ .

#### 1.2.4 Emission optique

L’élément le plus marquant du spectre optique (et UV) d’un système binaire WR+OB est la superposition sur ses raies d’émission d’un excès d’émission se déplaçant en fréquence au cours de l’orbite, et ce de manière périodique (Bartzakos et al., 2001, Hill et al., 2000). Ce phénomène est particulièrement visible sur les raies de la WR à profil plat, typiques de l’émission par une coquille de gaz optiquement mince et d’émissivité constante en expansion avec une certaine vitesse  $v_{exp}$ . Au lieu d’un excès d’émission, on observe parfois une lacune d’émission présentant le même comportement verrouillé en phase. Ces deux phénomènes se comprennent assez facilement de manière qualitative et sont illustrés à la Fig. 1.11.

Dans le premier cas, on explique l’excès d’émission comme se formant dans le cône de choc, à un endroit où le gaz a pu se refroidir jusqu’à une température permettant la formation de la raie considérée. Cette raie présentera donc une composante s’approchant de nous (décalée vers le bleu), et une composante s’éloignant de nous (décalée vers le rouge). À deux moments dans l’orbite du système

(les phases de conjonction : = 0 et 0.5 par définition pour une orbite circulaire), les décalages en fréquence des excès d'émission seront égaux, et on observera un seul excès, plus grand, décalé de  $v \cos(\theta)$  (phase 0) et  $-v \cos(\theta)$  (phase 0.5), où  $2\theta$  est l'angle d'ouverture du cône de choc. Nous avons ici considéré un angle d'inclinaison  $i = 90^\circ$ . Bien sûr, la région d'émission n'est pas constituée seulement de deux points mais d'un anneau sur le cône de choc, cependant, comme cela est démontré dans l'annexe I, le profil de raie dû à un anneau en expansion présente quand même deux maximums aux fréquences  $-v_{\text{exp}}$  et  $+v_{\text{exp}}$ . Lorsque l'on prend en compte l'inclinaison non-nulle du système, l'addition des excès d'émission aux phases de conjonction n'aura plus lieu mais sera remplacé par un minimum de séparation de  $2 \times v \sin(\theta) \cos(i)$ . La mesure de ce minimum nous donnera accès à l'inclinaison du système. Tous ces paramètres peuvent être pris en compte dans un modèle de raie d'émission tel que celui développé par Lührs (1997) que nous décrivons à la section 1.2.4.1.

Dans le cas du déficit d'émission de la WR, l'explication vient en considérant un trou circulaire dans la coquille d'émission correspondant à l'intersection entre cette dernière et le cône de choc (Voir Fig. 1.11). L'excès d'émission et le déficit peuvent dans certains cas s'additionner sur le même profil d'émission.

#### 1.2.4.1 Modèle de raie d'émission

Le modèle de Lührs (1997) a été construit en se basant sur des observations de la raie à profil plat C III 5695.9 Å du système binaire massif HD152270 (WR 79). Il s'agit d'un modèle géométrique prenant en compte le demi-angle d'ouverture du cône de choc,  $\theta$ , l'épaisseur de ce cône,  $\delta\theta$ , la vitesse de l'écoulement le long du choc (supposée constante),  $v_{\text{strm}}$ , l'inclinaison du plan orbital,  $i$ , la phase d'observation,  $\phi$ , et une correction de la phase modélisant l'effet des forces de Coriolis dues au mouvement orbital,  $\delta\phi$  ( $\phi^* = \phi - \delta\phi$ ). La configuration du modèle est illustrée à la Fig. 1.12(a). Lührs paramétrise également la distribution en densité dans l'épaisseur du choc avec un polynôme d'ordre 2 en  $\cos(\theta)$  s'annulant aux

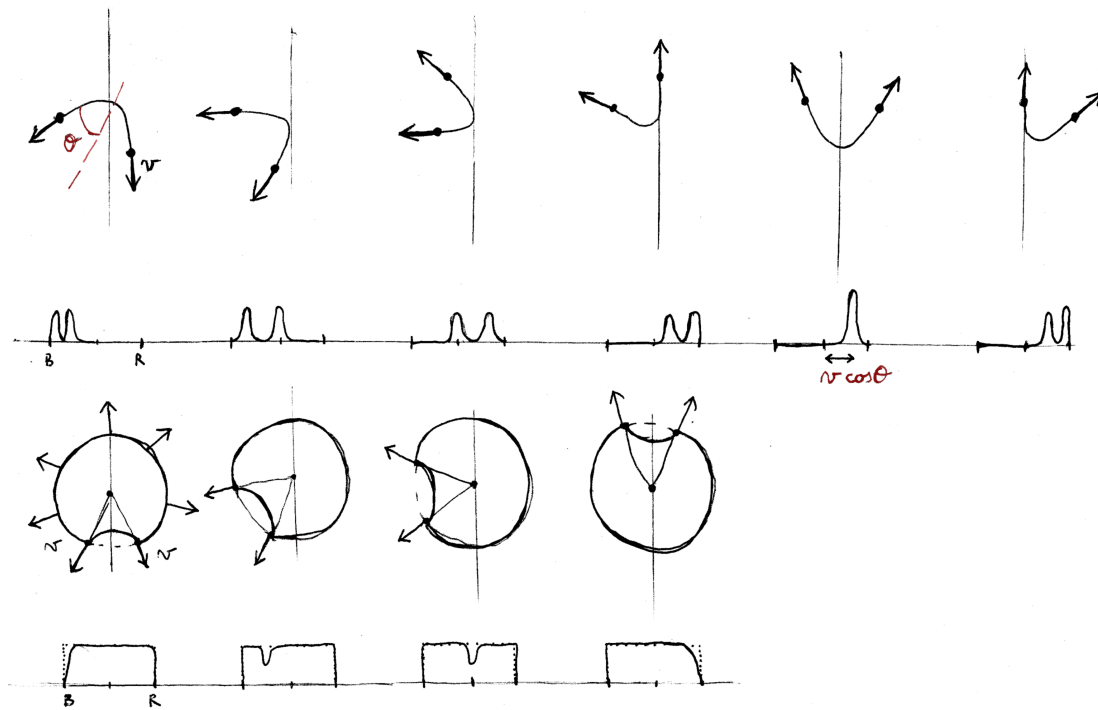


Figure 1.11 – Explication qualitative du profil d'excès d'émission dû à un cône de choc (haut) et du déficit d'émission dû à un trou circulaire dans la coquille d'émission en expansion correspondant à l'intersection entre cette dernière et le cône de choc (bas). Dans les deux cas, on considère pour simplifier une inclinaison du système de  $90^\circ$ . B et R désignent les côtés bleu et rouge du profil.

limites  $\theta_1$  et  $\theta_2$ , et étant maximal au milieu de ces deux bornes<sup>4</sup>. Le profil synthétique ( $I(v)$ ) est alors obtenu par l'intégration pour chaque vitesse  $v$  de l'émissivité par unité de masse  $I(\theta', v)$  sur le volume du cône de choc (Fig. 1.12(b)), soit :

$$I(v)dv = dv \times \int_{\theta_1}^{\theta_2} 2\pi \sin(\theta') d\theta' \rho(\theta') . \quad (1.39)$$

Des exemples de profils synthétiques pour différentes valeurs de paramètres sont montrés à la Fig. 1.13. Ces profils peuvent être ajustés aux observations pour déterminer les paramètres du système.

#### 1.2.4.2 Détermination des paramètres du système à partir des mesures de $\bar{v}$ et $v^*$

De manière alternative, la mesure de la position des maximums du profil de la raie au cours du temps ( $v_{\text{red}}$  et  $v_{\text{blue}}$ ) peut également servir à la détermination des paramètres du système. En effet, si l'on définit  $\bar{v}$  et  $v^*$  comme suit :

$$\bar{v} = \frac{v_{\text{red}} + v_{\text{blue}}}{2} \quad (1.40)$$

$$v^* = \frac{v_{\text{red}} - v_{\text{blue}}}{2} , \quad (1.41)$$

le modèle géométrique de Lührs donne la loi d'évolution suivante en fonction de la phase corrigée  $\phi^*$  :

$$\bar{v} = -v_{\text{strm}} \cos(\theta) \sin(i) \cos(\phi^*) \quad (1.42)$$

$$v^* = v_{\text{strm}} \sin(\theta) \sqrt{1 - \sin^2(i) \cos^2(\phi^*)} , \quad (1.43)$$

où  $\phi = \phi^* + \delta\phi$  est pris égal à 0 à la conjonction inférieure de l'étoile O (= la projection de l'étoile O sur la ligne de visée est devant la projection de la Wolf-Rayet). Le résultat de cette loi est tracé à la Fig. 1.14.

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<sup>4</sup>On peut montrer que dans le cas adiabatique, on a l'égalité  $\delta\theta = \theta_1 - \theta_2 = \theta$ .

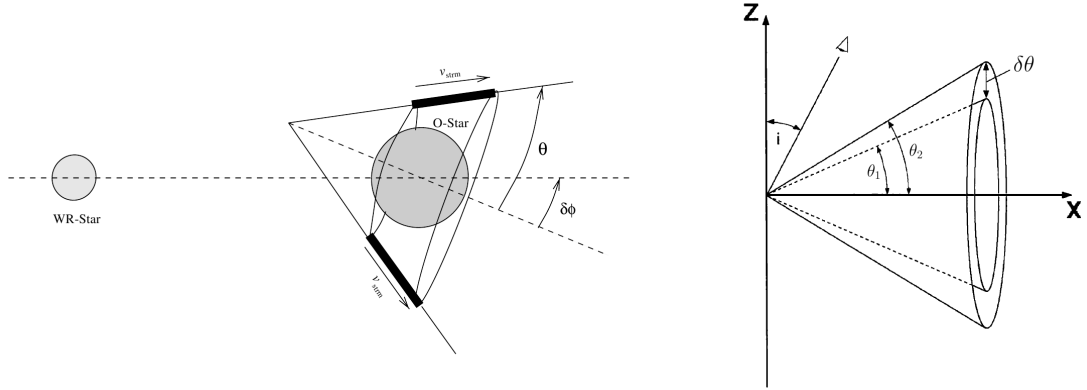


Figure 1.12 – Paramétrisation du cône de choc pour le modèle de Lührs.  $v_{\text{strm}}$  est la vitesse de l'écoulement, supposée constante, le long de la surface de contact. L'effet des forces de Coriolis est quantifié par un déphasage  $\delta\phi$  entre l'axe du cône et l'axe reliant les deux étoiles. La région d'émission de la raie est supposée être une bande autour de la surface de contact. L'épaisseur de la région de choc est notée  $\delta\theta$ . (Tiré de Bartzakos et al. 2001, Lührs 1997)

### 1.2.4.3 Détermination des paramètres du système à partir des mesures de $FW_{\text{ex}}$ et $RV_{\text{ex}}$

Lorsque les mesures de  $\bar{v}$  et  $v^*$  ne sont pas accessibles dans les données, il est encore possible de déterminer les paramètres du système par la mesure de la largeur à mi-hauteur du profil total de l'excès d'émission et de la vitesse radiale globale de ce même profil. Les lois d'évolution suivantes ont été calculées par Hill et al. (2000) à partir du modèle de Lührs :

$$FW_{\text{ex}} = C_1 + 2v_{\text{strm}} \sin(\theta) \sqrt{1 - \sin^2(i) \cos^2(\phi - \delta\phi)} \quad (1.44)$$

$$RV_{\text{ex}} = C_2 + v_{\text{strm}} \cos(\theta) \sin(i) \cos(\phi - \delta\phi) , \quad (1.45)$$

où  $C_1$  et  $C_2$  sont de simples constantes.

### 1.2.5 Production de poussière

Un excès infrarouge apparaît souvent dans les étoiles Wolf-Rayet de type WC tardives (Williams et al., 1987a). Cet excès est associé à la production de poussières

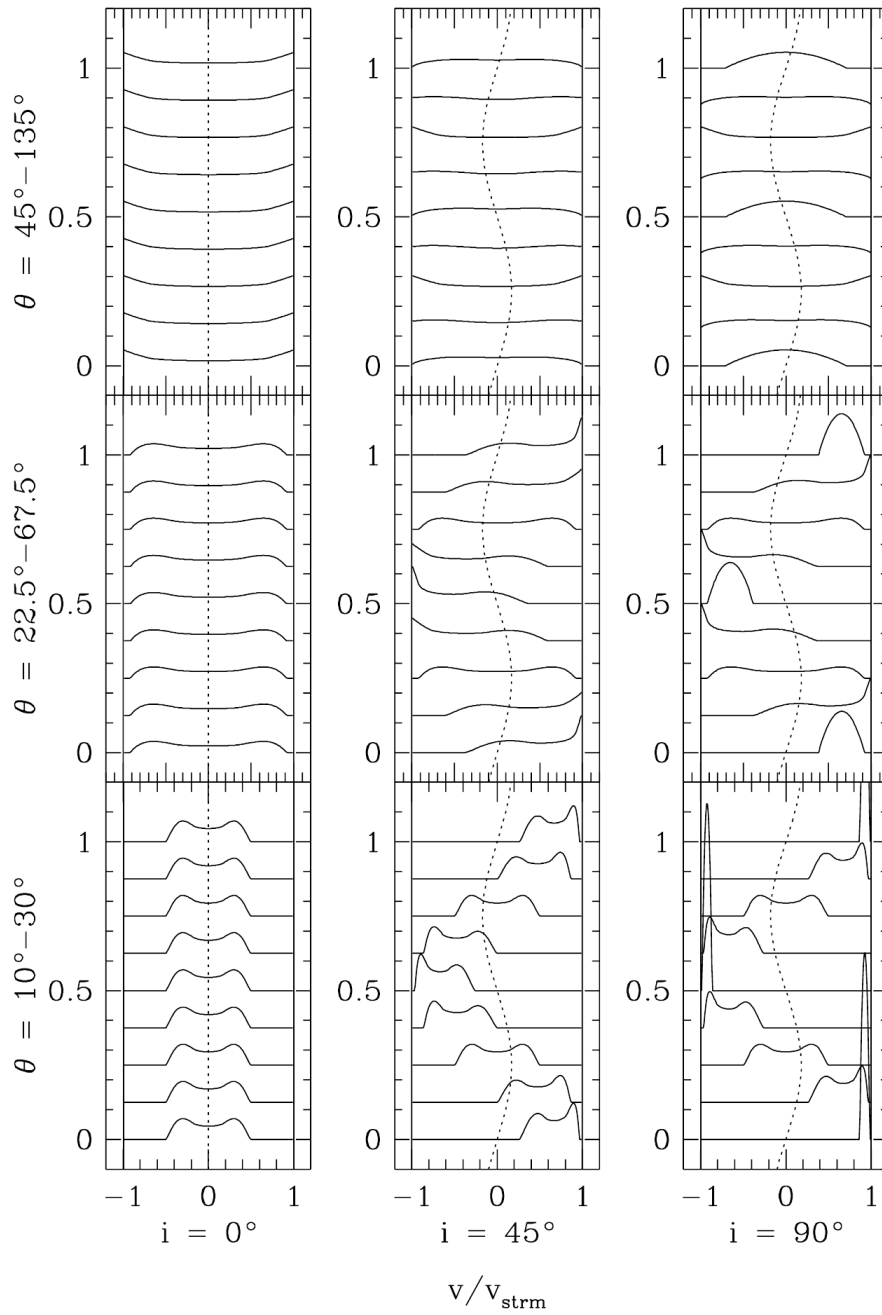


Figure 1.13 – Profils d’excès d’émission calculés par le modèle de Lührs pour différentes valeurs de l’inclinaison  $i$  du système et de la forme du cône de choc  $[\theta_1, \theta_2]$ . Pour  $i=0$ , le profil ne varie pas et s’élargit avec l’élargissement du cône. Pour  $i=90$ , on retrouve le résultat qualitatif décrit à la Fig. 1.11. (Tiré de Bartzakos et al. 2001)

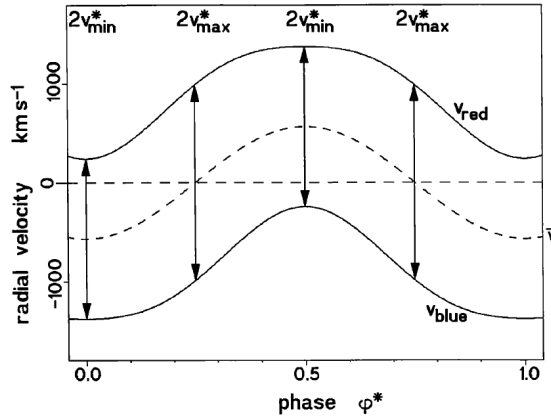


Figure 1.14 – Variation en fonction de la phase corrigée  $\phi^*$  de la position  $v_{red}$  et  $v_{blue}$  des maximums du profil d’excès d’émission, ainsi que de la position du milieu du profil  $\bar{v}$  et de la distance entre eux  $v^*$ , selon le modèle de Lührs. (Tiré de Lührs 1997)

de carbone amorphe, qui absorbent le rayonnement UV et optique de l’étoile et le ré-émet sous la forme d’un corps noir dans l’infrarouge aux longueurs d’onde micrométriques. Un exemple d’excès infrarouge est donné à la Fig. 1.15 pour WR 70, WR 76 et WR 80 (issu de Williams et al., 1987a). Cet excès permet, à l’aide d’un modèle d’émission, d’estimer le taux de formation de ces poussières (Williams et al., 1987a). Ce taux peut prendre des valeurs allant de  $\simeq 10^{-9} M_{\odot} \text{ an}^{-1}$  à  $\simeq 10^{-6} M_{\odot} \text{ an}^{-1}$ . Dans certains systèmes, la poussière est même résolue grâce à l’imagerie infrarouge, formant des spirales ou des portions de spirale de poussières autour de l’étoile et suggérant ainsi la présence d’un système double (Monnier et al., 1999, Tuthill et al., 1999).

Certaines étoiles forment de la poussière de manière périodique et sont des systèmes doubles avérés, comme c’est le cas pour WR 140 avec une période de presque 8 ans (Marchenko et al., 2003). Dans de tels systèmes, la formation de poussière semble se produire au moment du passage au périastre (Williams et al., 1990) (Fig. 1.16). Dans certains systèmes, par contre, la binarité n’a pu être prouvée de manière convaincante. Un résumé des types d’étoiles produisant de la poussière est présenté au Tab. 1.I.

Le phénomène de formation de poussière dans un environnement aussi hostile

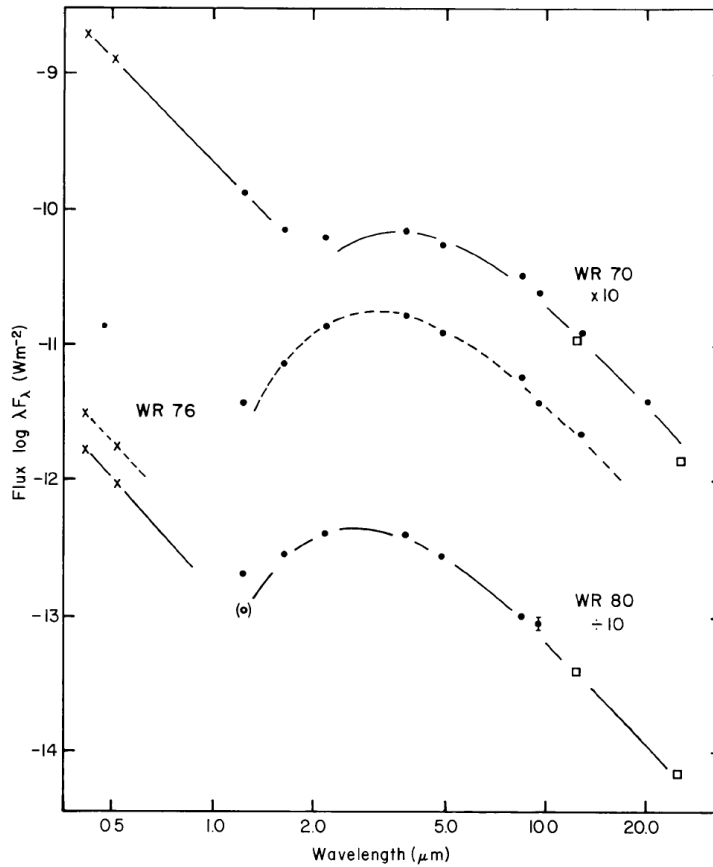


Figure 1.15 – Exemples d’excès d’émission infrarouge observés pour trois étoiles WC9 : WR 70, 76 et 80. Sont tracées également les courbes de corps noirs s’ajustant le mieux aux profils. (Tiré de Williams et al. 1987a)

Subtype	Persistent dust formation (WCd)	Variable dust formation (WCvd)	Periodic dust formation (WCpd)	Episodic dust formation (WCed)
WC4			WR 19	
WC7			WR 137,140	WR 125
WC8	WR 53,113	WR 98a		
WC9	WR 59, 65, 69, 73, 76, 80, 95, 96, 103, 104, 106, 112, 117, 118, 119, 121	WR 70		

Tableau 1.I – Distribution des types d’étoiles formant de la poussière (“dustars”) en fonction de leurs types spectraux. (Tiré de van der Hucht 2001)



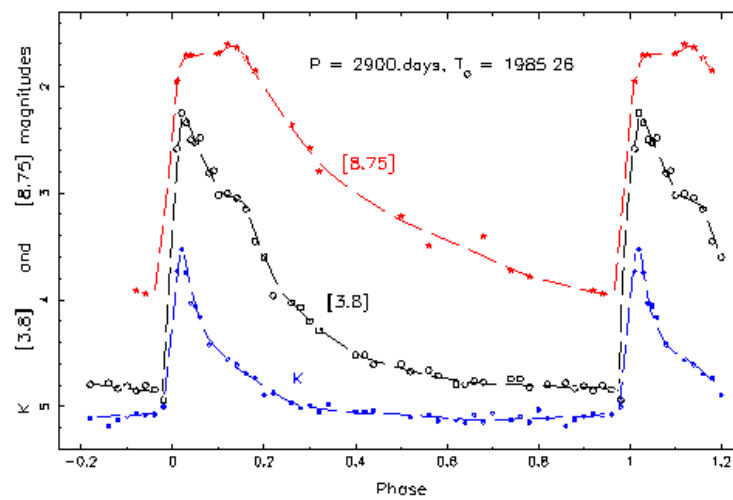


Figure 1.16 – Courbes de lumière infrarouge du système binaire WR 140 (WC7+O) dans trois bandes :  $K$ ,  $3.8 \mu\text{m}$  et  $8.75 \mu\text{m}$ . L'augmentation de la luminosité infrarouge coïncide avec le passage au périastre du système. La baisse progressive de la luminosité après le maximum correspond au refroidissement des grains de poussière, les plus petits grains, dominant l'émission dans la bande  $K$  se refroidissant plus vite que les gros grains, dominant l'émission dans la bande  $3.8 \mu\text{m}$  puis  $8.75 \mu\text{m}$ . (Tiré de Williams et al. 1990)

(intense rayonnement UV, rayons X) reste encore relativement incompris. Cependant, il existe des indices suggérant l'importance cruciale de la binarité et de la collision des vents, capable de produire les densités nécessaires à la condensation des poussières. La physique de la formation de poussière dans de tels systèmes devra nécessairement impliquer un refroidissement efficace du gaz (choc radiatif) ainsi qu'un mécanisme d'écrantage (par exemple par des clumps) du rayonnement UV et X provenant des étoiles.

### 1.3 Conclusion du premier chapitre

Deux étoiles massives évoluant dans un système binaire vont voir leurs vents interagir le long d'une région de collision dont la forme, approximativement celle d'un cône d'ouverture  $2\theta$ , dépend de la force relative des deux vents (le paramètre  $\eta$ ). Cette région de choc est composée d'une surface de contact et de deux surfaces de chocs situées de part et d'autre de cette première. La thermalisation de l'énergie cinétique des vents (de l'ordre de 1 keV) est capable de chauffer le gaz choqué à des températures de l'ordre de  $10^7$  K, produisant ainsi un fort rayonnement X. L'intensité de ce rayonnement ainsi que la géométrie même des surfaces de choc et la présence d'instabilités hydrodynamiques va dépendre crucialement de l'importance du refroidissement radiatif au sein de la région (le paramètre  $\chi$ ). On a montré que ce refroidissement, dominé par l'émission libre-libre, était proportionnel à  $\rho^2\sqrt{T}$ , ce qui nous a amené à la conclusion que les binaires les plus détachées (dont la période est supérieure à quelques jours) présentaient des chocs adiabatiques, tandis que les binaires serrées présentaient plutôt des chocs radiatifs. La présence du cône de choc peut également être observée par son effet sur les raies d'émission visibles et UV. Il génère en effet un excès d'émission à double pic dont la forme et la position varient avec la phase orbitale. Un modèle géométrique simple permet de reproduire ces profils, et d'avoir accès à la géométrie du cône de choc ainsi qu'à l'inclinaison du système (ce qui est crucial pour la détermination des masses stellaires). La collision des vents est aussi très probablement responsable de la formation de poussière de carbone amorphe dans les binaires contenant des

WC (bien que des mécanismes pour les étoiles simples aient été proposés ; Zubko 1998). Cette poussière est observée par l’excès infrarouge sous forme de corps noir qu’elle occasionne dans le spectre. La physique de la formation de ces poussières dans un environnement aussi hostile est cependant encore mal comprise, mais doit certainement impliquer un refroidissement efficace du gaz ainsi qu’un mécanisme d’écrantage de la radiation UV et X.

L’étude des systèmes binaires à collision de vents est cruciale pour la compréhension des étoiles massives car elle donne accès à une mesure directe des masses des étoiles ainsi qu’à une mesure des taux de perte de masse indépendante du clumping, deux paramètres décisifs pour l’élaboration de modèles d’évolution stellaire.

#### 1.4 Description du projet de doctorat

Le but de mon projet de doctorat est d’étendre l’étude des systèmes binaires à collision de vents aux systèmes de l’hémisphère sud. En effet, la plupart des systèmes de l’hémisphère nord ont déjà fait l’objet d’études poussées, mais un important travail reste encore à faire dans l’hémisphère sud, où l’on trouve la majorité des étoiles Wolf-Rayet, et la quasi totalité des étoiles WC riches en carbone. Ce doctorat est constitué de 3 articles publiés ou soumis dans des revues à arbitre et s’oriente selon deux axes privilégiés, pouvant parfois s’entremêler.

Le premier axe pourrait s’intituler : “les systèmes binaires comme outils de mesure des paramètres fondamentaux des étoiles”. Quelles sont les méthodes à utiliser pour avoir accès aux paramètres stellaires dans le cadre de la spectroscopie optique, et quelles en sont les limites ? Cet axe concerne les deux premiers articles (chapitres 2 et 3).

Le deuxième axe concerne la formation de poussière dans les WC tardives et le lien entre ce phénomène et la binarité. Dans quelle mesure peut-on relier la formation de poussière à la binarité ? Existe-t-il d’autres mécanismes, propres aux WC tardives, à l’origine de la formation de ces poussières ? Par une recherche systématique de variabilité photométrique dans ces étoiles, nous apportons des

éléments de réponse à cette question. C'est le sujet du dernier article (chapitre 4) de cette thèse.

Le premier projet constituant mon doctorat est le suivi spectroscopique de WR 140 dans le domaine visible lors de son dernier passage au périastre, qui a eu lieu en janvier 2009. Cet objet est considéré comme l'archétype des systèmes à collision de vents. La campagne d'observation a consisté en une collaboration entre astronomes professionnels et astronomes amateurs et a eu lieu sur une demi-douzaine de sites incluant l'observatoire du Teide (Îles Canaries), l'observatoire de Haute-Provence (France), le Dominion Astrophysical Observatory (Colombie Britannique) et l'observatoire du Mont-Mégantic (Québec). Cette campagne a permis d'obtenir une résolution temporelle jamais atteinte sur une telle durée pour ce système, et nous a donné accès, notamment grâce à l'étude des profils d'excès d'émission, à une estimation plus précise des paramètres de l'orbite, ainsi qu'à la géométrie et à la nature des chocs. Cette étude a fait l'objet d'un article accepté pour publication dans MNRAS en mai 2011.

Un deuxième projet a consisté en un suivi spectroscopique de 5 systèmes WR+O de l'hémisphère sud qui n'ont jamais été étudiés en terme de collision de vents : WR 12 (WN8h), 21 (WN5o+O7V), 30 (WC6+O7.5V), 31 (WN4o+O8) et 47 (WN6o+O5). Les observations ont été faites au Complexe Astronomique El Leoncito (CASLEO) à l'aide du télescope 0.6 m Helen Sawyer Hogg pendant un mois complet (avril 2009). Nous avons obtenu, de manière similaire au travail sur WR 140, une mise à jour des paramètres orbitaux ainsi qu'une estimation de la géométrie du cône de choc et de l'inclinaison orbitale. Une 6e étoile, WR 69, dont certains indices suggéraient qu'elle était un système binaire à collision de vent de courte période ( $P=2.293\pm 0.005$  j, Marchenko et al. 1998b), a également été observée. Il s'est avéré que la période orbitale du système est vraisemblablement beaucoup plus longue (supérieure au mois), et que la période publiée pour celui-ci n'est pas d'origine orbitale. Nous avons cependant identifié dans nos spectres une période probable de  $2.15\pm 0.04$  jours, très proche de la période de Marchenko. L'article issu de ce projet a été soumis à MNRAS en juin 2011.

Enfin, un troisième projet a consisté en un suivi en photométrie de 19 étoiles WC9 et 1 WC8 de l'hémisphère sud pendant 1 mois (juillet 2007) au télescope de 1 m Swope à Las Campanas, Chili. Le but de ce projet est la recherche d'éventuelles pulsations de grande amplitude dans ces étoiles, comme elles sont observées systématiquement dans leurs cousines riches en azote, les WN8 (Marchenko et al., 1998a, Veen et al., 1999). Ces pulsations pourraient alors être mises en lien, d'une part avec le poussage du vent stellaire, en complément de la pression de radiation, comme le suggèrent certains auteurs (Maeder, 1985). Et d'autre part, du fait des chocs transférés dans le vent (Grott et al., 2005, Townsend, 2007), avec la formation de poussière, ce qui constituerait un mécanisme alternatif à la collision de vent. Nous avons observé un niveau de variabilité étonnamment faible pour notre échantillon, ce qui montre que les pulsations dans les WC8/9 ne sont pas aussi courantes et aussi violentes que dans les WN8. Pour l'étoile la plus variable de notre échantillon, WR 76, la variabilité est clairement due à une extinction variable due à la poussière. Nous avons calculé une taille moyenne des grains de poussière de  $0.126 \pm 0.003 \mu\text{m}$ . L'article issu de ce projet a été publié dans MNRAS en janvier 2009.

## 1.5 Contribution aux articles

### 1.5.1 Spectroscopy of the archetype colliding-wind binary WR 140 during the 2009 January periastron passage

Malgré le grand nombre de co-auteurs, ma contribution à cet article est très largement dominante. La contribution des co-auteurs se résume essentiellement à la participation aux campagnes d'observation et à une collaboration lors de la phase de rédaction de l'article (corrections et quelques suggestions d'ajouts ou modifications). Cependant, ce projet d'échelle importante, n'aurait pu avoir lieu sans cette collaboration.

Ma contribution à l'article est détaillée ci dessous :

Participation à la campagne d'observation de Ténérife (Télescope MONS 50 cm)

pendant 1 mois complet.

Réduction et extraction des spectres de Mons, OMM, Three Hills O., Berthold Stober et Lothar Shanne.

Élaboration et mise en oeuvre d'une technique pour corriger le résidu de blaze variable (oscillation résiduelle instrumentale) observé dans les spectres échelle de l'OHP.

Élaboration et mise en oeuvre d'une technique originale d'élimination des raies d'absorption telluriques dans le cas où celles-ci sont résolues.

Normalisation homogène des spectres de toutes les sources d'observation (professionnels et amateurs).

Extraction des vitesses radiales et déduction des paramètres orbitaux.

Extraction et analyse de l'excès d'émission.

Écriture et application des routines de fit de l'orbite et fit de l'excès d'émission.

Élaboration d'une version modifiée du modèle de Lührs (1997).

Écriture et application d'une routine de shift-and-add (Demers et al., 2002).

Classification des spectres.

Élaboration des figures.

Cet article a été déjà cité plusieurs fois, vu son importance cruciale pour toute autre étude sur cet objet de base.

### **1.5.2 Colliding winds in five WR + O systems of the southern hemisphere**

Observations au télescope HSH du Complexe Astronomique El Leoncito (CAS-LEO), Argentine, pendant une durée de 1 mois.

Réduction et extraction des spectres.

Analyse (impliquant une partie du travail effectué dans le cadre de l'article précédent).

Élaboration et application d'une routine permettant le calcul des erreurs sur les périodes trouvées par Lomb-Scargle avec la méthode Monte-Carlo.

Élaboration des figures.

### **1.5.3 Photometric variability of WC9 stars**

Observations au télescope Swope de Las Campanas, Chili pendant une durée de 1 mois. Réduction des images.

Extraction des courbes de lumière à l'aide d'ISIS et IDL.

Adaptation et utilisation d'un code de simulation pour le calcul d'extinction par le modèle de Mie.

Élaboration des figures.





## CHAPTER 2

### SPECTROSCOPY OF THE ARCHETYPE COLLIDING-WIND BINARY WR 140 DURING THE 2009 JANUARY PERIASTRON PASSAGE

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## ABSTRACT

We present the results from the spectroscopic monitoring of WR 140 (WC7pd + O5.5fc) during its latest periastron passage in January 2009. The observational campaign consisted of a constructive collaboration between amateur and professional astronomers and took place at half a dozen locations, including Teide Observatory, Observatoire de Haute Provence, Dominion Astrophysical Observatory and Observatoire du Mont Mégantic. WR 140 is known as the archetype of colliding-wind binaries and has a relatively long period ( $\simeq 8$  years) and high eccentricity ( $\simeq 0.9$ ). We provide updated values for the orbital parameters, new estimates for the WR and O-star masses and new constraints on the mass-loss rates and colliding-wind geometry.

**Keywords :** (stars:) binaries: general, stars: fundamental parameters (classification, colors, luminosities, masses, radii, temperatures, etc.), stars: Wolf-Rayet , stars: winds, outflows

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### 2.1 Introduction

Binary systems provide the least model-dependent way to secure measurements of stellar masses. Understanding binary systems is therefore of paramount importance for many problems in stellar astrophysics. In the case of massive stars and most particularly WR+O systems, the strong, fast winds of the two stars will interact with each other along a shock surface, giving rise to an excess emission in a range of wavelengths going from radio to X-rays, possibly even  $\gamma$ -rays (Moffat and St-Louis, 2002). Understanding the underlying physics of this phenomenon gives us access to additional parameters of the system, like the wind momentum

ratio and the orbital inclination.

Among these colliding-wind systems, WR 140 is considered as the archetype. WR 140 is a WC7pd + O4-5 system with a large eccentricity ( $e = 0.88$ ) and long period ( $P = 7.94$  years) (Marchenko et al. 2003, hereafter M03). It is also the brightest Wolf-Rayet star in the northern hemisphere (Cygnus) with a visual magnitude of 6.9. It was first suggested to be a spectroscopic binary in 1924 (Plaskett, 1924) but searches for radial velocities were unsuccessful until a period was found in infrared photometry (Williams et al., 1987b) and a first spectroscopic orbit in 1987 (Moffat et al., 1987). A spectroscopic campaign around periastron in 2001 revealed that the spectral changes due to the wind-wind collision were much faster than expected and therefore, not sampled properly. New spectroscopic observations were thus requested in 2009 around periastron.

In this article, we present our analysis of an international optical spectrometric campaign which involved both professional and amateur astronomers at half a dozen locations. The campaign focuses on a short but crucial fraction of the orbit (four months) around the January 2009 periastron passage, during which the radial velocities vary most significantly.

In the next section, we will describe the multisite observational campaign and present the different instruments used. We will then devote a section to the data reduction procedure. We will also present the orbital solution deduced from the radial velocity measurements of the two components and look at what can be said about the spectral classification of the stars, in particular the luminosity class of the O-star. Finally, we will discuss the analysis of the excess emission induced by the colliding-wind region (CWR).

## 2.2 Observations

### 2.2.1 Pro-Am campaign

One of the highlights of this observational campaign is that it consisted of a constructive collaboration between amateur and professional astronomers. In-

deed, since the appearance of the first commercial spectrographs, like those of the French company Shelyak, a small, but growing, community of amateurs is willing to involve themselves in scientific collaborations with professionals. The Bess project of Be star spectroscopy under the supervision of Coralie Neiner and François Cochard, is an example of such a collaboration (Neiner et al., 2007). See websites<sup>1</sup> of VDS (Germany) and ARAS (France) for further information about amateur groups involved in Pro-Am collaboration. With progressive improvements, these projects are offering a serious alternative to traditional observational modes, especially for monitoring over long time scales.

### 2.2.2 Data

Among the data collected through Pro-Am collaboration, we first have the Mons project. Under the leadership of one of us (TE), a LHIRES III spectrograph was installed on a telescope previously owned by the University of Mons in Belgium, and nowadays used mainly for teaching purposes. This is a 50 cm telescope located at the Teide Observatory of the Instituto de Astrofísica de Canarias (IAC) in Tenerife. Four months of data were acquired with this instrumentation. Other amateurs contributed using their own personal instruments in Portugal (José Ribeiro), Germany (Berthold Stöber) and England (Robin Leadbeater at Three Hill Observatory<sup>2</sup>).

The rest of the data was obtained with the échelle spectrograph SOPHIE at OHP (Observatoire de haute Provence) and with more traditional single-grating long-slit spectrographs at DAO (Dominion Astrophysical Observatory) and at OMM (Observatoire du Mont-Mégantic). We are primarily interested in the yellow spectral region ( $\simeq 5500\text{-}6000 \text{ \AA}$ ) where the main CWR-sensitive optical emission lines of C III 5696  $\text{\AA}$  and He I 5876  $\text{\AA}$  lie. A list of the data sets of the campaign is presented Table 2.I.

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<sup>1</sup>VDS : <http://spektroskopie.fg-vds.de/>  
 ARAS : <http://www.astrosurf.com/aras/>

<sup>2</sup><http://www.threehillsobservatory.co.uk/>

Observatory	Dates	$\lambda\lambda$ (Å)	Recip. disp. (Å/pixel)	Res. (pixels)	No. spectra	Weight
Teide Mons 50 cm	1.12.08 - 23.03.09	5530-6000	0.35	5-6	34	0.12
OHP 193 cm	12.12.08 - 23.3.09	3900-6800	0.01	15	63	0.49
DAO 120 cm	22.4.08 - 9.1.09	5350-5900	0.37	4.5	13	0.16
OMM 160 cm	5.7.09 - 8.8.09	4500-6000	0.63	2.6	18	0.09
Three Hills O.* 20 cm (UK)	10.12.07 - 20.3.09	5600-6000	0.68	3.5	38	0.08
Berthold Stober* 36 cm (Germany)	26.8.08 - 29.2.09	5500-6100	0.53	2-3	12	0.06

Table 2.I: List of the different data sets in the 2009 campaign. Names marked with an asterisk are amateur astronomers/organisations equipped with their personal instrumentation. We also show the relative weight (= normalized  $1/\sigma^2$ , where  $\sigma$  is the dispersion around the fitted RV curve) used in the orbit fit. For the OMM data, a technical problem for the spectrograph focus gave lower quality spectra than expected, which explains the low weight of this data set.

## 2.3 Data reduction

### 2.3.1 Pre-reduction

Each set of data was first pre-reduced the standard way using the CCDPROC package (IRAF<sup>3</sup>) and following the procedure described in P. Massey’s “User’s Guide to Reducing Slit Spectra with IRAF” (Massey et al., 1992) : bias subtraction, normalized flat division and cosmic ray removal using the routine COSMICRAYS . For the amateur spectra, the neon comparison lamps present only a few lines in the main yellow spectral range. The calibration exposures were thus overexposed to access the more numerous very faint neon lines, leading to a saturation of the fewer stronger lines. About twelve lines were used to compute the dispersion solution. In spite of this, the wavelength calibration was sufficiently

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<sup>3</sup>IRAF is the Image Reduction and Analysis Facility, a general purpose software system for the reduction and analysis of astronomical data. IRAF is written and supported by the IRAF programming group at the National Optical Astronomy Observatories (NOAO) in Tucson, Arizona. NOAO is operated by the Association of Universities for Research in Astronomy (AURA), Inc. under cooperative agreement with the National Science Foundation (<http://iraf.noao.edu/>).

accurate for our purpose. Where it was present, we used the interstellar sodium doublet at  $\lambda\lambda$  5890,5895 Å as a fine-tuning of the dispersion solution. This was done by applying a constant shift in wavelength to this solution in order to get the doublet at the right position. In the vast majority of cases, this shift is less than two pixels (it’s always less than one pixel for the OHP data).

The échelle spectra were reduced using the automatic OHP reduction package adapted from the HARPS software of the Geneva Observatory. It includes a procedure to reconnect the orders of the echelle spectrum, thus leading to a single 1d spectrum. However, this reconnection procedure is not perfect and a flux residual of variable amplitude centred around 5% persists in the data. This defect was for the most part removed by using the fact that its shape was nearly constant. This shape was determined by dividing the spectra by each other, which nearly eliminated the stellar line-contribution, and by fitting an appropriate analytical model using one polynomial per order of the échelle spectrograph (39 in total), these polynomials being then smoothly reconnected with each other. We finally modified all the spectra with this correction so that they best matched a non-affected single-grating spectrum (we chose one from the OMM for its wide wavelength range).

### 2.3.2 Telluric line removal

As no “telluric” stars were observed in the OHP campaign, an alternative solution was used to remove the telluric absorption lines. These lines were particularly strong because the star was in a very low position on the sky at the time of the campaign (air masses of about 2). Nevertheless, SOPHIE’s resolving power of 40000 at 5500 Å (HE mode) can largely resolve these lines, whose width is about 1 Å. So, after removing the non-telluric structures with widths larger than a few Angstroms, the lines were fitted with a sum of Gaussians at positions indicated in the IRAF atlas of telluric lines. The results of the fit were then used to divide into each of the spectra to eliminate the telluric contributions.

## 2.4 Radial velocities

### 2.4.1 O-star

The O-star radial velocities (RVs) were obtained by measuring the centroid of each relatively narrow photospheric absorption line (allowing for a sloping pseudo continuum if superposed on WR lines), and taking the mean of the deduced velocities for each spectrum. We took into account a possible shift between the RVs determined from different O-star lines. This was done by choosing one RV set as a reference and by plotting the other sets versus this reference. By fitting a linear curve of slope 1 (i.e.  $y=x+b$ ) into the points, we could then deduce the velocity shift ( $=b$ ) to be subtracted from the considered set. Typical velocity shifts have values from 5 to 10 km s<sup>-1</sup>. Lines used were: H I 3970.09, 4101.74, 4340.47, 4861.33; He I 4026.19, 4471.48, 5875.62; He II 4199.87, 4541.63, 5411.56.

### 2.4.2 Wolf-Rayet star

The WR-star RVs were estimated by cross correlation with a reference spectrum, chosen to be unaffected by the wind-wind collision effects. This reference spectrum was constructed by combining all the OHP spectra outside the phase interval 0.98-1.02, where obvious CWR effects occur. The same reference spectrum was used for all the data sets. We used the whole spectral domain of each data source in the correlation and checked that the excess emission did not critically bias the measurements by testing by eye the shape similarity between the WR and the O RV curves.

### 2.4.3 Orbit fit

The measured radial velocities and errors of the 2009 campaign are presented in annex II. The typical errors ( $\sigma$ ) for each observatory are deduced from the internal dispersion of the differences (RV-Fit) between the measurements and the orbital solution (O and WR taken together). The fitted parameters are the period  $P$ , the eccentricity  $e$ , the time of periastron passage  $T_0$ , the longitude of periastron



passage in the plane of the orbit measured from the ascending node  $\omega$ , and the semi-amplitudes  $K(\text{WR})$  and  $K(\text{O})$  of the velocity curves. The radial velocities were fitted using the IDL<sup>4</sup> routine MPFIT.PRO (Markwardt, 2009), which uses the Levenberg-Marquardt technique to solve the least-squares problem. Errors on the parameters are computed by MPFIT using the covariance matrix method applied on the RV errors given in input. We fitted for each set of data an independent  $\gamma$  parameter (=systemic velocity) to best reconnect the different data sets with each other. Observed differences in  $\gamma$  between the sets are of order 5 to 10 km s<sup>-1</sup> and are most probably instrumental. A weight was attributed to each set based on the external rms dispersion of the measurements minus the model. Data from the previous observational campaigns (M03) were also included in the fit; only the weights were recalculated with our method in order to insure homogeneity. The resulting fit is given in Fig. 2.1 and the fitted orbit parameters in Table 2.II.

The differences between the new results and M03 are not formally significant at the 3 sigma level but we improved the precision on the orbital parameters. We also note that our new period ( $2896.5 \pm 0.7$  d instead of  $2899.0 \pm 1.3$  d) agrees better with recent X-ray data repeatability (Corcoran, private communication). Fits were also made for the WR- and the O-star separately which give respectively :  $P=2896.5$ ,  $e=0.896$ ,  $K_0=46155.8$ ,  $K[\text{WR}]=-75.7$ ,  $\omega=44.5$  and  $P=2896.5$ ,  $e=0.897$ ,  $T_0=46159.1$ ,  $K[\text{O}]=30.6$ ,  $\omega=45.2$ . This is consistent within the errors with the O/WR simultaneous fit.

## 2.5 Spectral Classification

The WR component has been previously confirmed as a WC7pd star, where  $p$  refers to its peculiarly broad emission lines (more like WC5) for its WC7 subtype and  $d$  refers to its dust-forming events. We see no revision to this classification but we present precise measurements of equivalent widths and FWHM of different WR lines (Table 2.III) and a comparison of these measurements with the classification criteria of Crowther et al. (1998) (Table 2.IV). On the other hand, the spectral

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<sup>4</sup>Interactive Data Language, ITT Visual Information Solutions (ITT VIS)

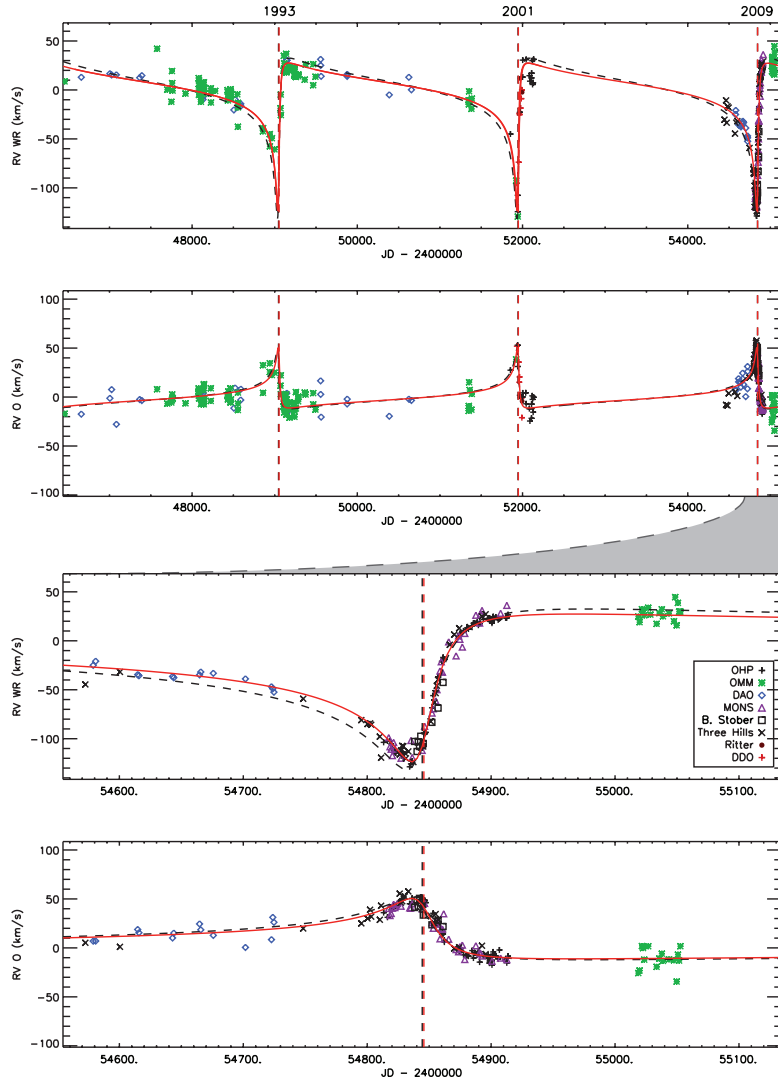


Figure 2.1: (top two panels) Measured radial velocities of the WR-star and of the O-star together with the fit for the orbital solution (full line) whose best-fit parameters are given in Table 2.II. We included data from the last periastron campaign in 2001 (M03) taken at David Dunlap Observatory (DDO), OHP, Ritter Observatory, DAO and OMM. The black dashed line is the orbital solution from M03. The dashed vertical lines show the position of the new periastron passage. Both “old” and “new” periastron passages are shown. (bottom two panels) Same plots but zoomed on the 2009 campaign (the gray-filled curve between panel two and three illustrates the X axis expansion).

Parameter	This paper+M03	This paper	M03
$K$ [WR] ( $\text{km s}^{-1}$ )	$-75.5 \pm 0.6$	$-75.3 \pm 0.7$	$-82.0 \pm 2.3$
$K$ [O] ( $\text{km s}^{-1}$ )	$30.9 \pm 0.5$	$30.7 \pm 0.6$	$30.5 \pm 1.9$
$a \sin(i)$ [WR] (A.U.)	$8.92 \pm 0.15$	$8.98 \pm 0.17$	$10.29 \pm 0.47$
$a \sin(i)$ [O] (A.U.)	$3.646 \pm 0.085$	$3.665 \pm 0.093$	$3.81 \pm 0.27$
$P$ (days)	$2896.5 \pm 0.7$	2896.5	$2899.0 \pm 1.3$
$e$	$0.8962 \pm 0.0014$	$0.894 \pm 0.002$	$0.881 \pm 0.005$
$T_0$ (HJD 2400000+)	$46156.2 \pm 0.3$	$46157.0 \pm 0.3$	$46147.4 \pm 3.7$
$\omega$ (deg.)	$44.6 \pm 0.6$	$47.5 \pm 0.9$	$46.7 \pm 1.6$

Table 2.II: New orbit parameters for WR 140 compared with the values from M03.

type of the O-star component is less clearly identified on our combined mean spectrum, and most especially its luminosity class. The determination of this luminosity class is of prime importance for the determination of the distance and to deduce information about the O-star wind. The quality of the OHP spectra provides an opportunity to extract this luminosity class.

Given the dominating strength of the WR spectrum, it is useful to apply a spectral decomposition technique to isolate the O-star spectrum. We therefore used the shift-and-add technique of Demers et al. (2002), with our new orbital solution. In this procedure, we first co-add all the spectra in the WR frame of reference, by correcting each spectrum from its orbital velocity (taken on the fitted curve and not the individual measurements), which “dilutes” the O-star absorption

Line or Blend	$\log(ew_\lambda)$	FWHM ( $\text{\AA}$ )
C III/ IV+He II 4650,4658,4686	$2.348 \pm 0.005$	$67 \pm 2$
O III-V 5590	$1.05 \pm 0.04$	$75 \pm 2$
C III 5696	$1.87 \pm 0.01$	$94 \pm 2$
C IV+He I 5808,5876	$2.413 \pm 0.004$	$71 \pm 2$
C IV 5808	$2.215 \pm 0.006$	$51 \pm 2$

Table 2.III: Equivalent widths and FWHM measured on different WR emission lines. The typical FWHM of C III/ IV+He II for a regular WC7 star would be  $31 \pm 5 \text{ \AA}$  (Torres et al., 1986) where we measure  $67 \pm 2$ , which is closer to WC5 typical values (i.e.  $64 \pm 15 \text{ \AA}$ ). This justifies the “peculiar” denomination in the classification.

Criterion	Measurement	WC7 value	WC5 value
FWHM C IV 5808 (Å)	51±2	45±20	50±20
FWHM C IV+He I 5808,5876	71±2		
C IV 5808 / C III 5696 (log(ew))	0.35±0.01	+0.1 to +0.6	-1.1 to +1.5
C IV+He I 5808,5876 / C III 5696	0.543±0.007		
C III 5696 / O III-v 5590	0.82±0.04	≥0.1	-0.4 to 0.5

Table 2.IV: Comparison between measurements and classification criteria from Crowther et al. (1998) for the WR component.

lines and leads to a first estimation of the WR spectrum. We then subtract this spectrum from all the original spectra and co-add the differences this time in the O-star frame of reference, which leads to an estimation of the O spectrum. This spectrum is subtracted from the original spectra and the differences are co-added again in the WR frame of reference, and so on, until the co-added spectrum for each star converges, which generally happens in less than ten iterations. The result of the spectral separation is shown in Fig. 2.2. The whole OHP data set was used. As noted above, no revision is required for the WR-star spectral class. Concerning the determination of the O-star spectral class, the main criterion is the equivalent width (EW) ratio of He I  $\lambda$ 4471 to He II  $\lambda$ 4541 ( $W'$ ) (Conti and Alschuler, 1971). We measure  $\log(W') \simeq -0.37 \pm 0.09$ , which implies the O star is O5.5-O6, with a preference for O5.5, which is consistent with the observed absence of Si IV  $\lambda$ 4089.

The luminosity class determination is limited by the deteriorated information in the He II 4686 Å region. Indeed, this region is dominated by the residuals of the variability effects of the C III/IV+He II emission blend of the WR-star. One cannot really distinguish between absorption and emission for this line, although we do note that there is no narrow residual component, either emission or absorption, as expected from the O-star. Using the strength of the N III and C III emissions lines (which are narrow as required if arising in the O-star) in the vicinity of 4645 Å and based on the criteria in Walborn et al. (2010) we deduce a type of O5.5fc for the O star, with luminosity class somewhere between III and I (there is no II

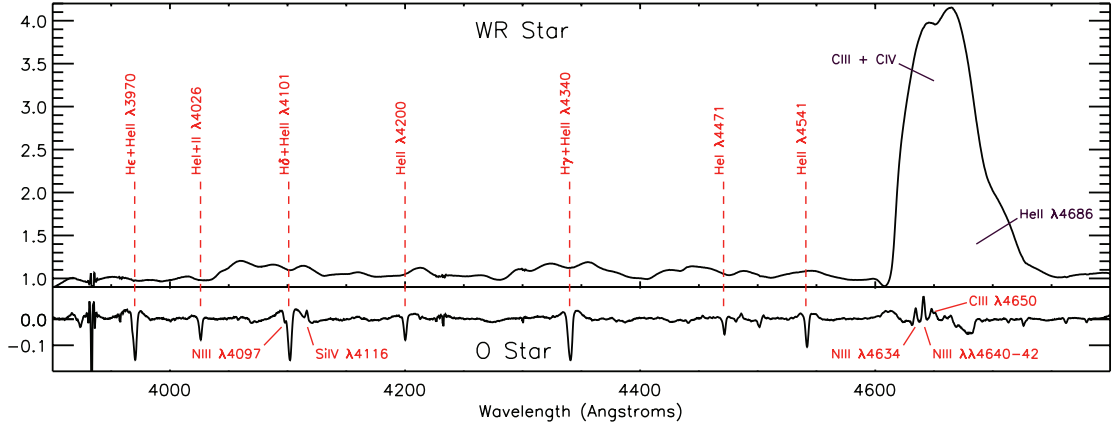


Figure 2.2: Result from the “shift-and-add” method for spectrum decomposition. The Y scale for the O-star has been expanded by a factor 3. Note that this technique does not yield absolute line strengths.

defined).

We now examine the line dilution of the O-star by the WR-star light. With conservation of line flux:

$$\text{Line Flux} = I_c(\text{O}+\text{WR}) \times ew_{\text{obs}} = I_c(\text{O}) \times ew_{\text{exp}} \quad (2.1)$$

where  $ew_{\text{obs}}$  and  $ew_{\text{exp}}$  are the observed and expected equivalent widths, respectively, for a single O-star of the same spectral type and  $I_c$  is the continuum intensity. We can then write :

$$\frac{I_c(\text{O})}{I_c(\text{O}+\text{WR})} = \frac{ew_{\text{obs}}}{ew_{\text{exp}}} \quad (2.2)$$

and,

$$\frac{10^{-0.4M_V(\text{WR})} + 10^{-0.4M_V(\text{O})}}{10^{-0.4M_V(\text{O})}} = 1 + 10^{-0.4[M_V(\text{WR})-M_V(\text{O})]} \quad (2.3)$$

where  $M_V$  designates here the absolute visual magnitude in the  $V$  band. We finally get the following estimation formula:

$$M_V(\text{WR}) - M_V(\text{O}) = -2.5 \log \left( \frac{ew_{\text{exp}}}{ew_{\text{obs}}} - 1 \right) \quad (2.4)$$

We measured the line EWs of the disentangled O-star spectrum and of two tem-

plate stars successively : HD 97253, a typical single O5.5 III star, and HD 163758, an O6.5f Ia with suggestion of a C over-abundance, which makes it quite close to our O star spectral type estimation. These spectra were both found in the UVES POP database (no O5.5/O6 I star was available). These measurements are shown in Table 2.V. We choose to apply equation 2.4 to the hydrogen lines only in order to minimize the effects of differences in helium content between the two stars. By averaging over the hydrogen lines, we obtain an equivalent width ratio  $ew_{\text{obs}}/ew_{\text{exp}} = 0.43 \pm 0.04$ , which yields a difference in magnitude of  $-0.3 \pm 0.2$  (WR slightly brighter than the O-star) with the O5.5 III template star. When making the same calculation on the He II lines, we obtain an average ratio of  $0.59 \pm 0.03$  and a difference in magnitude of  $0.4 \pm 0.1$  (O-star brighter).

If we take a typical O5.5 III absolute magnitude of  $M_V = -5.8$  from Martins et al. (2005), we get for the Wolf-Rayet star an absolute visual magnitude of  $-6.1 \pm 0.2$  with the H lines and  $-5.4 \pm 0.1$  with the He II lines, which is too luminous compared to a typical WC7 star magnitude of around  $-4.8$  (Smith et al., 1990) but might still be realistic given the peculiar nature the Wolf-Rayet in WR 140.

We now use the O6.5 Ia template. We get a difference in magnitude of  $0.7 \pm 0.4$  with the H lines only, and roughly the same with the He II lines. Given a typical O5.5 I magnitude of  $M_V = -6.3$  (Martins et al., 2005), this yields  $M_V = -5.6 \pm 0.4$  for the Wolf-Rayet which is again over-luminous compared to WC7 typical magnitudes.

We computed the corresponding photometric distances using an extinction of  $A_V = 2.27$  (van der Hucht, 2001)<sup>5</sup> in order to compare them with the value of Dougherty et al. (2005) (hereafter D05) (i.e.  $1.85 \pm 0.16$  kpc). We get for the class III template,  $d = 2.0 \pm 0.2$  kpc when using the H lines and  $d = 1.7 \pm 0.1$  kpc with the He II lines. For the class Ia template, we have  $d = 2.1 \pm 0.4$  kpc. These three values are compatible with D05.

It is also interesting to make the opposite reasoning and try to evaluate the

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<sup>5</sup>Here, we used the relation  $A_V = 1.1A_V$  as in Smith (1968)

Line	WR 140	HD 97253	ratio	HD 163758	ratio
H $\epsilon$ +He II $\lambda$ 3970	0.67 $\pm$ 0.01	1.54 $\pm$ 0.02	0.435 $\pm$ 0.009	0.98 $\pm$ 0.02	0.68 $\pm$ 0.02
He I+II $\lambda$ 4026	0.24 $\pm$ 0.01	0.40 $\pm$ 0.01	0.60 $\pm$ 0.03	0.60 $\pm$ 0.01	0.40 $\pm$ 0.02
H $\delta$ +He II $\lambda$ 4101	0.78 $\pm$ 0.01	1.83 $\pm$ 0.02	0.426 $\pm$ 0.007	1.42 $\pm$ 0.02	0.55 $\pm$ 0.02
He II $\lambda$ 4200	0.296 $\pm$ 0.005	0.48 $\pm$ 0.01	0.62 $\pm$ 0.02	0.43 $\pm$ 0.01	0.68 $\pm$ 0.02
H $\gamma$ +He II $\lambda$ 4340	0.778 $\pm$ 0.005	2.02 $\pm$ 0.01	0.385 $\pm$ 0.003	1.22 $\pm$ 0.01	0.637 $\pm$ 0.007
He I $\lambda$ 4471	0.187 $\pm$ 0.005	0.25 $\pm$ 0.01	0.75 $\pm$ 0.04	0.59 $\pm$ 0.01	0.32 $\pm$ 0.01
He II $\lambda$ 4541	0.395 $\pm$ 0.005	0.65 $\pm$ 0.01	0.61 $\pm$ 0.01	0.68 $\pm$ 0.01	0.58 $\pm$ 0.01
H $\beta$ +He II $\lambda$ 4861	0.891 $\pm$ 0.005	1.82 $\pm$ 0.01	0.490 $\pm$ 0.004	1.15 $\pm$ 0.01	0.775 $\pm$ 0.008
He II $\lambda$ 5412	0.536 $\pm$ 0.005	0.97 $\pm$ 0.02	0.55 $\pm$ 0.01	0.80 $\pm$ 0.02	0.67 $\pm$ 0.02
O V $\lambda$ 5590	0.177 $\pm$ 0.005	0.18 $\pm$ 0.02	1.0 $\pm$ 0.1	0.30 $\pm$ 0.02	0.59 $\pm$ 0.04

Table 2.V: Measured equivalent widths for the O-star component of WR 140 compared with the ones for the O5.5 III star HD 97253 and the O6.5 Ia star HD 163758.

line dilution of the WR-star by the O-star light. Based on the work of Torres et al. (1986) (their Fig. 2) and Smith et al. (1990) (their Table 5) and given the fact that WR 140 shows broader lines than standard WC7 stars (values more typical of WC5 stars: Torres et al. 1986), we chose an expected value of  $\log(ew_{\text{exp}}) \simeq 3.0$  for 4650 and 5880 (typical value for a “regular” WC7 star would be more like 2.2 to 2.5). The errors on these empirical values are about 0.3. Our measured values are  $\log(ew(4650)) \simeq 2.34$  and  $\log(ew(5808)) \simeq 2.21$  which leads to two estimations for the difference in magnitude between the stars: 1.4 and 1.1, respectively. This makes the O-star about three times brighter than the Wolf-Rayet in the optical band. If we assume a reasonable absolute  $V$  magnitude of  $-4.8$  for the WR-star (Smith et al., 1990), we have for the O-star  $M_V = -5.9$  to  $-6.2$ , which is compatible with both class I and III magnitudes (Martins et al., 2005).

As a conclusion, we are not able here to distinguish between class I and III for the O5.5/6 star. Note however that this task is complicated by the fact that the magnitude gap between these classes is very small for this particular spectral type (O5-6).

## 2.6 Excess emission

### 2.6.1 Lührs model fit

The shock cone around the O-star produces excess emission that can be easily observed on the C III 5696 and He I 5876 flat top lines<sup>6</sup>. This excess emission first appears just before periastron passage, on the blue side of the line, and then moves quickly to the red side just after periastron passage, before it disappears shortly thereafter. We measured this excess in the OHP data by subtracting a reference spectrum from each spectrum of the set. This reference spectrum has been constructed by co-adding the OHP spectra outside phases 0.98-1.02, unaffected by wind-wind collision effects. We took care, before doing the subtraction, to remove the O-star spectral signatures, notably the narrow emission feature on top of the C III 5696 line by fitting a Gaussian and subtracting it from each profile. The result of this procedure is shown in Fig. 2.3.

We can analyse the excess emission by using a simple geometric model of the CWR developed by Lührs (1997). The model takes into account the half-opening angle of the shock cone  $\theta$ , the velocity of the plasma along the cone  $v_{\text{strm}}$  and the orbital inclination  $i$ . The CWR is assumed to be curved in the plane of the orbit and its orientation over a limited range of the shock cone is parameterized by a single angle,  $\delta\phi_0$ , for simplicity (see Fig. 2.4). By a simple integration of the emissivity on a thick slice of the cone where the emission is supposed to arise, the model provides theoretical excess profiles showing double peaks moving with the orbit. However, depending on spectral resolution and the velocities and geometry of the shock cone, the double peaks might not be visible. Faced with this situation, one can still derive some valuable estimates of the colliding-wind region by measuring the entire excess emission. Hill et al. (2000, 2002) have shown that, for a circular orbit, the width (e.g. FWHM) and position of the excess emission

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<sup>6</sup>Flat top is the characteristic line shape of an emitting optically thin spherical shell in expansion.



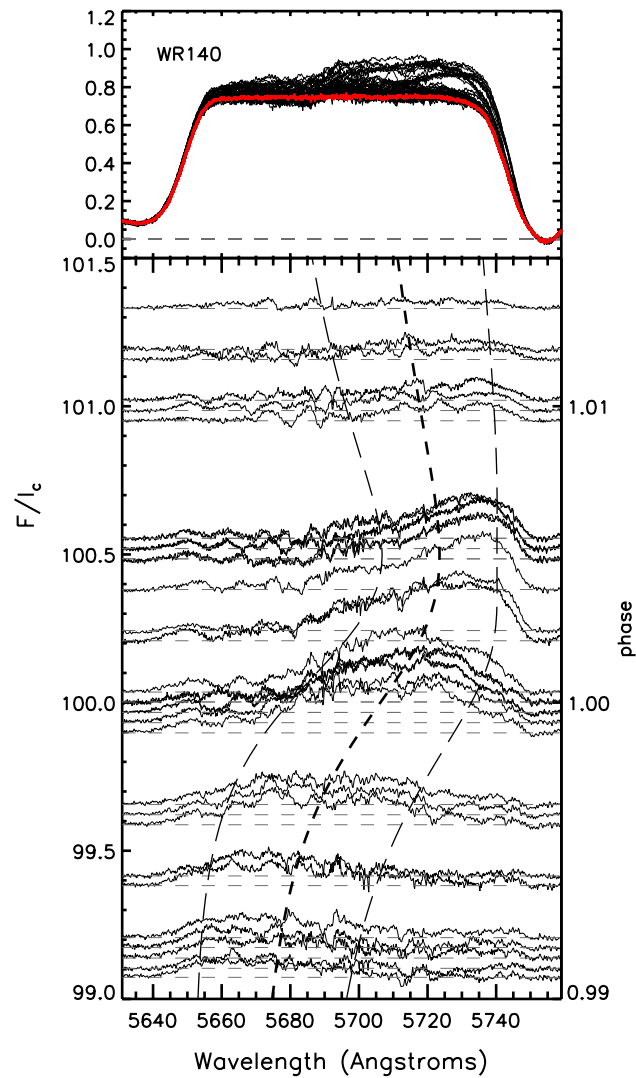


Figure 2.3: (top) C III 5696 reference profile compared with the ensemble of the spectra and (bottom) the resulting excess emission profiles as a function of phase. We also show the excess position (short dashed line) and width (horizontally between long dashed lines) arising from the modified Lührs model (see Fig 2.5 and Table 2.VI).

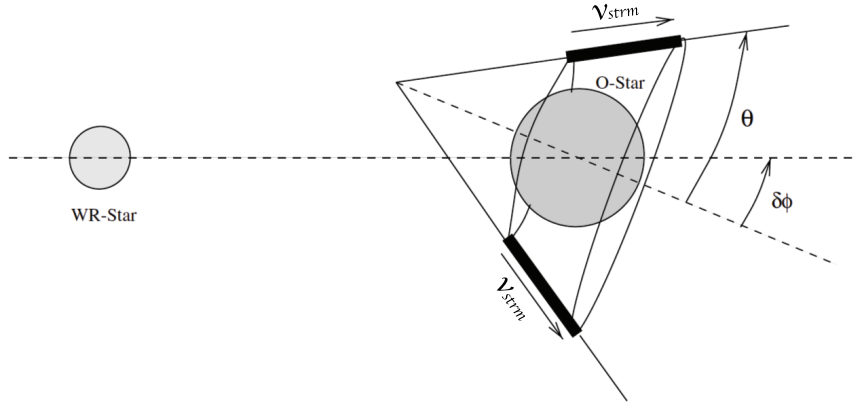


Figure 2.4: Schematic view of the geometric model by Lührs (1997) taken from Bartzakos et al. (2001) (their Fig. 2).

vary as

$$FW_{\text{ex}} = C_1 + 2v_{\text{strm}} \sin \theta \sqrt{1 - \sin^2 i \cos^2(\phi - \delta\phi_0)} \quad (2.5)$$

$$RV_{\text{ex}} = C_2 + v_{\text{strm}} \cos \theta \sin i \cos(\phi - \delta\phi_0) \quad (2.6)$$

where  $C_1$  and  $C_2$  are simple constants, and  $\phi$  is the orbital azimuthal angle (or true anomaly for  $\omega = 270^\circ$ ) in the case of a circular orbit :  $\phi = 360^\circ \times \text{phase}$  and the Wolf-Rayet star is in front at  $\phi = 0$ .

Unlike the cases for WR 42, WR 48, and WR 79 described by Hill et al. (2000, 2002), we face additional complications here. The orbit is not circular and we do not have full phase coverage. To circumvent the first complication, we can still fit versus true anomaly (calculated from the known orbit), and proceed as if the orbit was circular. We also need to replace  $\phi$  by  $\phi - [90^\circ - \omega]$  in equations 2.5 and 2.6, where  $\phi$  is now the true anomaly for an eccentric orbit :  $\phi = 0$  is the periastron passage and  $\phi = 90^\circ - \omega$  has the WR in front. The lack of phase coverage is not as bad as it appears at first sight, since the very limited phase interval over which excess emission is measurable actually corresponds to about 150 degrees of true anomaly. Using this model to fit the data, we obtain

$v_{\text{strm}} = 2090 \pm 90 \text{ km s}^{-1}$ ,  $i = 53 \pm 6^\circ$ ,  $\theta = 42 \pm 3^\circ$  and  $\delta\phi_0 = 21 \pm 3^\circ$  (see plot in Fig. 2.5). But we can go a bit further in modifying the Lührs model for an eccentric orbit by allowing the parameter  $\delta\phi_0$  (now denoted  $\delta\phi$ ) to vary throughout the orbital motion.

Two different effects will affect  $\delta\phi$ . The first one is the aberration effect (or Coriolis effect); this will produce an angular shift (Moffat et al., 1998):

$$\delta\phi = \arctan\left(\frac{v_\theta}{v_{\text{wind}}(\text{WR}) + v_{\text{wind}}(\text{O})}\right)$$

where  $v_\theta$  is the azimuthal component of the orbital velocity of the O-star in the WR frame of reference (or inversely). This velocity is deduced from the angular momentum conservation :

$$rv_\theta = J = \sqrt{G(M_{\text{WR}} + M_{\text{O}})a(1 - e^2)}$$

and

$$r = \frac{a(1 - e^2)}{1 + e \cos \phi}$$

However, in our case, this effect will only contribute to a maximum angular shift of  $2.6^\circ$  at periastron, and fall quickly to 0 subsequently. This is far from the  $22^\circ$  needed to fit the data.

The second factor contributing to  $\delta\phi$  is the curvature of the shock surface. If we assume that after being shocked at time  $t$ , the wind plasma moves in a rectilinear motion with a velocity  $v_{\text{strm}}$  along the ‘local’ Lührs cone, and that the distance from the apex to the excess forming region (=EFR) is constant throughout the orbit, it will take a time  $\delta t$  (or a phase fraction  $\delta p = \delta t/P$ ) for this plasma to reach the EFR. The ‘local’ Lührs cone emitting the excess at phase  $p$  will therefore appear shifted by an angle :

$$\delta\phi = \phi(p) - \phi(p - \delta p)$$

We can then use the parameter  $\delta p$  (i.e. a constant phase shift) to fit instead of  $\delta\phi$  (a shift in true anomaly), which has the advantage of being consistent with the orbit. Using equations 2.5 and 2.6 and the modifications described above (Coriolis + phase shift  $\delta p$ ), we have fitted the measurements of the width and position of the excess emission shown in Fig. 2.5. The evolution of  $\delta\phi$  with phase is shown in Fig. 2.6. Note that, even if the original Lührs model seems to give a slightly better fit, the modified model is more physically correct and allows an estimation of the distance of the EFR relative to the shock cone apex (see section 2.6.2).

The position of the excess was computed with the first moment formula :

$$RV = \frac{c(\bar{\lambda} - \lambda_0)}{\lambda_0} \text{ with } \bar{\lambda} = \int_{\lambda_1}^{\lambda_2} F(\lambda)\lambda d\lambda / \int_{\lambda_1}^{\lambda_2} F(\lambda) d\lambda \quad (2.7)$$

where  $\lambda_0$  is the central wavelength of the line and  $\lambda_1, \lambda_2$  are the limits of the excess profile.

Instead of the full width at half maximum, too sensitive to the irregularities of the profiles, the width was computed as the second central moment of the excess :

$$FW = \frac{2.35c}{\lambda_0} \left[ \int_{\lambda_1}^{\lambda_2} F(\lambda)(\lambda - \bar{\lambda})^2 d\lambda / \int_{\lambda_1}^{\lambda_2} F(\lambda) d\lambda \right]^{1/2} \quad (2.8)$$

where 2.35 is the factor linking the FWHM and  $\sigma$  for a Gaussian profile (FWHM=2.35  $\times$   $\sigma$ ). Of course, our profiles are not strictly Gaussians but are not extremely different either. The errors were obtained by applying the error propagation formulae to Eqs 2.7 and 2.8 and taking into account the Poisson noise and readout noise of  $F(\lambda)$ .

Reasonable fits are obtained for a very wide range of parameters, but some limits can be derived by making a few realistic assumptions. For example, we can require that the velocity of the gas streaming along the bow shock cone,  $v_{\text{strm}}$ , does not exceed the terminal velocity of the WR wind,  $v_{\infty}$ . Eenens and Williams (1994) find  $v_{\infty} = 2900 \text{ km s}^{-1}$  for the WR-star component of WR 140. Another possible constraint can be obtained by assuming that turbulence in the shock cone contributes to the constant  $C_1$  in equation 2.5. Hill et al. (2002) found that

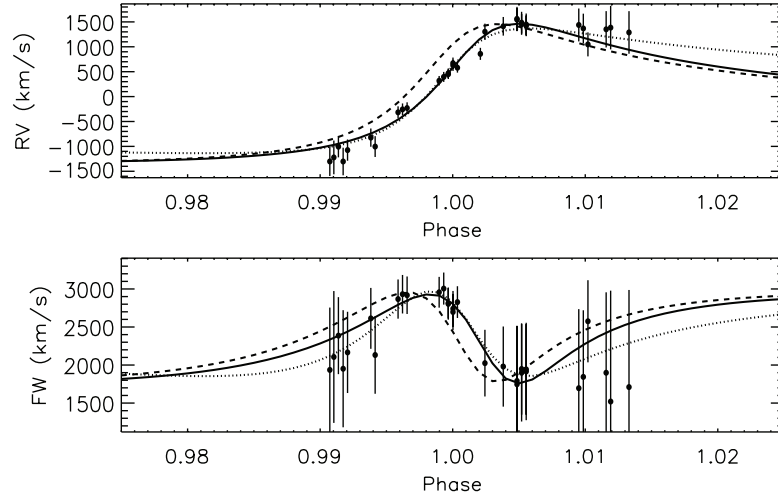


Figure 2.5: Fit of the RV and width of the excess using the original (dotted line) and the modified Lührs model (full line). The dashed line shows the solution from M03. The resulting parameters of the fit are listed in Table 2.VI. The measured RV and FW are presented in Table II.II.

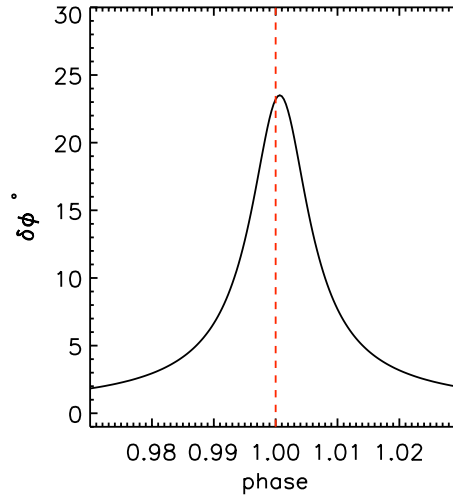


Figure 2.6: The angle  $\delta\phi$  as a function of phase resulting from the combined action of the Coriolis effect (nearly negligible) and a constant phase shift of 0.0014 (or 4.0 days) corresponding to the flow time from the apex to the EFR.

numbers of the order of 200-500  $\text{km s}^{-1}$  were appropriate for WR 42, WR 48, and WR 79 in this respect.

We finally find  $v_{\text{strm}} = 2170 \pm 100 \text{ km s}^{-1}$ ,  $i = 55 \pm 6^\circ$ ,  $\theta = 39 \pm 3^\circ$ ,  $\delta p = 0.0014 \pm 0.0002$  corresponding to  $\delta\phi$  at periastron =  $23^\circ$  (vs.  $v_{\text{strm}} = 2300 \pm 500 \text{ km s}^{-1}$ ,  $i = 50 \pm 15^\circ$ ,  $\theta = 40 \pm 15^\circ$ ,  $\delta\phi_0 = 40 \pm 20^\circ$  from M03). Our value of  $i$  is compatible with that from D05,  $i = 122 \pm 5^\circ$ , equivalent to  $180 - i = 58^\circ$ , given the fact that we cannot distinguish between  $i$  and  $180 - i$  with the Lührs fit alone. The errors on the parameters are those computed by MPFIT, using the errors on  $F(\lambda)$  given in input (typically around 0.006 in continuum units). Applying the approach of Cantó et al. (1996) (their equation 28), we find that our value of  $\theta$  leads to  $\eta = [\dot{M}(\text{O})v_\infty(\text{O})]/[\dot{M}(\text{WR})v_\infty(\text{WR})] = 0.039 \pm 0.010$  (vs. 0.045 from M03 and 0.02 from Pittard and Dougherty (2006)). The difference between the estimations of  $\eta$  are discussed in section 2.6.2. Our value for the inclination finally gives the following estimation for the stellar masses :  $M_{\text{WR}} = 16 \pm 3 M_\odot$  and  $M_{\text{O}} = 41 \pm 6 M_\odot$  (vs.  $19 M_\odot$  and  $50 M_\odot$  from M03). The O-star mass agrees well with O5.5 stars between luminosity class III and I given by Martins et al. (2005) (as deduced in section 2.5). The WR mass can be compared with measurements of other WC type stars of the “VIIth catalogue of galactic Wolf-Rayet stars” (van der Hucht, 2001). This latter gives for WC6-8 type stars, masses between  $9.5 M_\odot$  ( $\gamma$  Vel) and  $16 M_\odot$  (HD 94305). A list of all the parameters deduced in this study is presented in Table 2.VI.

The same analysis done for the He I 5876 line shows qualitatively the same behaviour of the excess emission. However, the RV and FWHM measurements are of lesser quality and no additional information was deduced. The inferior quality is due mainly to the blending of He I with the C IV 5806 emission feature (also restricting the placement of the continuum), along with residuals from imperfect removal of telluric lines and the blaze correction.

An important remark is that whereas the measurement of the position of the excess is quite robust (even for the He I 5876 line), its width is very dependent on the details of the measurement (boundaries in the moment integral, minimum

flux for pixel selection in the integral, etc.). The errors on the width have been artificially increased in order to reflect this source of uncertainty. We should also note several caveats when deducing the wind momentum ratio from half-opening angles. First, the opening angles of the O-star and WR-star shocks are in general different, and different from that of the contact discontinuity (only in highly radiative systems where the interaction region is very thin are all three identical). Usually, it is assumed that the measured half-opening angle is that of the contact discontinuity, and the wind momentum ratio is then derived using e.g. the equation in Cantó et al. (1996). However, it is not so obvious that this is correct. We know, for instance, that in an adiabatic system, the total emission is dominated by the stronger wind (Stevens et al., 1992). In our case, the CWR is suspected to be radiative at the time where the excess emission arises, as the  $1/r^2$  law in Fig. 2.7 suggests, but it would certainly turn adiabatic later in the orbit as the two stars get more distant. The approach of Gayley (2009) for purely adiabatic shocks predicts, for our value of  $\eta$ , a half-opening angle of  $\theta = 48^\circ \pm 3$  in the case where no mixing occurs through the contact discontinuity (his eq. 11), and closely the same value for homogeneous mixing (his eq. 17) with a reasonable value of the terminal velocities ratio of 0.9. The augmentation of  $\theta$  when turning adiabatic is due to the explosive gas pressure and the absence of significant radiative cooling (Gayley, 2009). However, the details of the WWC process, and especially the adiabatic or radiative nature of the shocks need further investigation.

### 2.6.2 Discussion

One can see that our estimation of the wind momentum ratio  $\eta$  is consistent with the estimation of M03. Our value is also in line with the estimation by Pittard and Dougherty (2006) (i.e.  $\eta = 0.02$ ) based on radio emission modelling. These latter authors computed mass-loss rates required to match the observed X-ray flux at 3 keV as a function of  $\eta$  (their Table 2). According to this study, our value of  $\eta$  gives  $\dot{M}(\text{WR}) \simeq 3 \times 10^{-5} \text{ M}_\odot \text{ yr}^{-1}$  and  $\dot{M}(\text{O}) \simeq 1 \times 10^{-6} \text{ M}_\odot \text{ yr}^{-1}$ . The terminal velocities assumed in the calculations are  $v_\infty(\text{WR}) = 2860 \text{ km s}^{-1}$  (D05)

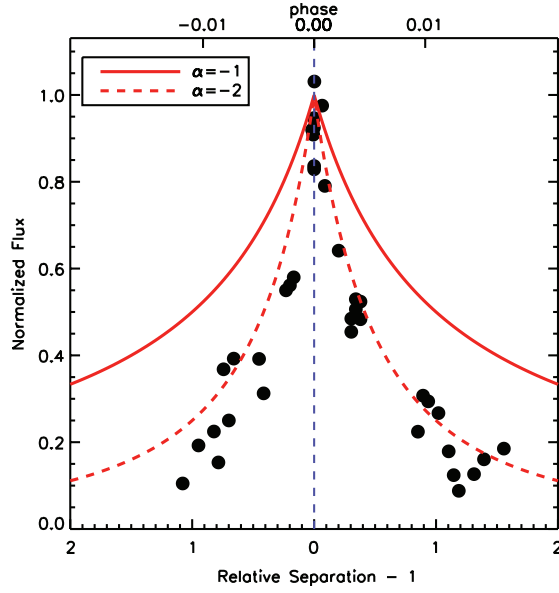


Figure 2.7: Normalized flux of the excess as a function of the relative separation of the two stars ( $[d - d_{\min}]/d_{\min}$ ). The full line shows a  $d^{-1}$  dependency, expected for an adiabatic emission process. The dashed line shows a  $d^{-2}$  dependency, possibly more in line with an isothermal radiative process.

Parameter	This paper		M03
	modified model	original model	
$i$ (deg.)	$55 \pm 6$	$53 \pm 6$	$50 \pm 15$
$C_1$	200	200	200 – 500
$C_2$	$71 \pm 55$	$117 \pm 61$	100 – 200
$v_{\text{strm}}$ (km s $^{-1}$ )	$2170 \pm 100$	$2090 \pm 90$	$2300 \pm 500$
$\theta$ (deg.)	$39 \pm 3$	$42 \pm 3$	$40 \pm 15$
$\delta\phi$ (deg.)	23 (at phase=0)	$21 \pm 3$	$40 \pm 20$
$\delta p$	$0.0014 \pm 0.0002$		
$a$ [WR] (A.U.)	$10.9 \pm 0.8$	$11.2 \pm 0.9$	$13.5 \pm 3.0$
$a$ [O] (A.U.)	$4.4 \pm 0.3$	$4.6 \pm 0.4$	$5.0 \pm 1.1$
$M$ [WR] ( $M_{\odot}$ )	$16 \pm 3$	$18 \pm 3$	19
$M$ [O] ( $M_{\odot}$ )	$41 \pm 6$	$44 \pm 8$	50
radiative $\eta$	$0.039 \pm 0.010$	$0.051 \pm 0.0$	$0.045 \pm 0.075$

Table 2.VI: New parameters for WR 140, deduced from the modified Lührs model and from the original model, compared with the values from M03.



and  $v_\infty(\text{O}) = 3100 \text{ km s}^{-1}$  (Setia Gunawan et al., 2001). One can argue that the comparison between the two determinations of  $\eta$  are biased by the fact that we are looking at diagnostics from two very different regions of the shock: very close to the apex for the radio emission, where the shocks are strongest (hence most efficient in accelerating particles), and much further away for the optical excess emission. The similarity between our value and the one from Pittard and Dougherty (2006) would therefore imply that the opening angle is rather constant along the cone. Again, note that these opening angles refer to the ‘local’ Lührs cones defined by the orientation of the flow velocities at a radius  $r(t)$  from the apex.

Our excess fit gives a value of the phase fraction  $\delta p$  for the plasma to reach the excess forming region which can be translated into a time  $\delta t = 4.0 \pm 0.6$  days. With our value for the streaming velocity  $v_{\text{strm}}$  (assumed constant), this would make the distance from the apex to the zone where the excess arises to be  $d_{\text{EFR}} = 7.6 \times 10^8$  km or 5.1 A.U. This value could be compared with simulations of colliding-winds taking into account adiabatic/radiative cooling of the CWR. However, we can also try to get a rough estimate by using a simpler analysis. The crucial parameter is  $\chi$  (defined by Stevens et al. 1992). This gives the ratio of the cooling timescale to the flow timescale (roughly the time for the gas to flow a distance  $d_{\text{sep}}$  downstream), which indicates how radiative or adiabatic the wind-wind collision should be.

$$\chi \approx v_8^4 d_{12} / \dot{M}_{-7} \quad (2.9)$$

where  $v_8 = v/1000 \text{ km s}^{-1}$ ,  $d_{12} = d_{\text{sep}}/10^{12} \text{ cm}$ , and  $\dot{M}_{-7} = \dot{M}/10^{-7} \text{ M}_\odot \text{ yr}^{-1}$ .

One can substitute parameters for either the WR or O-star as required. For WR 140, if we take  $\eta = 0.0353$  from Table 2 in Pittard and Dougherty (2006) then  $\dot{M}(\text{WR}) = 3.3 \times 10^{-5} \text{ M}_\odot \text{ yr}^{-1}$ . Using  $v_\infty(\text{WR}) = 2860 \text{ km s}^{-1}$  and  $d_{\text{peri}} = 1.59 \text{ A.U.}$  ( $d_{12} = 23.8$ ) then  $\chi \approx 4.8$ . So the gas should flow downstream a distance of about  $4.8 d_{\text{peri}}$  before cooling (7.6 A.U.). This is relatively close to the estimate in this paper (5.1 A.U.) but is probably fortuitous since there are many other factors,

which mean it should only be used as a rough guide (e.g. the determination of  $\chi$  in Stevens et al. (1992) assumes solar abundances - in reality the WC7 gas will cool more quickly; clumping will also decrease the cooling time (at least within the clumps) - but the clumps may be destroyed even faster - see Pittard (2007); gas colliding away from the line-of-centres of the stars will be of lower density and will take longer to cool).

## 2.7 Conclusion

The 2009 periastron campaign on WR 140 provided increased accuracy on the orbital parameters, new estimates for the WR and O-star masses and new constraints on the relative mass-loss rates. For this purpose, we have modified the geometrical model of Lührs (1997) in order to adapt it to an eccentric orbit. This new version of the model also gives an estimation of the distance of the excess forming region from the apex of the shock cone of about 5.1 A.U. which could be compared with hydro/radiative simulations of the wind-wind collision. The observational campaign is also a very encouraging success in terms of professional-amateur collaboration and we hope it will give rise to similar initiatives in the future. A more sophisticated theoretical investigation should also be done to confirm our results about the geometry of the shock cone. Meanwhile, the  $\sim d^{-2}$  dependency of the excess, shown in Fig. 2.7, strongly suggests that some kind of isothermal radiative process is involved here, rather than adiabatic, despite the large separation of the two stars even at periastron. Links with observations in other spectral domains (X, IR and radio) will certainly provide valuable clues about the physics. Finally, we have attempted to isolate the WR spectrum from the O-star spectrum from our data in order to identify the spectral type of the latter more precisely. Additional photometric and spectropolarimetric data obtained during the campaign will be presented in a future publication.

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## CHAPTER 3

### COLLIDING WINDS IN FIVE WR + O SYSTEMS OF THE SOUTHERN HEMISPHERE

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## ABSTRACT

We present the results of an optical spectroscopic follow-up of five WR+O systems of the southern hemisphere that have not yet been previously studied for wind-wind collision: WR 12 (WN8h), 21 (WN5o+O7V), 30 (WC6+O7.5V), 31 (WN4o+O8) and 47 (WN6o+O5). Observations were made over a whole contiguous month in march/april 2009 using the Garrison spectrograph attached to the 0.6-m Helen Sawyer Hogg (HSH) telescope at Complejo Astronomico El Leoncito (CASLEO). We obtain improved orbital elements for these five systems and provide an estimation of the orbital inclination and shock-cone opening angle for four of them by analysing excess emission profiles with the geometrical model of Lührs (1997). We argue that the line-variability in one of the stars (WR 12) might be dominated by some dynamic instabilities in the shocked plasma, making the model inappropriate in this case. A sixth star was observed in the program, WR 69 (WC9d+OB), based on its published photometric period of  $2.293 \pm 0.005$  d (Marchenko et al., 1998b). While we found a probable spectroscopic period of  $2.15 \pm 0.04$  d compatible with Marchenko et al., 1998b, the period is unlikely related to orbital effects, rather to (relatively fast) rotation of the WC9d component of a much longer binary system.

**Keywords :** (stars:) binaries: general, stars: fundamental parameters (classification, colors, luminosities, masses, radii, temperatures, etc.), stars: Wolf-Rayet , stars: winds, outflows

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### 3.1 Introduction

Binary systems provide the least model-dependent way to secure measurements of stellar masses. Understanding binary systems is therefore of paramount

importance for many problems in stellar astrophysics. In the case of massive stars and most particularly WR+O systems, the strong, fast winds of the two stars will interact with each other along a shock surface, giving rise to excess emission in a range of wavelengths from radio to X-rays, possibly even  $\gamma$ -rays (Moffat and St-Louis, 2002). Understanding the underlying physics of this phenomenon allows us to choose the lines best suited to obtain orbital masses and gives us access to additional parameters of the system, like the wind momentum ratio and the orbital inclination. We present the results from an optical spectroscopic campaign on a selection of five well known short-period WR+O binaries not yet studied for colliding-wind effects, and, in an appendix, detection of a non-orbital short period in WR 69.

### 3.2 Observational data

Observations were made using the Garrison spectrograph (Garrison et al., 1993) attached to the 0.6-m Helen Sawyer Hogg (HSH) telescope at Complejo Astronomico El Leoncito (Argentina) between march 15 and april 15 2009. This telescope is owned by the University of Toronto, Canada, and was originally in operation (between 1971 and 1997) at Las Campanas Observatory, Chile. The targets were chosen to have a known (or suspected) period of less than a month in the catalog of van der Hucht (2001) (i.e. the duration of the campaign) and had not yet been studied for wind-wind collision. We ended with the following selection : WR 12, WR 21, WR 30, WR 31 and WR 47. A sixth star, WR 69 (WC9d+OB) was also selected based on its short period; however it was found not to be an orbital period as presented in the appendix. Four additional stars were observed but not treated in this paper : three long-period (or suspected) eccentric systems, R144, R145 and WR 25 (Schnurr et al., in prep.) as well as the rotating single star WR 110 (Chené et al., 2011). Each star was observed every clear night with a signal-to-noise ratio of about 100, in the spectral range of 3850 to 5300 Å, in order to include He II 4686, known to be particularly sensitive to wind-wind collision effects. We used a dispersion of 1 Å/pixel, which lead to



a resolution of about  $3 \text{ \AA}$  ( $\simeq 3$  pixels). A summary of the data acquired in the campaign is presented in Table 3.I.

### 3.3 Data reduction

#### 3.3.1 Pre-reduction

Each set of data was first pre-reduced the standard way using the CCDPROC package (IRAF<sup>1</sup>) and following the procedure described in P. Massey’s “User’s Guide to Reducing Slit Spectra with IRAF” (Massey et al., 1992) : bias subtraction, normalized flat division and cosmic ray removal using the routine COSMICRAYS.

#### 3.3.2 Continuum normalization

Instead of the more classical method that consists of computing the continuum in each spectrum individually, an alternative method was chosen to normalize the spectra to their continuum, more adapted to Wolf-Rayet stars and their very few (and uncertain) continuum regions. In a first step, we constructed, for each star, a reference spectrum by averaging the best spectra of the set (no defects, no moon, etc.) and divided all the spectra by this reference. We obtained a series of nearly

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<sup>1</sup>IRAF is the Image Reduction and Analysis Facility, a general purpose software system for the reduction and analysis of astronomical data. IRAF is written and supported by the IRAF programming group at the National Optical Astronomy Observatories (NOAO) in Tucson, Arizona. NOAO is operated by the Association of Universities for Research in Astronomy (AURA), Inc. under cooperative agreement with the National Science Foundation (<http://iraf.noao.edu/>).

Star Name	HD/CD	Spectral type	ref.*	Period (d.)	ref.*	No. spectra
WR 12	CD-45 4482	WN8h	(s96)	$23.923 \pm 0.001$	(r96)	25
WR 21	HD 90657	WN5o+O4-6	(s96)	$8.2546 \pm 0.0001$	(g99)	24
WR 30	HD 94305	WC6+O6-8	(n83)	$18.82 \pm 0.01$	(n83)	19
WR 31	HD 94910	WN4o+O8V	(s96)	$4.8306 \pm 0.0001$	(g99)	19
WR 47	HD 311884	WN6o+O5V	(s96)	$6.2393 \pm 0.0001$	(g99)	25

\* s96 – Smith et al. (1996), n83 – Niemela et al. (1983), r96 – Rauw et al. (1996), g99 – Gamen and Niemela (1999).

Table 3.I: Summary of the five selected objects observed during the 2009 campaign at CASLEO.

flat profiles that we fitted with low order Legendre polynoms (degree 3 or 4). Each spectrum was then divided by its corresponding Legendre fit. At this point, all the spectra are at the same level and their continua are superposed when plotted on the same graph. In a second step, we determined the continuum of the reference spectrum using the CONTINUUM routine in IRAF. This continuum function (we used a cubic spline of order 6 to 7) was then used to divide into the whole set of spectra. This method insures a maximum of homogeneity in the continuum normalization, which is of particular importance when measuring variable features over strong emission lines like those in WR spectra. The normalized mean spectra for each star are shown in Fig. 3.1

### 3.4 Radial velocities

Radial velocity measurements and orbit determinations have already been performed, with varying degrees of accuracy, for our 5 stars. We combined our measurements with those previously published, which led to an increased precision in some cases or simply to a confirmation of the orbit in other cases. Given the relatively short periods and their clearly sinusoidal shapes, all the RVs were fitted assuming circular orbits. We chose the primary to be the WR star and phase 0 to be the time when the WR is at inferior conjunction.

#### 3.4.1 Wolf-Rayet star

The Wolf-Rayet star radial velocities were determined by measuring the centroids of emission lines, avoiding those that vary like He II 4686. Among the non-varying lines, N IV 4057.80 is known to be the best indicator of the motion of WNL components in binary systems (Shylaja, 1987). When this line was too faint to give satisfactory results, we used N V 4603.2 or N III 4640.64. For WR 30, the only strong line available was the C III/IV 4640 blend.

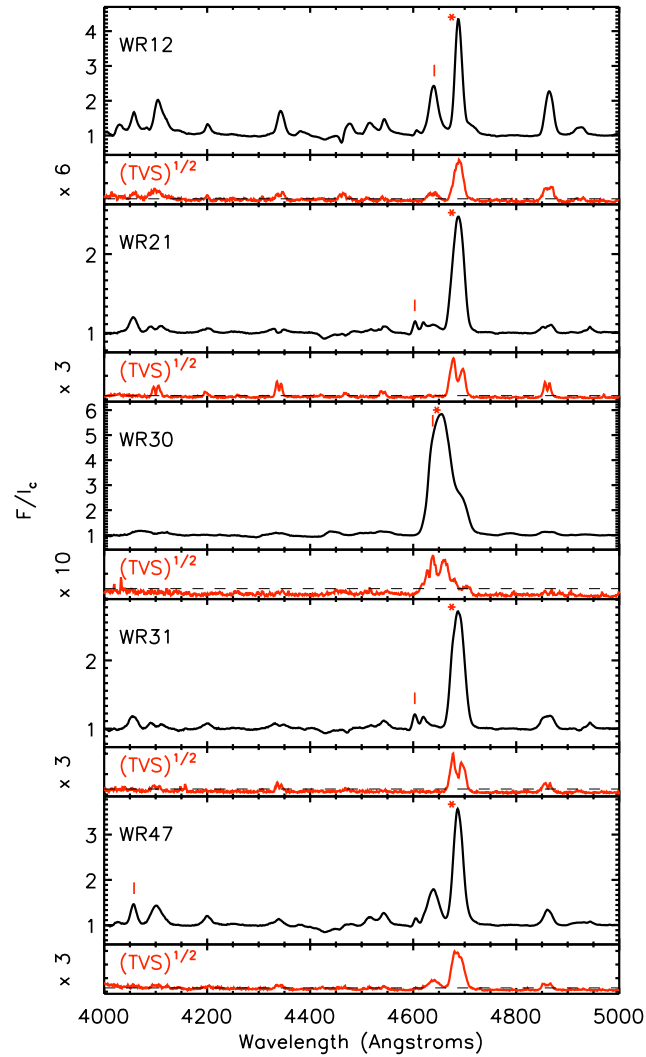


Figure 3.1: Temporal variance spectrum (TVS) analysis for each of the WR stars of the Leoncito campaign and comparison with the mean spectra computed in the WR rest frame. For the TVS curve, we indicated the enlargement factor of the y-axis relative to the above mean spectrum curve ( $\times 6$ ,  $\times 3$ , etc.). The dashed line represents the 99% confidence level for significant profile variability. The lines used for RV determination are identified with a ‘-’, and the ones used for the CWB analysis are marked with a ‘\*’.

### 3.4.2 O-star

The O-star RVs could only be successfully measured using the current data on one star, WR 21, the absorption lines being too weak to measure in all the other cases. For this star though, the RVs were obtained by measuring the centroid of each relatively narrow photospheric absorption line (allowing for a sloping pseudo continuum if superposed on WR lines), and taking the mean of the deduced velocities for each spectrum. We took into account a possible shift between the RVs determined from different O-star lines. This was done by choosing one RV set as a reference and by plotting the other sets versus this reference. By fitting a linear curve of slope 1 (i.e.  $y=x+b$ ), we could then deduce the velocity shift (i.e.  $b$ ) to be subtracted from the considered set. Typical velocity shifts have values from 5 to 10  $\text{km s}^{-1}$ . Lines used were: H I 3970.09, 4101.74, 4340.47, 4861.33; He I 4026.19, 4471.48; He II 4199.87, 4541.63. For the other four stars, we used previously published O-star RVs.

A plot of the radial velocities for each star is presented in Fig. 3.2. The corresponding circular orbital elements are presented in Tables 3.II-3.VI.

<b>WR 12</b>		g99-[N III 4641]	r96-[N IV 4058]	tp-[N III 4641]	Overall combined
P	[d]	$23.923 \pm 0.001$	$23.923 \pm 0.001$	$23.923 \pm 0.001$	$23.92336 \pm 0.00001$
$\gamma$ WR	[ $\text{km s}^{-1}$ ]	$-129.6 \pm 2.1$	$-25.0 \pm 2.0$	$-8.0 \pm 4.2$	-
$K$ WR	[ $\text{km s}^{-1}$ ]	$144.8 \pm 3.1$	$139.2 \pm 3.8$	$143.2 \pm 6.0$	$141.9 \pm 2.2$
$E_0$	[RJD]	$49811.3 \pm 0.1$	$49810.61 \pm 0.11$	$49811.79 \pm 0.16$	$49811.188 \pm 0.065$
$a \sin i$ WR	[ $R_\odot$ ]	$69 \pm 2$	$66 \pm 2$	$68 \pm 3$	$67 \pm 1$
$\sigma$ WR	[ $\text{km s}^{-1}$ ]	18	22	18	
$f(m)$ WR	[ $M_\odot$ ]	$7.5 \pm 0.5$	$6.7 \pm 0.6$	$7.3 \pm 0.9$	$7.1 \pm 0.3$

g99 – Data from Gamen (2004)  
r96 – Data from Rauw et al. (1996)  
tp – This paper. Data from Leoncito.

Table 3.II: Circular orbital elements of WR 12.

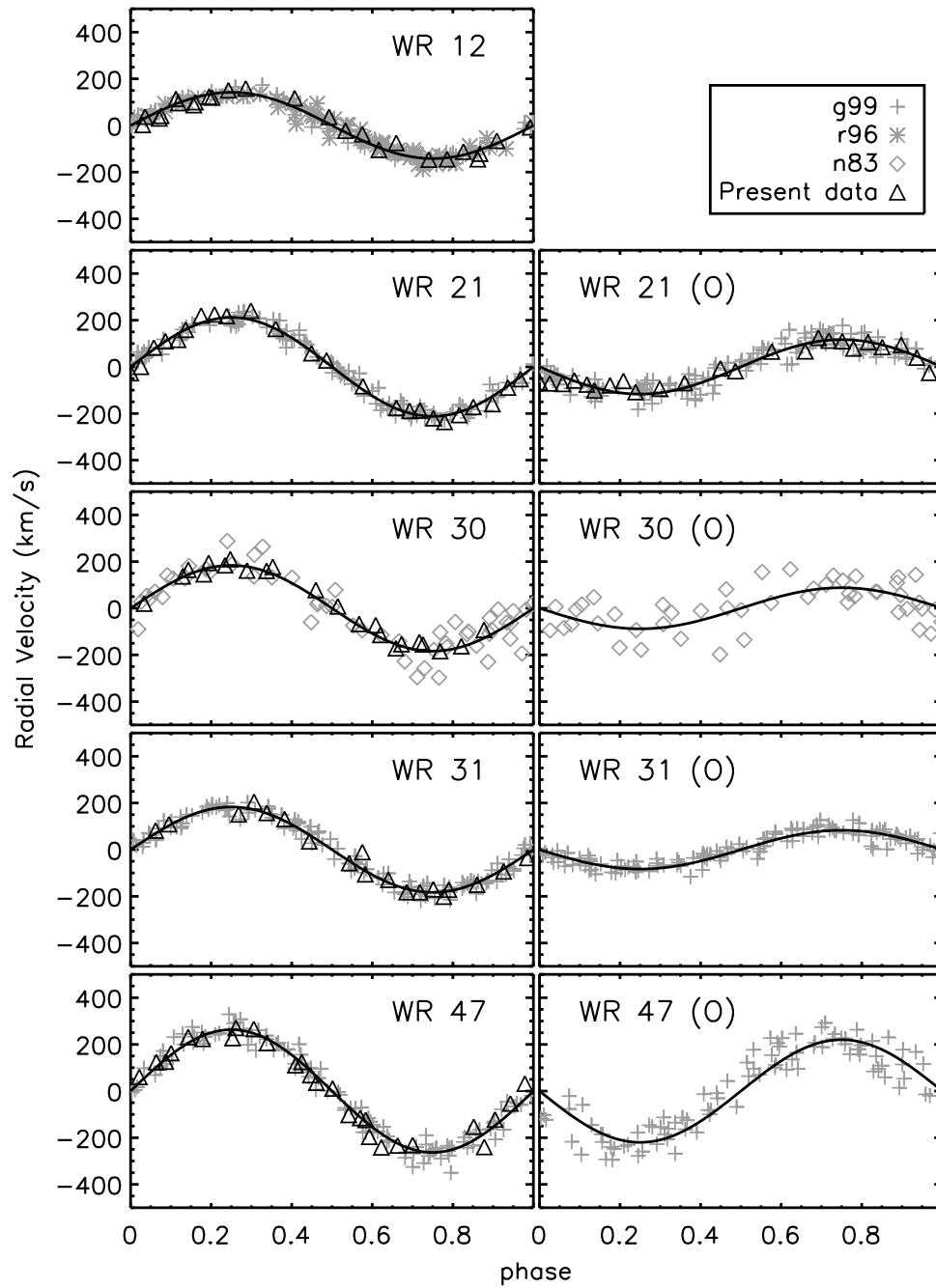


Figure 3.2: Measured radial velocities of the WR and O-star together with our orbital solution (full lines) whose simultaneous best-fit parameters are given in Tables 3.II-3.VI. The data source references are also given in these tables.

<b>WR 21</b>		g99-[N v 4603 + abs.]	tp-[N v 4603 + abs.]	Overall combined
P	[d]	8.2546 ± 0.0001	8.2546 ± 0.0001	8.25443 ± 0.00003
$\gamma$ WR	[km s <sup>-1</sup> ]	36 ± 3	106 ± 4	-
$\gamma$ O	[km s <sup>-1</sup> ]	-20 ± 2	50 ± 5	-
$K$ WR	[km s <sup>-1</sup> ]	209 ± 5	221 ± 6	211.8 ± 2.5
$K$ O	[km s <sup>-1</sup> ]	-119 ± 3	-105 ± 7	-116.9 ± 2.5
$E_0$	[RJD]	43907.28 ± 0.02	43907.136 ± 0.037	43907.277 ± 0.024
$a \sin i$ WR	[R <sub>⊙</sub> ]	34.1 ± 0.8	36 ± 1	34.6 ± 0.4
$a \sin i$ O	[R <sub>⊙</sub> ]	19.4 ± 0.5	17 ± 1	19.1 ± 0.4
$\sigma$ WR	[km s <sup>-1</sup> ]	21	16	
$\sigma$ O	[km s <sup>-1</sup> ]	33	24	
$f(m)$ WR	[M <sub>⊙</sub> ]	7.8 ± 0.6	9.3 ± 0.8	8.1 ± 0.3
$f(m)$ O	[M <sub>⊙</sub> ]	1.4 ± 0.1	1.0 ± 0.2	1.37 ± 0.09

g99 – Data from Gamen and Niemela (1999), Niemela (1976) and Niemela and Moffat (1982)  
 tp – This paper. Data from Leoncito.

Table 3.III: Circular orbital elements of WR 21.

<b>WR 30</b>		n83-[C IV 4441 + H $\delta$ abs.]	tp-[C III 4640]	Overall combined
P	[d]	18.82 ± 0.01	18.82 ± 0.01	18.8061 ± 0.0004
$\gamma$ WR	[km s <sup>-1</sup> ]	138 ± 7	10.8 ± 3.9	-
$\gamma$ O	[km s <sup>-1</sup> ]	-69 ± 10	50 ± 5	-
$K$ WR	[km s <sup>-1</sup> ]	195 ± 11		183.8 ± 4.6
$K$ O	[km s <sup>-1</sup> ]	-91 ± 15		-88 ± 15
$E_0$	[RJD]	44260.66 ± 0.1	44260.76 ± 0.2	44261.0 ± 0.2
$a \sin i$ WR	[R <sub>⊙</sub> ]	73 ± 4	67 ± 2	68 ± 1
$a \sin i$ O	[R <sub>⊙</sub> ]	34 ± 6		33 ± 5
$\sigma$ WR	[km s <sup>-1</sup> ]	58	18	
$\sigma$ O	[km s <sup>-1</sup> ]	68		
$f(m)$ WR	[M <sub>⊙</sub> ]	15 ± 3	12 ± 1	12.1 ± 0.5
$f(m)$ O	[M <sub>⊙</sub> ]	1.5 ± 0.7		1.3 ± 0.6

n83 – Data from Niemela et al. (1983)  
 tp – This paper. Data from Leoncito.

Table 3.IV: Circular orbital elements of WR 30.

<b>WR 31</b>		g99-[N v 4603 + abs.]	tp-[N v 4603]	Overall combined
P	[d]	4.8306 ± 0.0001	4.8306 ± 0.0001	4.830657 ± 0.000013
$\gamma$ WR	[km s <sup>-1</sup> ]	18 ± 3	77 ± 5	-
$\gamma$ O	[km s <sup>-1</sup> ]	-39 ± 2		-
$K$ WR	[km s <sup>-1</sup> ]	183 ± 4	184 ± 6	183 ± 3
$K$ O	[km s <sup>-1</sup> ]	-83 ± 3		-83 ± 3
$E_0$	[RJD]	45378.90 ± 0.02	45378.95 ± 0.03	45378.934 ± 0.014
$a \sin i$ WR	[R <sub>⊙</sub> ]	17.5 ± 0.4	17.6 ± 0.6	17.5 ± 0.3
$a \sin i$ O	[R <sub>⊙</sub> ]	7.9 ± 0.3	±	7.9 ± 0.3
$\sigma$ WR	[km s <sup>-1</sup> ]	21	23	
$\sigma$ O	[km s <sup>-1</sup> ]	24		
$f(m)$ WR	[M <sub>⊙</sub> ]	3.0 ± 0.2	3.1 ± 0.3	3.07 ± 0.15
$f(m)$ O	[M <sub>⊙</sub> ]	0.29 ± 0.03		0.29 ± 0.03

g99 – Data from Gamen and Niemela (1999) and Niemela et al. (1985)

tp – This paper. Data from Leoncito.

Table 3.V: Circular orbital elements of WR 31.

<b>WR 47</b>		g99-[N IV 4058 + abs.]	tp-[N IV 4058]	Overall combined
P	[d]	6.2393 ± 0.0001	6.2393 ± 0.0001	6.23930 ± 0.00002
$\gamma$ WR	[km s <sup>-1</sup> ]	-79.0 ± 3.4	-59.8 ± 5.3	-
$\gamma$ O	[km s <sup>-1</sup> ]	-41 ± 7		-
$K$ WR	[km s <sup>-1</sup> ]	262.9 ± 4.7	261 ± 8	263 ± 4
$K$ O	[km s <sup>-1</sup> ]	-218 ± 9	±	-218 ± 9
$E_0$	[RJD]	43912.48 ± 0.02	43912.41 ± 0.03	43912.48 ± 0.03
$a \sin i$ WR	[R <sub>⊙</sub> ]	32.4 ± 0.6	32 ± 1	32.4 ± 0.5
$a \sin i$ O	[R <sub>⊙</sub> ]	27 ± 1		27 ± 1
$\sigma$ WR	[km s <sup>-1</sup> ]	36	27	
$\sigma$ O	[km s <sup>-1</sup> ]	66		
$f(m)$ WR	[M <sub>⊙</sub> ]	11.8 ± 0.6	11.5 ± 1.1	11.8 ± 0.5
$f(m)$ O	[M <sub>⊙</sub> ]	6.9 ± 0.9		6.9 ± 0.9

g99 – Data from Gamen and Niemela (1999), Niemela et al. (1980) and Mandrini (1983)

tp – This paper. Data from Leoncito.

Table 3.VI: Circular orbital elements of WR 47.

### 3.4.3 Orbit fit

The measured radial velocities and their errors of the 2009 campaign are presented in annex III. The typical errors ( $\sigma$ ) for each data set are deduced from the external dispersion of the differences (RV–Fit) between the measurements and the orbital solution. For the Leoncito data, the errors obtained are consistent with what one would expect from the data quality a priori (see Eq. 3.7 and discussion in section 3.7.1). We then fitted new circular orbits to the data (WR and O simultaneously), including data from the previous observational campaigns (g99, r96, n83); only the weights were recalculated with our method in order to insure homogeneity. The fitted parameters were the time of inferior conjunction of the WR star  $E_0$ , the semi-amplitudes  $K[\text{WR}]$  and  $K[\text{O}]$ , the systemic velocities  $\gamma_{\text{WR}}$  and  $\gamma_{\text{O}}$ , and the period. We used the IDL<sup>2</sup> routine MPFIT.PRO (Markwardt, 2009), which uses the Levenberg-Marquardt technique to solve the least-squares problem. Errors in the parameters were computed by MPFIT using the covariance matrix method applied to the RV errors given as input. We fitted for each set of data an independent  $\gamma$  parameter (=systemic velocity) to best reconnect the different data sets with each other. A weight was attributed to each set based on the external inverse mean variance of the measurements minus the model. The resulting fits are given in Fig. 3.2, and the fitted orbit parameters in Tables 3.II–3.VI. We tried to deduce the O star  $K$  parameters from the present data alone by performing a series of “shift-and-add” procedures (see section 3.5) using a varying  $K[\text{O}]$  parameter. We would then expect to get, for the right value of  $K[\text{O}]$ , a best visibility of the combined O star absorption lines (i.e. greatest depth and smallest width). Unfortunately, this technique failed and was abandoned.

For WR 12, WR 21, WR 31 and WR 47 we observe large differences (of order  $100 \text{ km s}^{-1}$ ) in  $\gamma(\text{WR})$  values between the data sets. These differences are most probably due to the different techniques of measurement used by the authors. This means that in spite of the rather small statistical errors, there is an additional systematic error of the order of  $\pm 100 \text{ km s}^{-1}$ , i.e. in absolute terms the  $\gamma$  values

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<sup>2</sup>Interactive Data Language, ITT Visual Information Solutions (ITT VIS)



are uncertain by this margin.

### 3.4.4 Lomb-Scargle

As an additional check, we performed a Lomb-Scargle analysis (Horne and Baliunas, 1986, Scargle, 1982) on the ensemble of the WR RV data, see Fig. 3.3 and Table 3.VII. Errors on the periods were computed via Monte-Carlo simulations. i.e. we re-shuffled the data points according to their respective errors and deduced the best periods for each (tirage).  $\sigma(P)$  was then computed as the standard deviation of  $P$  after about 200 simulations. The resulting periods are consistent with those obtained via orbital fitting. For WR 30, two periods are in fact possible due to aliasing, i.e. caused by the data sampling,  $18.80678 \pm 0.00004$  and  $18.84152 \pm 0.00003$  d but we arbitrarily adopted the first one in order to continue our analysis. Note however that this choice does not affect our conclusions about wind-wind collision effects.

## 3.5 Spectral Classification

Given the dominating strength of the WR spectrum, it is useful to apply a spectral decomposition technique to isolate the O-star spectrum. We therefore used the “shift-and-add” technique of Demers et al. (2002), with our new orbital solutions. In this procedure, we first co-add all the spectra in the WR frame of reference, by correcting each spectrum for its orbital velocity (taken from the fitted curve and not the individual measurements), which “dilutes” the O-star absorption

Star Name	Period (d.)
WR 12	$23.92323 \pm 0.00007$
WR 21	$8.25441 \pm 0.00003$
WR 30	$18.8068 \pm 0.0004$ or $18.8415 \pm 0.0004$
WR 31	$4.83066 \pm 0.00001$
WR 47	$6.23934 \pm 0.00002$

Table 3.VII: Periods obtained by the Lomb-Scargle analysis. Errors were estimated via Monte-Carlo simulations. .

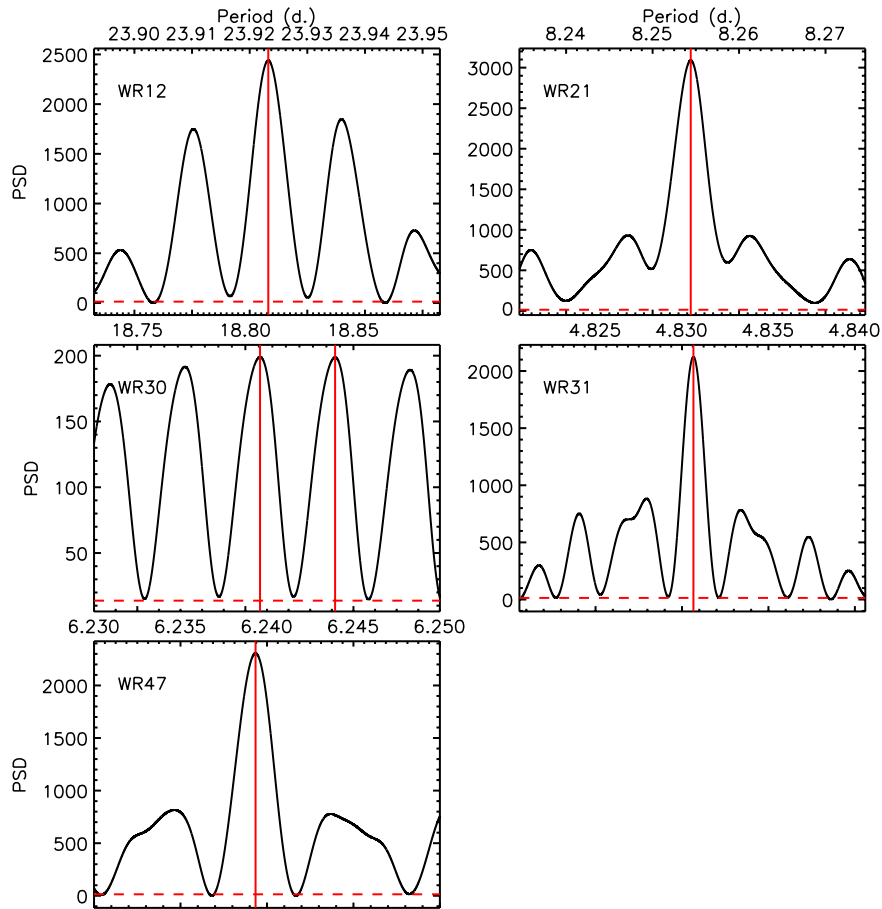


Figure 3.3: Power spectrum density (PSD) given by the Lomb-Scargle analysis applied on the WR radial velocities. The resulting periods are given in Table 3.VII. The dashed horizontal line represent the 99% significance threshold for period detection.

lines and leads to a first estimation of the WR spectrum. We then subtract this spectrum from all the original spectra and co-add the differences, this time in the O-star frame of reference, which leads to an estimation of the O spectrum. This spectrum is subtracted from the original spectra and the differences are co-added again in the WR frame of reference, and so on, until the co-added spectrum for each star converges, which generally happens in less than five iterations. The results of the spectral separations are shown in Fig. 3.4. We present in Tables 3.VIII and 3.IX the WR and O classification deduced from our spectra and the criteria from Smith et al. (1996) for the WN stars and Walborn and Fitzpatrick (1990) and Conti and Alschuler (1971) for the O stars. The classification of the WC star WR 30 (and WR 69 in the appendix) was not possible since the wavelength range does not cover the yellow region where the crucial lines for classification lie (Smith et al., 1990). We therefore adopted the previously published spectral types.

### 3.6 Temporal variance analysis

In systems such as colliding-wind binaries (CWB), we would expect to see mainly two sources of variability, one due to the displacement of the lines caused by the orbital motion, and another due to the varying excess emission produced by the colliding winds on the top of the most density-sensitive emission lines (mainly He II 4686 in WNs and C III 4650 in WCs for our spectra). In order to clearly characterize the variability in our spectra and unambiguously identify the colliding wind effects, we performed a temporal variance spectrum (TVS) analysis (Fullerton et al., 1996) on our data in the WR frame of reference. The TVS is given by the following equation :

$$TVS_j = \frac{1}{N-1} \sum_{i=1}^N \alpha_i d_{ij}^2 \quad (3.1)$$

where  $j$  designates the index of a pixel within a spectrum and  $i$  is the index of the spectra.  $d_{ij} = S_{ij} - \bar{s}_j$ , is the deviation from the mean spectrum at pixel  $j$  and  $\alpha_i = \sigma_{ic}^2 / \sum \sigma_{ic}^2$  is the normalized weight based on the continuum noise standard

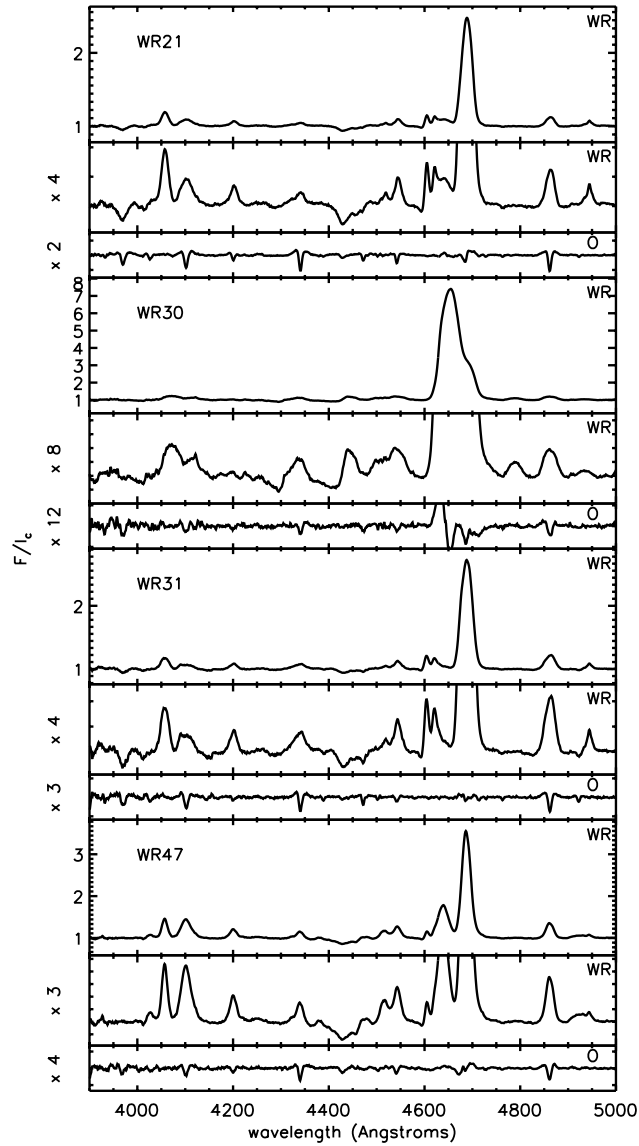


Figure 3.4: Results from the “shift-and-add” method for spectrum decomposition. For each star, we show a general view of the WR spectrum, an enlarged view of the latter, and the O star spectrum. The y-axis factors ( $\times 4$ ,  $\times 2$ , etc.) represent the y-scale enlargement relative to the above WR general view. Note that this technique does not yield absolute line strengths. This technique could not be applied on WR 12 since we do not have an orbit for the companion star.

Ion. sequence criteria		Strength-width criteria	Hydrogen criteria	Resulting sp. class
$\frac{N_{V\ 4603}}{N_{III\ 4641}}$	$\frac{N_{IV\ 4058}}{N_{V-III\ 4603-4641}}$	FWHM(4686) [ $\text{\AA}$ ]	$\frac{4340}{\sqrt{4200 \times 4541}} - 1$	
<b>WR 12</b>				
0.08	0.38	18.7±0.1	0.92±0.07	WN8 h
<b>WR 21</b>				
1.08	1.3	25.1±0.1	-0.4±0.08	WN5 o
<b>WR 31</b>				
5	0.85	26.9±0.1	-0.2±0.1	WN4 o
<b>WR 47</b>				
0.22	0.62	22.1±0.1	-0.36±0.04	WN6 o

Table 3.VIII: Classification deduced from our spectra and the criteria from Smith et al. (1996) for the WN stars. The classification of the WC star WR 30 was not possible since the wavelength range does not cover the yellow region where the crucial lines for classification lie.

Star	Signal/Noise	$EW(4471)$	$EW(4541)$	$\log EW(4471)/EW(4541)$	Sp. type
WR 21	1000	0.31±0.01	0.37±0.01	-0.08±0.02	O7
WR 30	250	0.31±0.01	0.31±0.01	0.00±0.02	O7.5
WR 31	500	0.29±0.02	0.21±0.02	0.14±0.05	O8
WR 47	500	0.14±0.02	0.33±0.02	-0.37±0.08	O5.5

Table 3.IX: Classification of the O star spectra obtained by the “shift and add” procedure and based on the criteria of Conti and Alschuler (1971) and Walborn and Fitzpatrick (1990).

deviation  $\sigma_{ic}$ . Another formulation of this TVS can also be used in order to take into account the variation of the Poisson noise throughout the spectrum, which can be critical in the strong emission lines of WR stars.

$$TVS_j = \frac{1}{N-1} \sum_{i=1}^N \alpha_i \beta_{ij} d_{ij}^2 \quad (3.2)$$

where  $\beta_{ij} = (\sigma_{ij}/\sigma_{ic})^{-2}$  in which  $\sigma_{ij}$  is computed for each spectrum in the continuum regions ( $\sigma_{ic}$ ) and scaled with the normalized flux throughout the spectrum according to the Poisson law, i.e.  $\sigma_{ij} = \sigma_{ic} \times \sqrt{S_{ij}}$ . We present in Fig. 3.1 the plots of the  $\sqrt{TVS}$  compared to the mean spectrum in the WR rest-frame and the significance threshold corresponding to a 99% confidence level for significant profile variability.

Most of the variability peaks (if not all) outside the “density-sensitive” lines (He II 4686 and C III 4650) can be attributed to the orbital motion of the O star absorption lines. To test this affirmation, we simulated a H $\gamma$  4340 gaussian line moving with the known orbit of WR 21 as seen in the WR frame of reference and with the same characteristics as H $\gamma$  in WR 21 (i.e. EW=0.8 Å, fwhm=6 Å and signal-to-noise of 100). We obtained a TVS profile nearly identical to that seen in WR 21 (see Fig. 3.5).

However, the variability of the “density-sensitive” lines cannot be explained by the companion orbital motion only. We simulated this case on the He II 4686 line in order to compare it with the WR 21 TVS profile. Given the spectral type found for WR 21, the He II 4686 absorption line is expected to be of intermediate EW between H $\gamma$  and He II 4542 or 4200 (Walborn and Fitzpatrick, 1990). We therefore superposed a He II 4686 absorption line of EW=0.6 Å and fwhm=6 Å over an emission line of EW=-40 Å and fwhm=26 Å, with a signal-to-noise ratio of 100. The result of the simulation and the comparison with the TVS profile obtained for WR 21 is shown in Fig. 3.5. This clearly shows that, given our assumption about the absorption line EW, the contribution of the O star orbital motion to the TVS of He II 4686 is nearly negligible. This analysis also applies

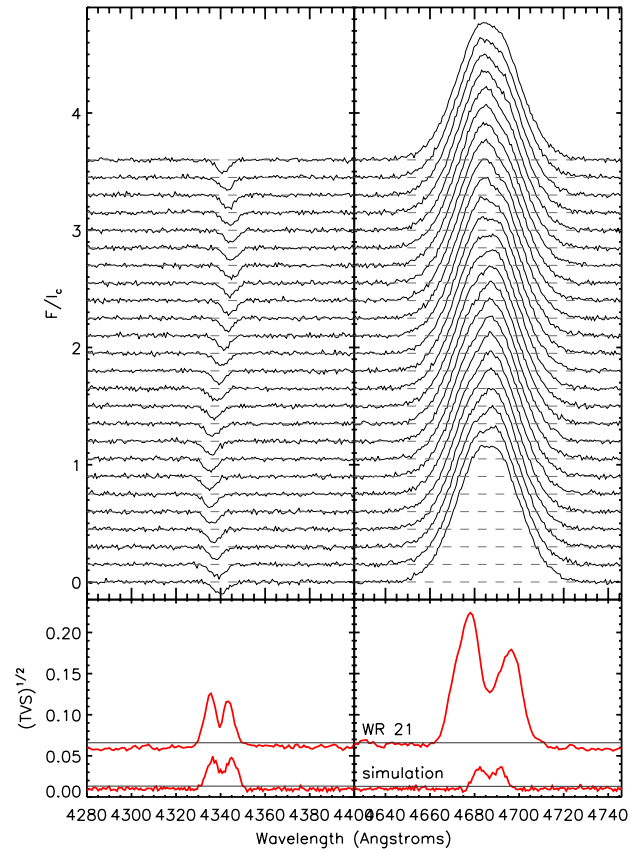


Figure 3.5: (left) Simulation of a  $H\gamma$  line moving with the orbit of WR 21 seen in the WR frame of reference and comparison of the resulting TVS with the one obtained for WR 21. (right) Simulation of a He II 4686 emission line + moving absorption line and comparison of the TVS with WR 21.

on the other stars of the campaign, since they all present even less prominent absorption components than WR 21.

### 3.7 Excess emission

#### 3.7.1 Measurement

The shock cone pushed by the stronger WR wind around the O-star produces excess emission that is responsible for most of the variability of He II 4686 and C III 4650. This excess emission arises from a thick slice of the shock surface (approximated by a cone) at a distance from the apex where the temperature of the plasma is appropriate for the He II and C III ions to exist. This excess emission is expected to have a maximum positive velocity (redshifted) near phase 0, when the WR star is in front and a maximum negative velocity (blueshifted) at phase 0.5, when the O star is in front. In order to isolate the emission due to the wind-wind collision from the underlying emission line produced by the WR expanding wind, we need to first have an estimation of the unperturbed reference profile of the emission line (assuming it is indeed unperturbed; previous experience shows this to be a fair assumption : Hill et al. 2000). Getting such an estimation is not straightforward, and comprises some degree of arbitrariness. We constructed these reference profiles by fitting simultaneously several gaussians to a minimum profile of the emission line using the deblending tool of the SPLIT routine in IRAF. We then adjusted the gaussians by eye so that the resulting excess looked reasonable. The chosen reference profiles compared with the ensemble of the spectra are shown in Fig.3.6 together with the resulting excess emission profiles as a function of the orbital phase. The gaussian parameters used for the reference profiles are presented in Table 3.X. This technique gave better results than simpler approaches like taking a profile built with the minimum flux of the set for each wavelength bin. This is due to the fact that we can more easily adjust the gaussians to the less perturbed wings of the lines. The chosen reference profiles compared with the ensemble of the spectra are shown in Fig.3.6 together with the resulting



excess emission profiles as a function of the orbital phase. The gaussian parameters used for the reference profiles are presented in the electronic version of this paper. Note that, except for the complex 4640 line of WR 30, the line cores are fitted with only one gaussian, the other components concerning the neighbouring lines, of which only the wings have influence.

After subtracting off the base profile, the position of the excess was computed with the first moment formula :

$$\bar{\lambda} = \int F(\lambda)\lambda d\lambda / \int F(\lambda) d\lambda, \quad (3.3)$$

where the integral is evaluated between  $\lambda_1$  and  $\lambda_2$ , the limits of the excess profile. Its width is calculated with the second central moment formula :

$$FW = 2\sigma = 2 \times \sqrt{\int F(\lambda)(\lambda - \bar{\lambda})^2 d\lambda / \int F(\lambda) d\lambda}. \quad (3.4)$$

Or, in their discretized version :

$$\bar{\lambda} = \sum \lambda_i F_i \Delta\lambda / \sum F_i \Delta\lambda = B/A \quad (3.5)$$

in which  $\Delta\lambda$  is the size of one pixel in  $\text{\AA}$ , and

$$FW = 2 \times \sqrt{\sum F_i (\lambda_i - \bar{\lambda})^2 \Delta\lambda / \sum F_i \Delta\lambda} = 2 \times \sqrt{\frac{C}{A}}, \quad (3.6)$$

where the sum  $\Sigma$  is from  $i=0$  to  $N-1$ , again corresponding to the extension of the excess. The errors on these measurements were estimated the following way. Let  $\sigma_{\text{pix}}$  be the flux error for one pixel, estimated for each individual excess profile. We have :

$$\sigma(\bar{\lambda}) = \sqrt{\left(\frac{\sigma(B)}{A}\right)^2 + \left(\frac{B\sigma(A)}{A^2}\right)^2} \quad (3.7)$$

and

$$\sigma(FW) = \frac{\sigma(C/A)}{\sqrt{C/A}} = \sqrt{\frac{A}{C}} \times \sqrt{\left(\frac{\sigma(C)}{A}\right)^2 + \left(\frac{C\sigma(A)}{A^2}\right)^2}. \quad (3.8)$$

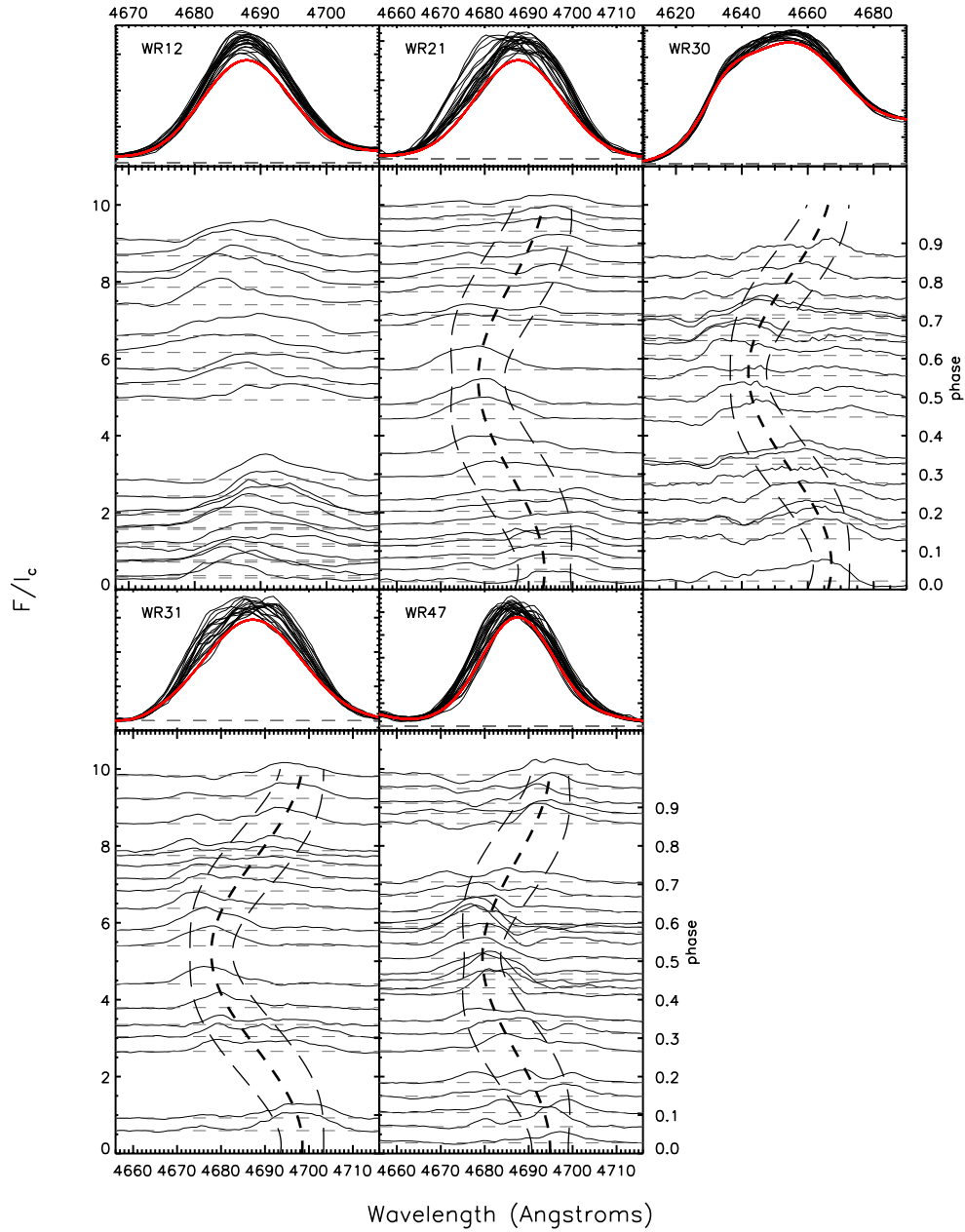


Figure 3.6: Reference profiles compared with the ensemble of the spectra and the resulting excess emission profiles as a function of phase. We also show the excess phase position (short dashed line) and width (horizontally between long dashed lines) arising from the Lührs model (see Fig 3.8 and Table 3.XI).

Star	$\bar{\lambda}$ (Å)	Flux	$\sigma$ (Å)	Star	$\bar{\lambda}$ (Å)	Flux	$\sigma$ (Å)
WR 12	4599.42	-0.90	8.29	WR 31	4596.23	-0.69	8.88
	4605.52	2.30	14.00		4602.10	2.80	11.60
	4639.53	30.00	20.00		4619.60	1.20	9.5
	4661.34	3.70	18.00		4626.27	1.80	28.07
	4687.72	47.57	16.00		4660.00	-0.45	8.89
	4709.50	8.50	26.88		4687.12	41.90	26.13
WR 21	4603.36	1.30	9.00	WR 47	4593.71	-0.31	10.91
	4620.09	0.90	8.70		4604.51	1.24	8.17
	4639.00	3.00	27.00		4638.70	21.50	26.19
	4687.65	31.22	23.32		4687.34	43.35	19.00
					4693.46	11.51	41.01
WR 30	4626.68	-7.00	46.02				
	4635.50	17.00	16.43				
	4634.35	1.00	8.80				
	4654.83	205.36	42.00				
	4695.50	30.50	25.00				
	4720.00	2.00	40.00				

Table 3.X: Gaussian parameters used for the reference profiles shown in Fig. 3.6.  $\bar{\lambda}$  : mean position, Flux : integrated flux of the gaussian and  $\sigma$  : standard deviation.

And we also have :

$$\sigma(A) = \sigma_{\text{pix}} \Delta \lambda \sqrt{N} \quad (3.9)$$

$$\sigma(B) = \sqrt{\sum (\lambda_i \sigma_{\text{pix}} \Delta \lambda)^2} = \sigma_{\text{pix}} \Delta \lambda \sqrt{\sum \lambda_i^2} \quad (3.10)$$

$$\sigma(C) = \sqrt{\sum ((\lambda_i - \bar{\lambda})^2 \sigma_{\text{pix}} \Delta \lambda)^2} \quad (3.11)$$

$$= \sigma_{\text{pix}} \Delta \lambda \sqrt{\sum (\lambda_i - \bar{\lambda})^4} \quad (3.12)$$

### 3.7.2 Lührs model fit

We can analyse the excess emission by using a simple geometric model of the colliding-wind region (CWR) developed by Lührs (1997). The model takes into account the half-opening angle of the shock cone  $\theta$ , the velocity of the plasma along the cone  $v_{\text{strm}}$  and the orbital inclination  $i$ . The CWR is assumed to be curved in the plane of the orbit and its orientation over a limited range of the shock cone is parameterized by a single angle,  $\delta\phi$ , for simplicity (see Fig. 3.7). By a

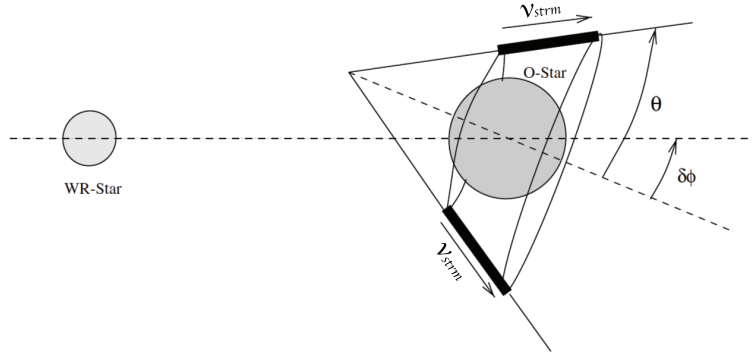


Figure 3.7: Schematic view of the geometric model by Lührs (1997) taken from Bartzakos et al. (2001) (their Fig. 2).

simple integration of the emissivity on a thick slice of the cone where the emission is supposed to arise, the model provides theoretical excess profiles showing double peaks moving with the orbit. However, depending on spectral resolution and the velocities and geometry of the shock cone, the double peaks might not be easily visible. Faced with this situation, one can still derive some valuable estimates of the colliding-wind region by characterizing the entire excess emission in a simple way. Hill et al. (2000, 2002) have shown that, for a circular orbit, the width (e.g. FWHM) and mean position of the excess emission vary as

$$FW_{\text{ex}} = C_1 + 2 v_{\text{strm}} \sin \theta \sqrt{1 - \sin^2 i \cos^2(\phi - \delta\phi)} \quad (3.13)$$

$$RV_{\text{ex}} = C_2 + v_{\text{strm}} \cos \theta \sin i \cos(\phi - \delta\phi) \quad (3.14)$$

where  $C_1$  and  $C_2$  are simple arbitrary constants (whose meaning is not perfectly clear but can be attributed, at least partly, to turbulence, as detailed in the next paragraph), and  $\phi$  is the orbital azimuthal angle (or true anomaly for  $\omega = 270^\circ$ ) in the case of a circular orbit :  $\phi = 360^\circ \times \text{phase}$  and the Wolf-Rayet star is in front at  $\phi = 0$ . Reasonable fits are obtained for a very wide range of parameters, but some limits can be derived by making a few realistic assumptions. For example, we can require that the velocity of the gas streaming along the bow shock cone,

$v_{\text{strm}}$ , does not exceed the terminal velocity of the WR wind,  $v_{\infty}$ . Hamann et al. (1995) give  $v_{\infty} = 1100 \text{ km s}^{-1}$  for WR 12. For the other stars, we took mean values for WN4-6 and WC6 stars based on van der Hucht (2001) (their Table 27). Another possible constraint can be obtained by assuming that turbulence in the shock cone contributes to the constant  $C_1$  in equation 3.13. Hill et al. (2002) found that values of the order of 200-500  $\text{km s}^{-1}$  were appropriate for WR 42, WR 48, and WR 79 in this respect. Based on this work, we chose to force the value of  $C_1$  between 0 and 500  $\text{km s}^{-1}$ .

The resulting fits are shown in Figs. 3.6 and 3.8. The fit of WR 12 did not lead to a reasonable result and we suggest that the wind-wind collision geometry might be particularly perturbed in this case, so that the Lührs model does not describe it well. We also observe that the time of maximum velocity of the excess does not coincide with phase 0 (see Fig. 3.6), which is not compatible with the expected behaviour of a wind-wind collision excess as described by Lührs. The perturbations of the shock geometry might arise from unstable flows like those described in Stevens et al. (1992) for highly radiative shocks.

### 3.7.3 Discussions and conclusions

We find that our values for the inclination of WR 31 and WR 47 are clearly compatible with previous estimations by Lamontagne and Moffat (1987) (photometry) and Moffat et al. (1990) (photometry and polarimetry) (see Table 3.XI). Given the large uncertainty obtained, WR 30 is marginally compatible with the value of Lamontagne et al. (1996) (photometry). One star, WR 21, is discordant with Lamontagne et al. (1996), giving a much larger inclination value. We therefore performed two additional fits of this star with  $i$  fixed to its previously published value. In the second fit, we also suppressed the constraint on  $C_1$ . We obtain larger values of  $\theta$  and  $v_{\text{strm}}$  in the first case (see Table 3.XI column (\*)), and even larger in the second one (column (\*\*)). However, the three values of  $\theta$  are all contained in an interval of  $10^\circ$  centered on  $i=40^\circ$ .

Several caveats of our method should be emphasized. First, it is very difficult

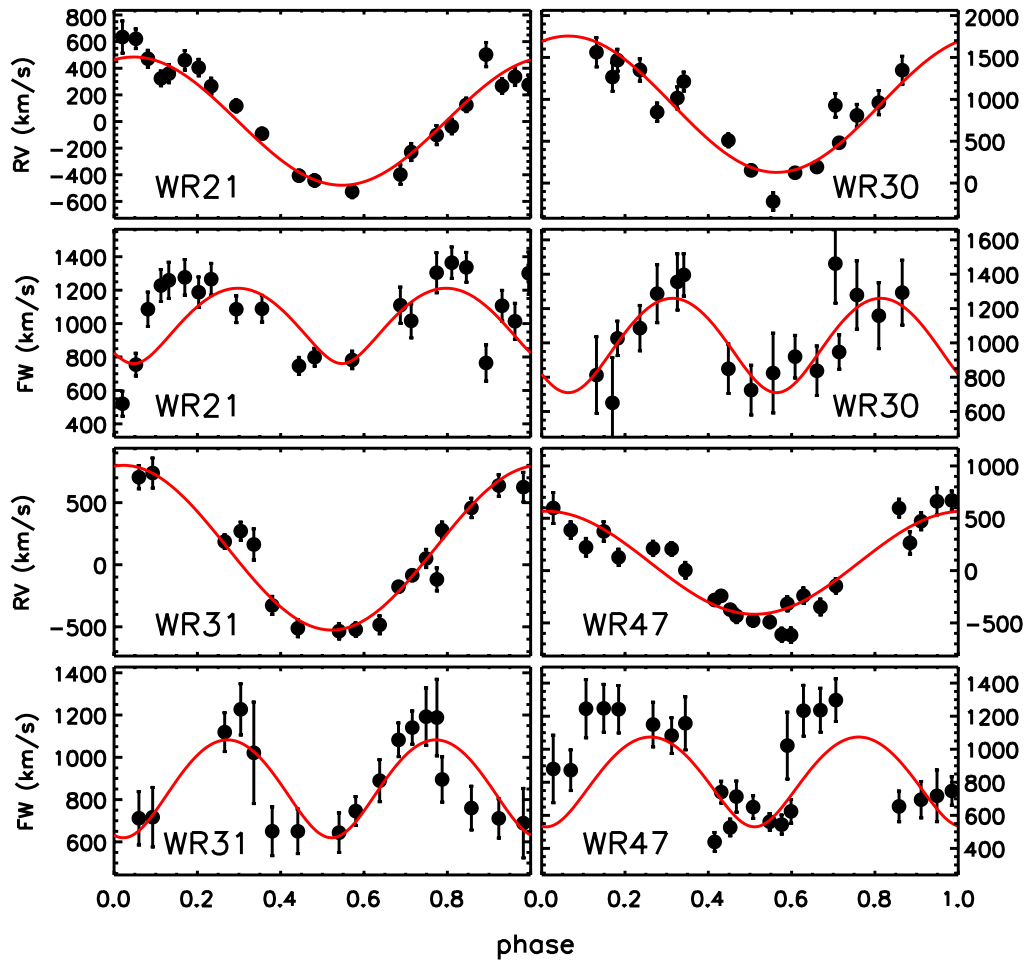


Figure 3.8: Resulting fit of the position and width of the excess emission with the Lührs model. The fitted parameters are shown in Table 3.XI

		WR 21	(*)	(**)
$C_1$	[km s <sup>-1</sup> ]	500	200	-30 ± 110
$C_2$	[km s <sup>-1</sup> ]	3 ± 2	1 ± 12	2 ± 12
$v_{\text{strm}}$	[km s <sup>-1</sup> ]	628 ± 18	800 ± 20	890 ± 50
$\theta$	[°]	34.5 ± 1.8	37 ± 1	45 ± 3
$\delta\phi$	[°]	17 ± 2	18 ± 2	17 ± 2
$i$	[°]	68.5 ± 3.3	49.6 (fixed)	49.6 (fixed)
Previously published $i$		49.6 ± 3.7 <sup>1</sup>	-	-
		WR 30	WR 31	WR 47
$C_1$	[km s <sup>-1</sup> ]	200	200	200
$C_2$	[km s <sup>-1</sup> ]	942 ± 32	137 ± 19	75 ± 16
$v_{\text{strm}}$	[km s <sup>-1</sup> ]	1068 ± 63	870 ± 40	782 ± 19
$\theta$	[°]	30 ± 3	30 ± 2	38 ± 2
$\delta\phi$	[°]	23 ± 2	8 ± 2	5 ± 2
$i$	[°]	61.2 ± 6.8	61.7 ± 4.5	67 ± 3
Previously published $i$		78.3 ± 5.8 <sup>1</sup>	61.6 ± 1.7 <sup>2</sup>	67.1 ± 1.9 <sup>3</sup>

References : <sup>1</sup> Lamontagne et al. (1996), <sup>2</sup> Lamontagne and Moffat (1987), <sup>3</sup> Moffat et al. (1990)

Table 3.XI: Resulting parameters of the Lührs model fit. For WR 21, two additional have been performed, one with inclination fixed to the published value from Lamontagne et al. (1996) (\*) and one with the constraint on variable  $C_1$  removed (\*\*).

to isolate the unperturbed profile with certainty in these gaussian shaped lines; in this respect, the Lührs method for determining  $i$  is more appropriate for flat top lines like C III 5696 in WR 140. Second, other physical phenomena can perturb the idealistic view of the Lührs model, e.g. clumps and turbulence, but also, in the case where the shock cone reaches the forming region of the same WR emission line. A lack of emission due to the hole thus formed has to be taken into account, which we did not do in this study. Furthermore, in the case of WR 30, the blending of C III 4640 with C IV 4648, which probably also shows intrinsic variability, leads to a more perturbed profile, hence more difficult to measure with accuracy. Finally, the continuum normalization is a very sensitive step since we are measuring a relatively small excess signal over a strong line. A tiny error on the continuum function will then be multiplied by the amplitude of the line. However, we expect that our method for continuum normalization (section 3.3.2) has minimized this effect. All these factors lead to rather “rugged” excess profiles

which explains the relatively large dispersion of the FW in Fig. 3.8. Meanwhile, Fig 3.6 shows that the resulting fits reproduce reasonably well the excess emission behaviour in time.

Stellar masses and orbital radii computed from our values of the inclination, together with wind-wind momentum ratios deduced from  $\theta$  and the approach of Cantó et al. (1996) (their equation 28) are presented in Table 3.XII.

In order to deduce values for the mass-loss and terminal velocities from  $\eta = [\dot{M}(\text{O})v_{\infty}(\text{O})]/[\dot{M}(\text{WR})v_{\infty}(\text{WR})]$  and to discuss the plausibility of the values, one has to adopt values for some of the variables.  $v_{\infty}$  has been measured for the WR component in one of our measured stars, WR 30 (Torres et al., 1986). For the other stars, we adopted mean values for the spectral class taken from van der Hucht (2001). For the O star components, WR 31 and WR 47 have been previously classified as class V (Smith et al., 1996). For WR 31 and WR 47, we considered the cases of class V and class I O-stars. We adopted O-star terminal velocities deduced from the models of Martins et al. (2005) and the mass loss rate predictions of Vink et al. (2001) at solar metallicities with  $v_{\infty} = 2.6 \times v_{\text{esc}}$ . Given these hypotheses, the WR star mass-loss rates resulting from our values of  $\eta$  are given in Table 3.XIII. Because of the large source of uncertainties, these values cannot be considered as very accurate and should only be used as rough guides. However, we have shown that the mass-loss rates obtained are realistic for Wolf-Rayet stars of these types.

### 3.8 APPENDIX : The case of WR 69 (HD 136488)

As noted in the introduction, WR 69 was included in the observation program based on its published photometric period of  $P=2.293\pm 0.005$  d (Marchenko et al., 1998b) and its classification among the dust-forming WC stars (Williams and van der Hucht, 2000), suggesting it to be potentially a short-period colliding-wind binary. The period was revealed by a photometric follow-up over three years with the HIPPARCOS satellite and the 100-point lightcurve shows an amplitude of 0.044 mag with a dispersion of 0.032 mag. Williams and van der Hucht (2000)



		WR 21	(*)	(**)
$M_{WR}$	$[M_{\odot}]$	$13.44 \pm 0.64$	$24.5 \pm 3.6$	-
$M_O$	$[M_{\odot}]$	$24.34 \pm 1.00$	$44.4 \pm 6.4$	-
$\eta$		$0.026 \pm 0.005$	$0.033 \pm 0.003$	$0.063 \pm 0.015$
$a$	[A.U.]	$0.2681 \pm 0.0061$	$0.327 \pm 0.018$	-
		WR 30	WR 31	WR 47
$M_{WR}$	$[M_{\odot}]$	$18.9 \pm 5.4$	$4.32 \pm 0.40$	$42.6 \pm 3.6$
$M_O$	$[M_{\odot}]$	$39.4 \pm 6.0$	$9.52 \pm 0.76$	$51.0 \pm 2.8$
$\eta$		$0.017 \pm 0.005$	$0.017 \pm 0.004$	$0.036 \pm 0.006$
$a$	[A.U.]	$0.536 \pm 0.035$	$0.1342 \pm 0.0057$	$0.3010 \pm 0.0067$

Table 3.XII: Stellar masses and wind-wind momentum ratios ( $\eta$ ) computed from the Lührs-deduced inclinations and the algebraic solution of the two-wind interaction problem by Cantó et al. (1996). (\*) and (\*\*) refers to the different fits presented in Table 3.XI.

		WR 21	WR 31	
Adopted Spectral class		WN5o+O7V	WN4o+O8(V)	WN4o+O8(I)
$\eta$		$0.026 \pm 0.005$	$0.017 \pm 0.004$	-
$v_{\infty}(O)$	$[\text{km s}^{-1}]$	2527	2454	1762
$\dot{M}(O)$	$[M_{\odot} \text{ yr}^{-1}]$	$2.4 \times 10^{-7}$	$9.6 \times 10^{-8}$	$1.3 \times 10^{-6}$
$v_{\infty}(WR)$	$[\text{km s}^{-1}]$	1500	1800	-
$\dot{M}(WR)$	$[M_{\odot} \text{ yr}^{-1}]$	$(1.6 \pm 0.3) \times 10^{-5}$	$(8 \pm 2) \times 10^{-6}$	$(8 \pm 2) \times 10^{-5}$
		WR 30	WR 47	
Adopted Spectral class		WC6+O7.5V	WN6o+O5(V)	WN6o+O5(I)
$\eta$		$0.017 \pm 0.005$	$0.036 \pm 0.006$	-
$v_{\infty}(O)$	$[\text{km s}^{-1}]$	2493	2579	2025
$\dot{M}(O)$	$[M_{\odot} \text{ yr}^{-1}]$	$1.5 \times 10^{-7}$	$1.5 \times 10^{-6}$	$5.7 \times 10^{-6}$
$v_{\infty}(WR)$	$[\text{km s}^{-1}]$	2200	1800	-
$\dot{M}(WR)$	$[M_{\odot} \text{ yr}^{-1}]$	$(1 \pm 0.3) \times 10^{-5}$	$(6 \pm 1) \times 10^{-5}$	$(1.8 \pm 0.3) \times 10^{-4}$

Table 3.XIII: WR mass loss rates deduced from our values of  $\eta$ . WR terminal velocities were taken from van der Hucht (2001). O-star mass loss rates and  $v_{\infty}$  were deduced from the models of Martins et al. (2005) and the mass-loss predictions of Vink et al. (2001) with  $v_{\infty} = 2.6 \times v_{\text{esc}}$ . For WR 31 and WR 47 O-star components, we considered the two cases : class V and class I luminosity stars. We used an error of  $100 \text{ km s}^{-1}$  on the terminal velocities and of about 10 % on the O-star mass loss rate in order to compute the errors on  $\dot{M}(WR)$ .

also clearly show the presence of an absorption component in the visible spectrum, identified as an OB star. The star, though, had never been thoroughly followed in spectroscopy. We show that the period is not due to binarity but WR 69 reveals a probable period of  $P=2.15\pm 0.04$  that might be compatible with that from Marchenko et al. (1998b).

### 3.8.1 Observations and data reduction

We collected 22 spectra of signal-to-noise of about 100. The observations and data reduction were made the exact same way as the other program stars.

### 3.8.2 Radial velocities

Radial velocities were measured by cross correlation of our spectra with a reference spectrum between wavelength 4000 and 5200 Å using the IRAF routine FXCOR. We detected a radial velocity trend of amplitude  $\simeq 30 \text{ km s}^{-1}$  with an external dispersion of about  $20 \text{ km s}^{-1}$  (Fig. 3.9). This is consistent with an orbit of period much longer than a month. However, we are unable to provide a more precise value in this paper and a longer time-baseline follow-up would be needed for this purpose. In any way, the published period of  $2.293\pm 0.005 \text{ d}$  is clearly not that of the orbital motion (Fig. 3.9).

### 3.8.3 Temporal variance analysis

In order to explore the intrinsic variability and not that from the observed velocity drift, we fitted a straight line through the measured RVs and used this law to doppler-shift the spectra in the WR frame of reference. We then performed a TVS analysis which revealed a significant variability in the the P-Cygni profile of C III 4650 (see Fig. 3.10). Lower levels of variability are also detected in lines  $\lambda 4257$ ,  $\lambda 4325$ ,  $\lambda 4475$ ,  $\lambda 4530$  and  $\lambda 4865 \text{ Å}$ , but the low amplitude signals could not be analysed in a satisfactory manner given our relatively low signal-to-noise.

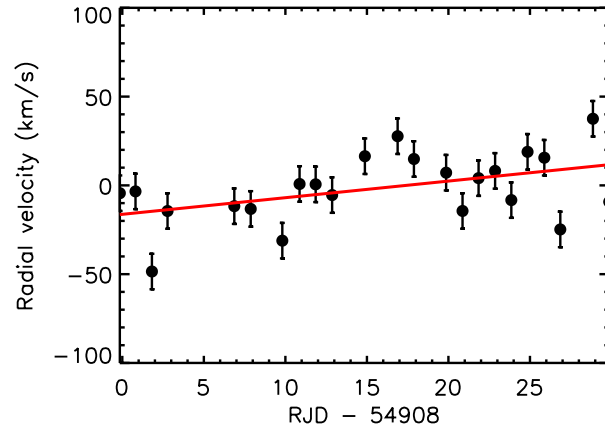


Figure 3.9: Measured radial velocities of WR 69 and the linear function fitted.

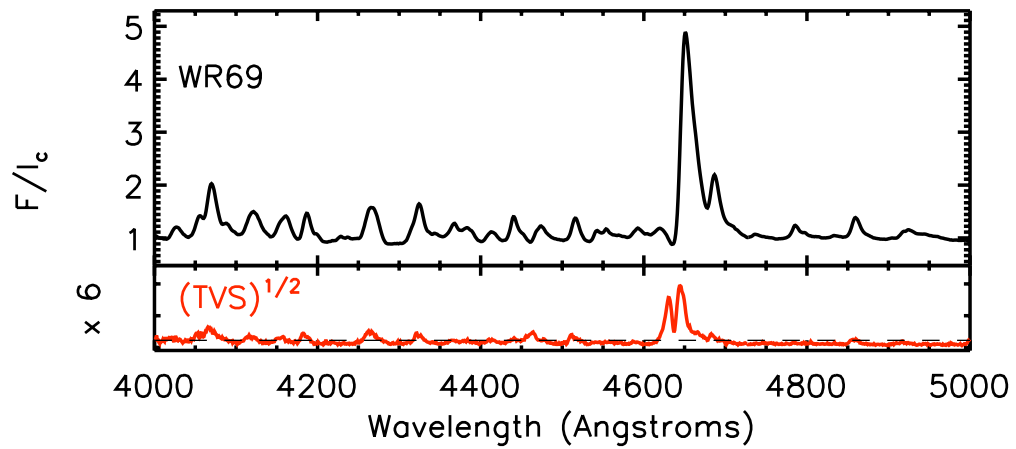


Figure 3.10: TVS analysis for WR 69. The ‘ $\times 6$ ’ factor represents the enlargement factor of the TVS y-axis relative to the above spectrum curve. The dashed line represents the 99% confidence level for significant variability.

### 3.8.4 C III 4650 P-Cygni profile variation

In order to carry out a period-search, we computed the mean position, width and equivalent width of the P-Cygni absorption profile of C III 4650 and their errors with the same equations as in section 3.7.1. We choose the continuum level and left boundary of the integral to be the first maximum on the left of the absorption profile. The right boundary was then defined as the first point on the right side of the profile to reach the continuum level. Thereafter, we performed a Lomb-Scargle analysis on each of the time series thus obtained, looking for periods between 1 and 10 days. The result of this analysis is presented in Fig. 3.11. The most probable periods for each quantity and their errors are given in Table 3.XIV. In Fig. 3.12, we show the spectra as a function of phase for  $P=2.15$ . A Lomb-Scargle analysis on lines  $\lambda 4257$  and  $\lambda 4325 \text{ \AA}$  shows that the period is possibly seen in these lines, but this is highly uncertain, given the low level of variability.

### 3.8.5 Discussion

Even if the Lomb-Scargle peak at  $P \simeq 2.25$  d seems clear and stable for the three quantities : RV, FW and EW, the plot of the 4640 line as a function of phase (Fig. 3.12) is not so convincing. Indeed, the profiles do not superpose well from cycle to cycle and some clear outliers are observed. However, the consistence of our period with the photometry of Marchenko et al. (1998b) is encouraging.

The behaviour of the variability of 4640 doesn't seem compatible with co-rotating interacting regions (CIR) models (Mullan, 1986, Prinja and Howarth, 1986) as we observe that when the absorption profile becomes wider, it also be-

Quantity	Period (d.)
Position	$2.13 \pm 0.04$
Width	$2.16 \pm 0.04$
Equivalent Width	$2.15 \pm 0.04$

Table 3.XIV: Periods obtained by the Lomb-Scargle analysis on the P-Cygni absorption part of C III 4640, respectively its position, width and equivalent width. Errors were estimated via Monte-Carlo simulations. .

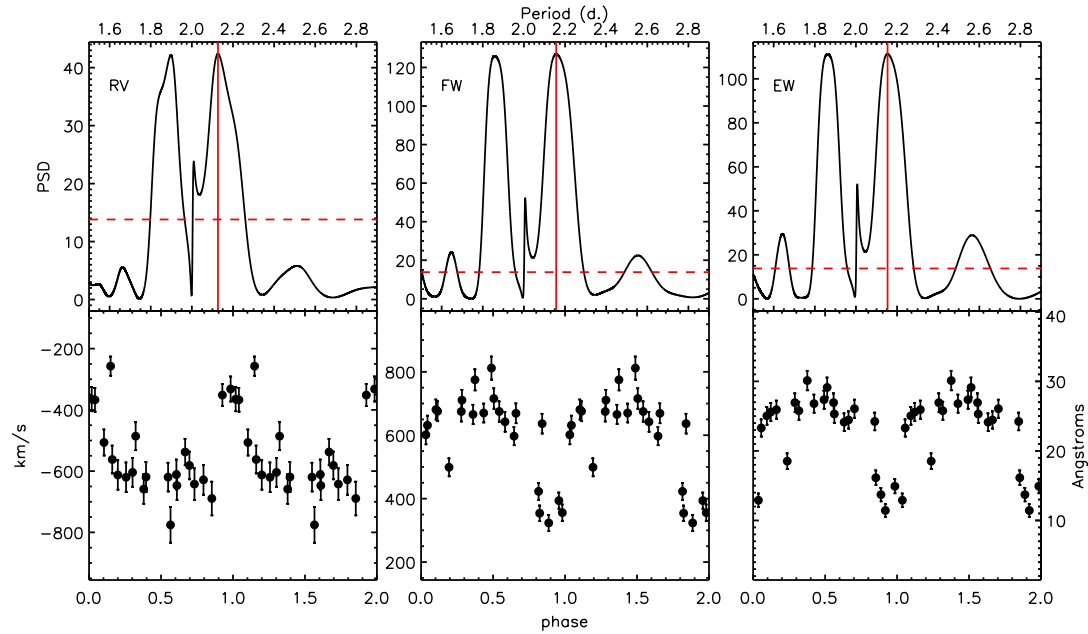


Figure 3.11: Power spectrum density (PSD) given by the Lomb-Scargle analysis applied on the position (RV), width (FW) and equivalent width (EW) of the absorption part of the P-Cygni profile of the C III 4640 line and the corresponding plots folded in phase. As we don't have any information about the time zero-point for the phase calculation, we chose an arbitrary value of  $T_0 = 0$  (reduced julian day). The resulting periods and their errors are given in Table 3.XIV. The dashed horizontal line represents the 99% significance threshold for period detection. The PSD peak symmetric to  $P \simeq 2.15$  about  $P = 2$  d is an alias of the method due to the regular sample rate of 1 day.

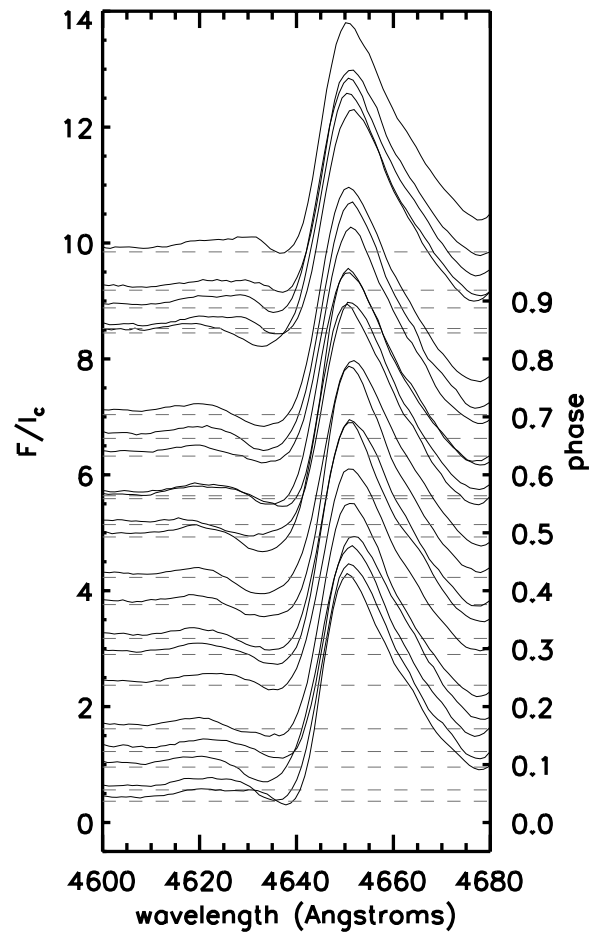


Figure 3.12: The C III 4640 line of WR 69 as a function of phase for P=2.15 d

comes blue-shifted. However, we cannot exclude a rotation-linked phenomenon, since there are no significant associated RV variations as expected from such a short-period binary. If we suppose a radius of  $8.4 R_{\odot}$  and a mass of  $10 M_{\odot}$  for the WC9d star (A. Sander & W.-R. Hamann, in prep.), the rotation period of 2.15 days would correspond to about 40 % of the break-up velocity which would make WR 69 a relatively rapid rotator. The rotation origin of the variability would though demand that the rotation period is seen in other parts of the spectrum and is stable in time. The first point might be true in lines  $\lambda 4257$  and  $\lambda 4325 \text{ \AA}$  but exploration of other spectral domains is needed. This would warrant future spectroscopic campaigns on this star. In any case, other physical phenomena (turbulence, instabilities, etc.) visibly perturb the ideal view of rotation-modulated variability.

### Acknowledgements

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## CHAPTER 4

### PHOTOMETRIC VARIABILITY OF WC9 STARS

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## ABSTRACT

Do some Wolf-Rayet stars owe their strong winds to something else besides radiation pressure? The answer to this question is still not entirely obvious, especially in certain Wolf-Rayet subclasses, mainly WN8 and WC9. Both of these types of Wolf-Rayet stars are thought to be highly variable, as suggested by observations, possibly due to pulsations. However, only the WN8 stars have so far been vigorously and systematically investigated for variability. We present here the results of a systematic survey during 3 consecutive weeks of 19 Galactic WC9 stars and 1 WC8 star for photometric variability in two optical bands,  $V$  and  $I$ . Of particular interest are the correlated variations in brightness and colour index in the context of carbon-dust formation, which occurs frequently in WC9 and some WC8 stars. In the most variable case, WR 76, we used this information to derive a typical dust grain size of  $\sim 0.1 \mu\text{m}$ . However, most photometric variations occur at surprisingly low levels and in fact almost half of our sample shows no significant variability at all above the instrumental level ( $\sigma \sim 0.005 - 0.01 \text{ mag}$ ).

**Keywords :** stars: Wolf-Rayet – circumstellar matter – stars: variable: other – stars: carbon – stars: individual (WR 76)

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### 4.1 Introduction

Population I Wolf-Rayet (WR) stars are the descendants of the most massive stars on the main sequence, namely O stars. They present the signatures of strong, dense winds, in which typical terminal velocities are of order  $10^3 \text{ km s}^{-1}$ . WR spectra come in two sequences, those with strong lines of helium and nitrogen (WN subtypes) and those with strong lines of helium, carbon and oxygen (WC and WO). WC9 are the coolest stars among population I WC subtypes,

with spectroscopic temperatures of order 50000 K at the stellar (hydrostatic) radius, deduced from non-LTE, clumped, line-blanketed models (Barniske et al., 2006, Crowther et al., 2002, Dessart et al., 2000). The particularity of WC9 stars is their frequent very bright luminosity in the infrared that is associated with dust formation (Williams et al., 1987a). Some WC9 stars also show strong spectroscopic and photometric variability (Crowther, 1997, Eenens and Corral, 2003, Kato et al., 2002, Lépine, 1996). The brightest WC9 star, WR 103, was observed to show periodic behaviour, suggesting a possible binary (Isserstedt and Moffat, 1981, Moffat et al., 1986), although later more intensive monitoring (Veen et al., 1999) suggests that stellar pulsations are at play, as also confirmed by recent MOST satellite data (Moffat et al., 2008). A few other WC9 stars have been monitored in the optical by Veen (2000), revealing behaviour that reminds one of the all-variable WN8 stars, which typically show a level of photometric variability of  $\sigma \simeq 0.016$  to  $0.05$  mag in the visible (Antokhin et al., 1995, Lamontagne and Moffat, 1987, Lefèvre et al., 2005, Marchenko et al., 1998a). Photometric variability of other WC9 stars both in the infrared and in the visible can be explained by episodic or periodic dust formation (NIR emission and optical absorption) associated in most cases with binarity (Veen et al., 1998, Williams, 1996, Williams and van der Hucht, 2000, Williams et al., 2005).

The goal of this study is to provide a systematic, uniform survey of WC9-stars using precision optical photometry in an attempt to answer the questions raised by previous studies, namely: Do all WC9 stars show variability like that of WN8 stars? And : Is variability associated with dust formation or with pulsations or both?

## 4.2 Observations

We selected a complete sample of 19 WC9 stars and 1 WC8 star in the southern sky from the “VIIth Catalogue of Galactic Wolf-Rayet Stars” and its annex (van der Hucht, 2001, 2006) with *V*-band magnitude from 12 to 16, in order not to go below the minimum practical exposure time of the CCD and to stay below

a reasonable maximum exposure time of 600 seconds. Observations were made using the Swope 1-meter telescope at Las Campanas (Chile) between UT 2007 June 8 and July 7. We used the direct CCD camera SITe#3 and the two optical bands  $V$  and  $I$  in order to be sensitive to both magnitude and colour variations, which are of particular interest to constrain the local dust properties. Our field of view is  $8.5 \times 10$  arcminutes with a pixel size of  $0.435''$ . We chose net exposure times such that we obtained a good signal-to-noise ratio (at least  $\sim 300$ ) per observation with one or two observations per clear night for each star. Standard star fields were not observed, since we are only interested in differential photometry.

### 4.3 Data reduction

Extraction of stellar magnitudes from the images was made in the usual way, using nightly flat-field images in each band, nightly bias images and overscan pixels in each image. We then chose to perform aperture photometry on our images, which is more efficient and precise than the alternative point-spread extraction for well-isolated stars in the field. We first interpolated all the images to match a reference frame so that the target star always has the same pixel position. This was done with the ISIS routine `interp.csh` (Alard, 2000). We then chose an adequate number of reference stars ( $N_{\text{Ref}}$ ) for which we measured the instrumental flux  $F$ . To do so, we used the IDL function `APER` with a fixed aperture radius  $R = 6$  pixels ( $2.610''$ ) in  $I$  band and  $R = 8$  pixels ( $3.480''$ ) in  $V$  band, which correspond to  $2 \times \text{FWHM}$  for typical images, together with a sky annulus of radii  $[1.5 \times R, 2.5 \times R]$ . The instrumental flux is then computed as

$$F = \sum_{r < R} P(r) - \pi R^2 \times F_{\text{sky}}$$

where  $P(r)$  is the pixel value at radius  $r$  and  $F_{\text{sky}}$  is the mean flux computed in the sky annulus. The differential magnitude  $\Delta M_i$  of each star in the field is finally

given by :

$$\Delta M_i = -2.5 \log \left( \frac{F_i}{\sum_{j \neq i}^{N_{\text{Ref}}} F_j} \right)$$

We then sorted the stars in order of variance and chose the  $N_{\text{Ref}}$  least variable among them to be the new reference stars. We then re-performed the process until it converged.  $N_{\text{Ref}}$  is chosen in order to minimize the variance of the reference stars and ranges from 10 for a very diffuse field to 50 or more for a crowded field. We defined the photometric precision (and the characteristic threshold for variability) as the dispersion of a reference star of similar magnitude than the WR. It goes from  $\sigma(\Delta M) = 0.005$  to 0.012 magnitude depending on the quality of the images, essentially determined by the average airmass at which the star could be observed.

#### 4.4 Results

We found 12 WC9 stars to be variable when compared to reference stars of similar magnitude : WR 76, WR 48b, WR 59, WR 75c, WR 81, WR 88, WR 121, WR 77t, WR 80, WR 92, WR 96 and WR 119 (Fig.4.1 and 4.2). Their variability ranges from  $\sigma = 0.012$  to 0.06 mag, the last value referring to the by far most variable star of the sample, WR 76. Aside from this star, the level of variability of our sample is quite small (going from  $\sigma$  0.012 to 0.02 mag) relative to the variable star WR 103 described by Veen et al. (1999), who confirm a value of  $\sigma(\Delta M)$  of  $\simeq 0.035$  mag in  $V$  band (0.05 mag peak to valley) initially found by Isserstedt and Moffat (1981) and confirmed by Moffat et al. (1986) and Balona et al. (1989). The timescale of the (stochastic) variability of WR 76 lies in the range of  $\sim$  a day to several days, like that seen in WR 103.

Eight stars were not detected as variables above the instrumental value (Fig.4.3 and 4.4). Among these stars we estimated the upper bound for  $\sigma_V$  to be around 0.01 for WR 106, WR 95 and WR 65 and around 0.005 for WR 53, WR 73, WR 75a, WR 75b and WR 117.

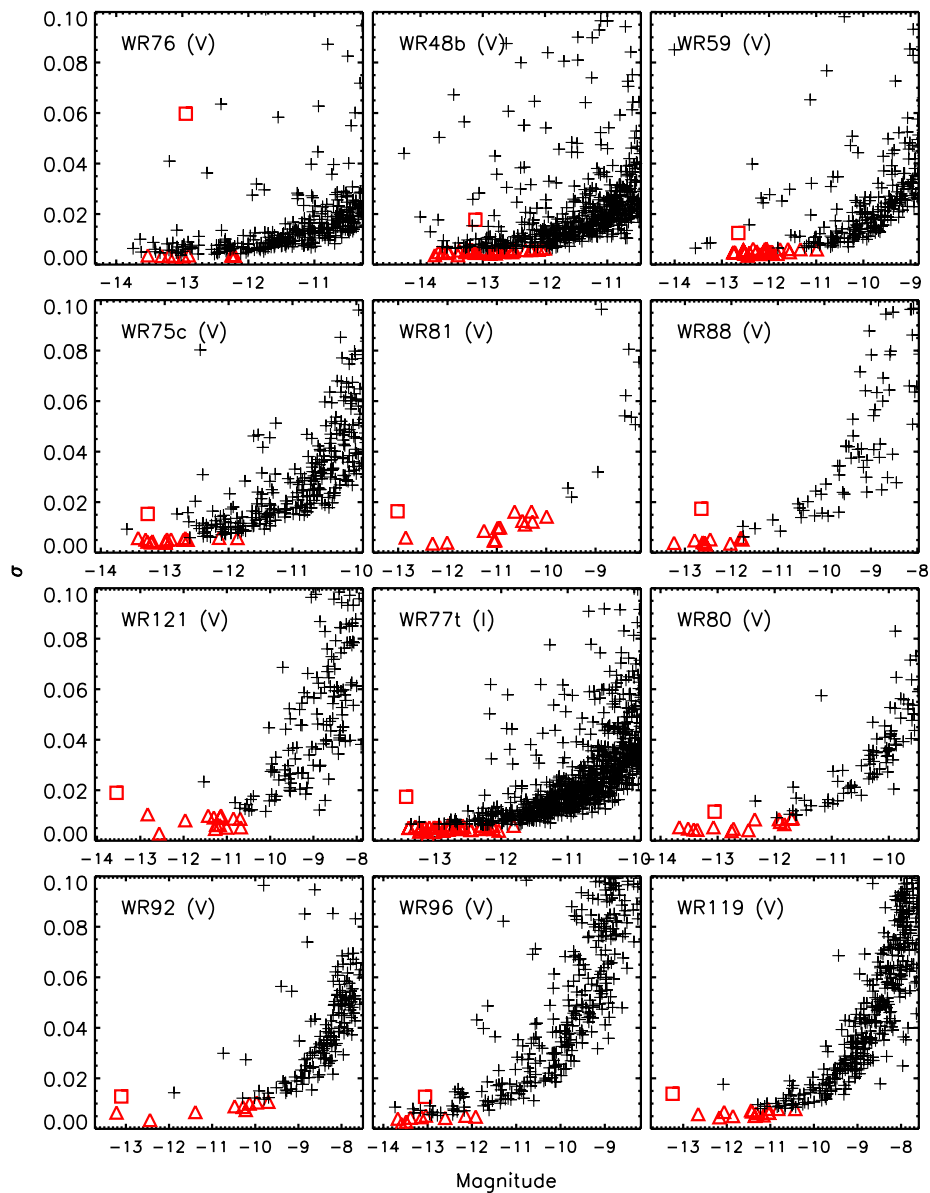


Figure 4.1: Dispersion plots for the fields containing the 12 variable WC9 stars: WR 76, WR 48b, WR 59, WR 75c, WR 81, WR 88, WR 121, WR 77t, WR 80, WR 92, WR 96 and WR 119. These are for the  $V$  band except for WR 77t, which has been observed only in  $I$  band. The reference stars are indicated by triangles and the Wolf-Rayet by a square. On the x-axis is the instrumental magnitude of the stars measured with arbitrary zero point from an image chosen as reference. On the y-axis is the standard deviation in magnitudes of  $\Delta M$  with time ( $\sigma$ ).

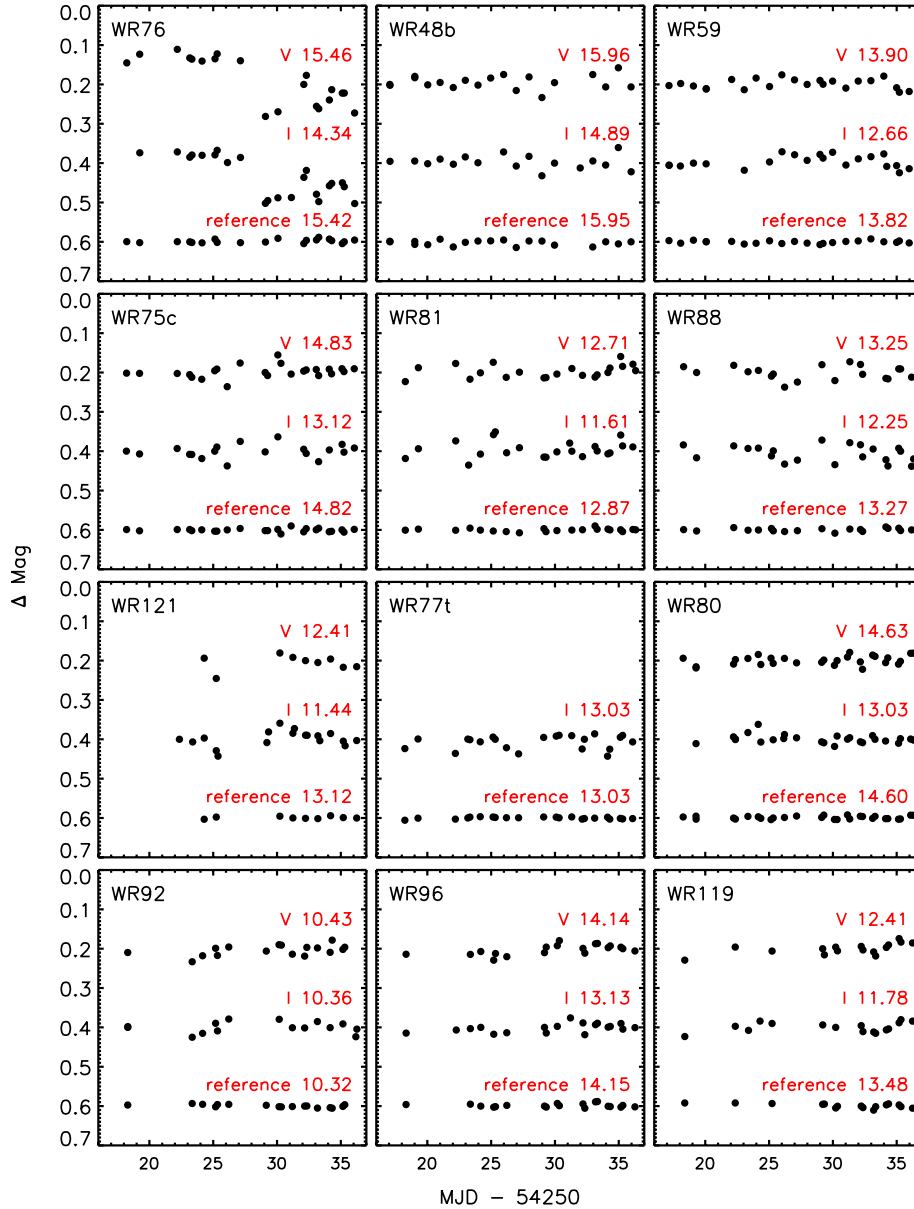


Figure 4.2:  $V$  and  $I$  lightcurves for the 12 variable WC9 stars compared with a reference star of approximately the same magnitude in  $V$ .  $\sigma$  is the rms standard deviation of  $\Delta Mag$  in time.  $V$  and  $I$  apparent magnitudes of the stars from van der Hucht (2001, 2006) are indicated. When not available, these apparent magnitudes are estimated values deduced from an extrapolation of what was available in the catalogue. MJD is for modified Julian day.



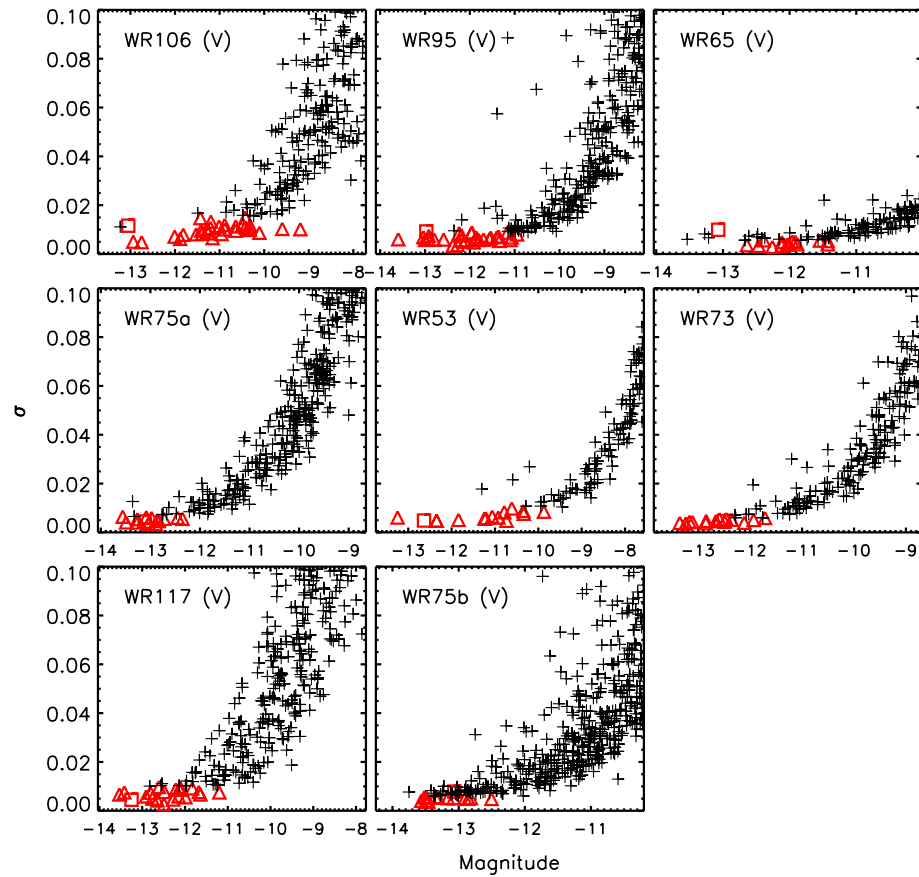


Figure 4.3: Dispersion plots for the fields containing the 7 non-variable WC9 stars WR 106, WR 95, WR 65, WR 73, WR 75a, WR 75b and WR 117, and the non-variable WC8 star WR 53 as in Fig.4. These are all for the V band.

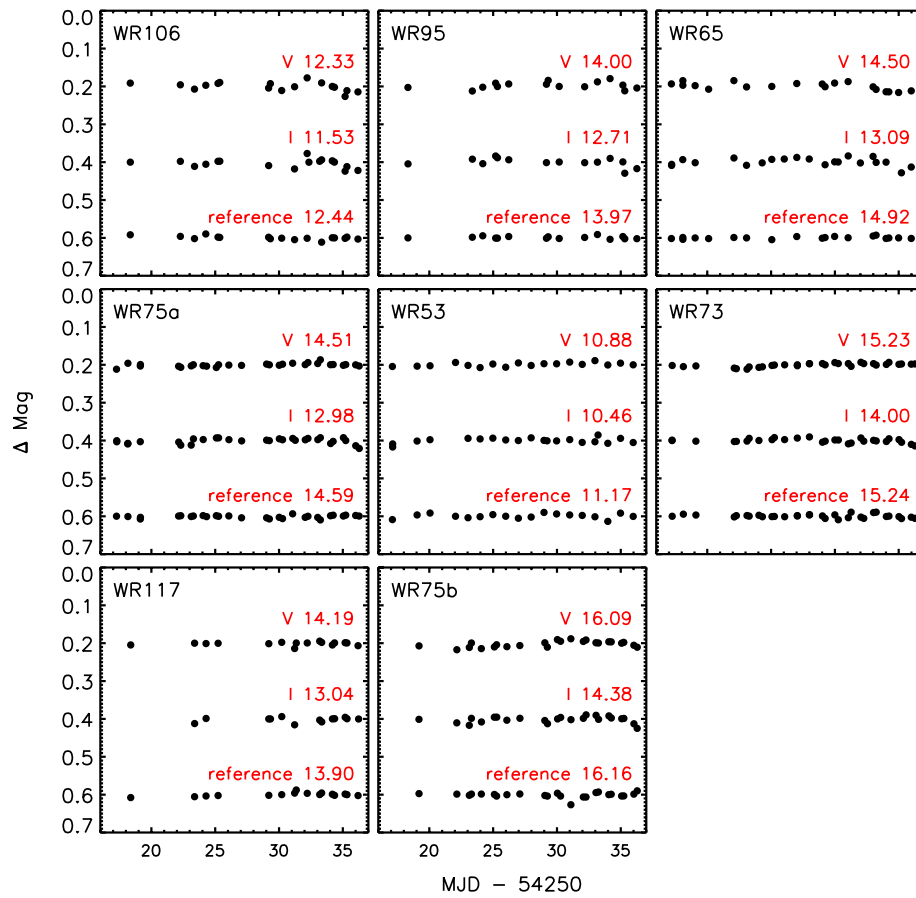


Figure 4.4:  $V$  and  $I$  lightcurves for the 8 non-variable WC9 stars compared with a reference star of approximately the same magnitude in  $V$ .

Star name	Spectral Type	$\sigma_V$ of WR	$\sigma_V$ of typical ref. star	$\sigma_I$ of WR	$\sigma_I$ of typical ref. star	Slope
WR 76	WC9d	0.060	0.004	0.051	0.006	4.47±0.60
WR 48b	WC9d	0.018	0.005	0.017	0.005	-
WR 59	WC9d	0.012	0.005	0.017	0.007	-
WR 75c	WC9	0.015	0.005	0.017	0.005	-
WR 81	WC9	0.016	0.010	0.021	0.010	-
WR 88	WC9	0.017	0.005	0.021	0.006	-
WR 121	WC9d	0.019	0.008	0.020	0.009	-
WR 77t	WC9d	-	-	0.017	0.005	-
WR 80	WC9d	0.012	0.006	0.011	0.007	-
WR 92	WC9	0.013	0.008	0.014	0.007	-
WR 96	WC9d	0.013	0.004	0.011	0.007	-
WR 119	WC9d	0.014	0.006	0.012	0.005	-
WR 106	WC9d	0.012	0.010	0.012	0.008	-
WR 95	WC9d	0.009	0.007	0.012	0.005	-
WR 65	WC9	0.001	0.004	0.010	0.007	-
WR 75a	WC9	0.004	0.005	0.007	0.007	-
WR 53	WC8d	0.005	0.007	0.007	0.006	-
WR 73	WC9d	0.005	0.005	0.005	0.005	-
WR 117	WC9d	0.005	0.007	0.006	0.008	-
WR 75b	WC9	0.008	0.005	0.009	0.005	-

Table 4.I: Measured  $V$  and  $I$  variability compared to typical values for the reference stars. The latter is estimated from the rms mean of the reference stars  $\sigma$ . ‘Slope’ is the slope of the ( $V$ ) vs ( $V - I$ ) plot (Fig. 4.5), which was only significant for the most variable star of the survey, WR 76.

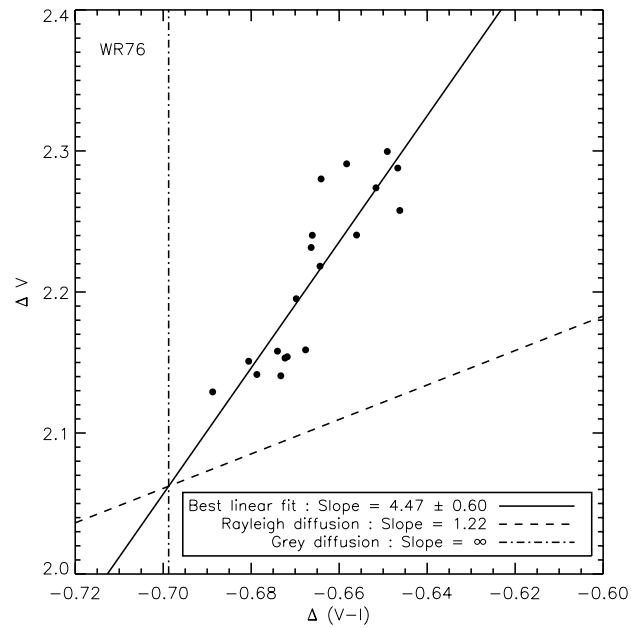


Figure 4.5:  $V$  band differential magnitude versus  $V - I$  differential colour index for WR 76, by far the most variable star of our survey. The expected curves for Rayleigh scattering (valid for dust grains smaller than  $\lambda_V/10 \simeq 0.05 \mu\text{m}$ ) and grey scattering (valid for dust grains larger than  $\lambda_V \simeq 0.5 \mu\text{m}$ ) are shown for comparison. The Rayleigh slope has been computed for Mie scattering with  $Q_{ext}(\lambda, a) \propto \lambda^{-4}$  (see text).

## 4.5 Dust formation

We searched for a correlation of  $V$  vs.  $(V - I)$  but only found one for WR 76, which is by far the most variable star of our survey (Fig.4.5). The computed correlation coefficient for this star is  $R=0.92$  and the Spearman rank index, which expresses the probability that the two variables are uncorrelated, is  $2.68 \times 10^{-8}$ . We observe that WR 76 becomes redder when the  $V$  extinction increases (the star gets fainter). This is consistent with circumstellar extinction by small dust grains, which is more efficient in  $V$  band than in  $I$  band. Moreover, WR 76 is known to have a very high dust formation rate measured from infrared photometry (Williams et al., 1987a): its dust formation rate is by far the highest among our sample. Nevertheless, the correlation between the dust formation rate and the level of variability of the star is completely dominated by this star, as seen in Fig.4.6. For the remaining stars, the correlation is quite poor. In particular, we found variable stars of low amplitude with no dust formation detected (WR 75c, WR 81, WR 88, WR 92) together with non variable stars known to produce dust because of their strong infrared excess (WR 53, WR 73, WR 117). This might be at least partly due to the shorter timescale of our variability search compared to that for typical dust ejection.

As some non-dusty WC9 stars show low-level variability, presumably some of that observed from WR 76 may also be intrinsic. However, on account of its high dust formation rate, variable extinction is thought to play a dominant role.

Estimations of the dust formation rates were not available for all our target stars, although the analysis of this correlation does show quite clearly that the production of dust cannot explain the behaviour of all WC9 stars. Like in the strongly variable WC9 star WR 103 (Moffat et al., 2008, Veen et al., 1999), we believe that pulsations are likely to play an important role in some cases, sometimes combined with the effects of dust formation.

## 4.6 Dust properties

It is known that dust in WC9 stars is essentially composed of amorphous carbon grains (Marchenko et al., 2002, Williams et al., 1987a, Zubko, 1998). We therefore used Mie theory (Bohren and Huffman, 1983) and a model of monosize carbon spheres, with optical constants from Zubko et al. (1996), in order to interpret the measured slope of Fig. 4.5 as a dominant size in the dust size distribution. To do so, we calculated the weighted extinction coefficients  $Q_{\text{ext}}(\lambda, a)$  (function of the wavelength  $\lambda$  and the dust sphere radius  $a$ ) as in Marchenko et al. (2003), the weights being given by  $W(\lambda) = T(\lambda)F(\lambda)$ , where  $T(\lambda)$  is the transparency of the corresponding filter ( $V$  or  $I$ ) and  $F(\lambda)$  is the energy distribution of WR 76 in the optical and the near infrared (Williams et al., 1987a). We then used the fact that the extinction,  $A(\lambda)$ , is directly proportional to  $Q_{\text{ext}}$ . Finally,  $A(\lambda)$  is linked to the variations of magnitude that we interpret as due to an increase of the extinction by the dust ( $A(\lambda) \propto \Delta Mag$ ). The result of our model is shown in Fig.4.7 for  $a$  ranging from 0.0001 to 0.2  $\mu\text{m}$ . Our estimated dust size, corresponding to a slope of  $4.47 \pm 0.60$ , is  $a = 0.126 \pm 0.003 \mu\text{m}$ . Note that the very small error might not be real but simply due to the over simplistic hypothesis of monosized grains, the reality being more a distribution of sizes. However, the parameter  $a$  gives an idea of the characteristic size of this distribution. Our value is quite close to the estimation of dust size in WR 140 (WC7pd+O5) during its periastron transit in 2001:  $a = 0.069_{-0.001}^{+0.002} \mu\text{m}$  (Marchenko et al., 2003) but is lower than other estimations in WR 112 (WC9d+OB?) and WR 118 (WC9d):  $a = 0.4 - 1.0 \mu\text{m}$  (Chiar and Tielens, 2001, Marchenko et al., 2002, Yudin et al., 2001). However, the last values concern very different systems from WR 140 and are based on different techniques from ours. A more homogeneous study would be needed to conclude about this point.

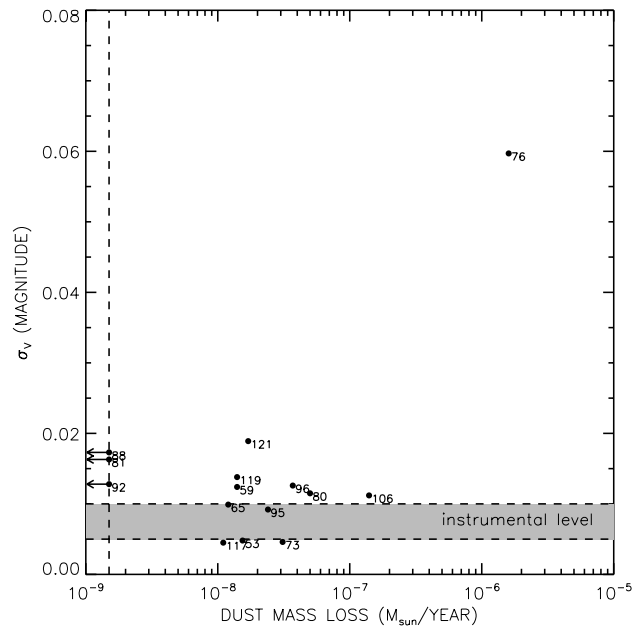


Figure 4.6: Our measured level of photometric variability  $\sigma(V)$  versus the estimated dust formation rate by Williams et al. (1987a). We show the range of values for the instrumental level, which is determined, for each field, by the  $\sigma$  of the reference star of closest magnitude to the WC9 star. The dust formation rate estimates are based on infrared photometry. We see that any correlation between the level of variability and the dust formation rate of the WC9 stars is completely dominated by only one star, WR 76, combined with a cloud of points at low variability level.

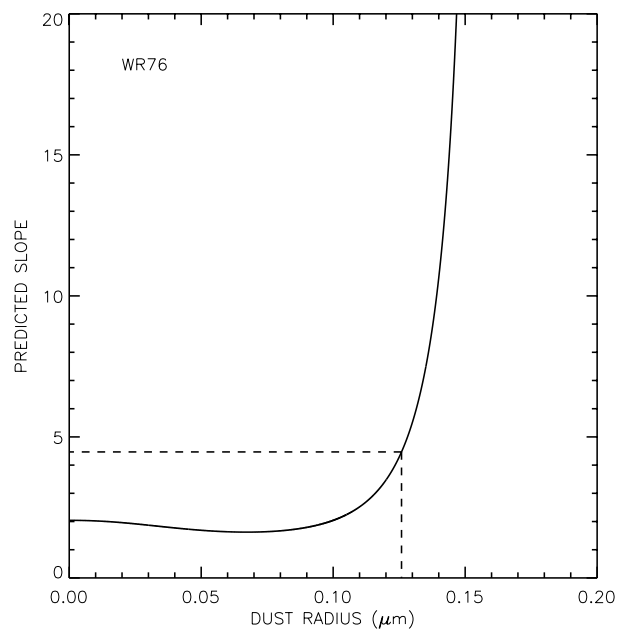


Figure 4.7: Model prediction of the  $V$  vs.  $(V - I)$  slope for WR 76. We used Mie theory for mono-sized spherical grains plus optical constants for amorphous carbon by Zubko et al. (1996). The best model to reproduce our slope of  $4.47 \pm 0.60$  is a dust size of  $a = 0.126 \pm 0.003 \mu\text{m}$



## 4.7 Conclusions

The most surprising result of this study is the large fraction of WC8/9 stars that are not significantly variable above the instrumental level (i.e.  $\sigma \sim 0.005$  to  $0.01$  mag). Even among the 12 that do vary, it is mostly at a relatively modest level,  $\sigma \lesssim 0.02$  mag, in contrast to their highly variable WN cousins, the WN8 stars. This shows that, contrary to the “violent” WN8 stars, pulsations are not present in WC8/9 stars, which argues against any important role for pulsations in wind driving or dust formation. We see a significant variation in both  $V$  and  $I$  filters for only one object, WR 76, which we ascribe to variable circumstellar dust extinction. We estimated the dominant dust grain size to be close to  $0.1 \mu\text{m}$ , which is  $\sim$  an order of magnitude smaller than estimated in two other WC9d stars.

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## CHAPITRE 5

### CONCLUSION

L'étude des systèmes binaires d'étoiles massives constitue une opportunité unique d'accès aux paramètres fondamentaux des étoiles massives, d'une part du fait de la binarité qui permet une mesure keplérienne des masses, et d'autre part, du fait du phénomène de collision des vents, qui permet une mesure des taux de perte de masse des étoiles. Ce dernier phénomène est responsable d'un excès d'émission en radio, IR (avec la formation de poussière de carbone), optique, rayons X, et même  $\gamma$ . Dans cette thèse nous nous sommes intéressés au domaine optique avec deux campagnes d'observation en spectroscopie et une en photométrie.

La première campagne a concerné l'objet archétype parmi les systèmes binaires à collision de vent : WR 140. Il s'agit d'un système WC7+O5.5fc de période  $P=7.93$  ans, de très grande excentricité ( $e=0.896$ ) et produisant de la poussière de manière périodique. Nous avons obtenu, grâce à une collaboration internationale entre professionnels et amateurs, une résolution temporelle inégalée pour ce système. Des paramètres orbitaux plus précis ont été calculés ainsi qu'une estimation de la géométrie de la zone de collision des vents grâce à une version modifiée du modèle géométrique de Lührs (1997). Ce modèle propose une vision géométriquement simplifiée d'un cône de choc se déplaçant au cours de l'orbite et dont l'intégration le long d'une tranche correspondant à la zone de formation de l'excès d'émission, fournit des profils théoriques et des prédictions sur la largeur et la position de l'excès. Ce modèle, conçu au départ pour une orbite circulaire a été adapté à une orbite elliptique telle que celle de WR 140. Une classification plus précise de l'étoile O a également été obtenue, même si la détermination exacte de la classe de luminosité n'a pas été possible.

La deuxième campagne a concerné 5 systèmes binaires de l'hémisphère sud, WR 12, WR 21, WR 30, WR 31 et WR 47. Un travail similaire à celui fait sur

WR 140 a été effectué. Là encore, une mise à jour des paramètres orbitaux a été obtenue, ainsi qu'une estimation de la géométrie de collision des vents à l'aide du modèle de Lührs. Une sixième étoile a également été observée, soupçonnée d'être un système binaire de courte période, il s'agit de WR 69 (WC9d). Cette étoile est bien une binaire mais nous avons établi que sa période orbitale est supérieure à un mois. Une période non-orbitale courte ( $P=2.15$  j) a cependant été détectée, compatible avec la période photométrique de Marchenko et al. (1998b), possiblement reliée à la rotation. Ce point reste cependant à confirmer dans une étude future.

Dans la troisième campagne, nous nous sommes intéressés à un phénomène associé à la collision de vents dans le cas où l'étoile WR est de type WC tardive, à savoir, la production de poussière de carbone amorphe. Le lien entre la formation de poussière et la collision des vents n'est en effet pas totalement clair. La poussière se forme-t-elle exclusivement au cours du processus de collision des vents, ou peut-elle se former dans des étoiles simples ? Cela impliquerait alors des chocs, probablement induits par des pulsations, telles que celles observées systématiquement dans les WN8. 19 étoiles WC9 et 1 WC8 ont donc été observées en photométrie pendant un mois, à la recherche de ces éventuelles pulsations. Le niveau de variabilité de ces étoiles s'est révélé étonnamment faible ( $\sigma \sim 0.005 - 0.01$  mag pour la majorité d'entre elles), ce qui suggère que les pulsations ne sont pas courantes chez les WC tardives, contrairement à ce que l'on observe chez leurs cousines riches en azote, les WN8. Le rôle des pulsations dans la formation de poussière est donc relativement modeste. Cela montre, par défaut, que la collision des vents en est le mécanisme prédominant.

## 5.1 Travail futur

La vision géométrique du modèle de Lührs est capable de fournir des informations fondamentales sur les étoiles massives constituant les systèmes binaires à collision de vents. Il est l'outil le mieux adapté pour obtenir ces informations d'une manière simple à l'aide de la spectroscopie optique. Cependant, cette vision

simple a ses limites. Premièrement, elle ne tient pas compte de l'éventuel défaut d'émission pouvant se produire sur la raie à modéliser si le cône de choc crée un trou dans la coquille d'émission de cette même raie associée au vent de la WR, comme cela est décrit dans l'introduction (section 1.2.4). Deuxièmement, il faut être prudent quant à l'interprétation de l'angle d'ouverture du choc, et notamment, sa traduction en terme de ratio des moments cinétiques des vents. En effet, l'angle d'ouverture du choc de la WR et celui de l'étoile O sont en général différents, et différents de celui de la surface de contact (voir Fig. 1.7). Nous avons fait l'hypothèse implicite que le cône mesuré est celui de la surface de contact, mais ce n'est pas nécessairement exact, en particulier si les chocs sont adiabatiques (Fig. 1.10). Il est nécessaire de construire des modèles plus élaborés incluant une description hydrodynamique et radiative de la zone de collision de vent, et étant capables de fournir des profils théoriques d'excès d'émission à comparer aux observations. Nous avons également constaté que l'extraction de l'excès d'émission dans le cas où la raie est de forme gaussienne est très incertaine, du fait de la difficulté à isoler le profil de base (non perturbé) avec certitude. Un modèle théorique fournissant des profils globaux "raie WR + excès d'émission" serait donc très utile dans ce contexte.

Nous avons apporté un indice supplémentaire appuyant le lien entre la formation de poussière circumstellaire et le phénomène de collision des vents. Cependant, la recherche systématique de compagnons dans les systèmes comprenant des WC tardives doit être poursuivie, et une compréhension théorique quantitative du mécanisme reste à être établie.



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## Annexe I

### Profil de raie produit par un anneau en expansion

Considérons un anneau en expansion avec une vitesse  $V_{exp}$  uniforme (Fig. I.1) émettant dans une raie de fréquence  $\nu_0$ . Supposons également que cet anneau est optiquement fin, qu'il a une émissivité linéique  $I_\theta$  constante et qu'il est vu par la tranche. Le profil de la raie, noté  $I_\nu$ , où  $\nu = c\Delta\lambda/\lambda_0$  correspond au décalage doppler, s'obtient en écrivant le changement de variable suivant :

$$I_\theta d\theta = I_\nu d\nu .$$

Puis, avec  $\nu = -V_{exp} \cos(\theta)$  :

$$d\nu = V_{exp} \sin(\theta) d\theta$$

et finalement

$$I_\nu = I_\theta \frac{d\theta}{d\nu} = \frac{I_\theta}{V_{exp} \sin(\theta)} = \frac{I_\theta / V_{exp}}{\sqrt{1 - (\nu/V_{exp})^2}} . \quad (\text{I.1})$$

Le profil de raie obtenu est un profil à double pic de maximums  $-V_{exp}$  et  $+V_{exp}$ . Il est intéressant de noter que l'on obtient exactement le même profil en considérant un coquille en rotation uniforme avec une vitesse  $V_{rot}$ ,  $V_{exp}$  étant alors simplement remplacé par  $V_{rot}$  dans I.1.

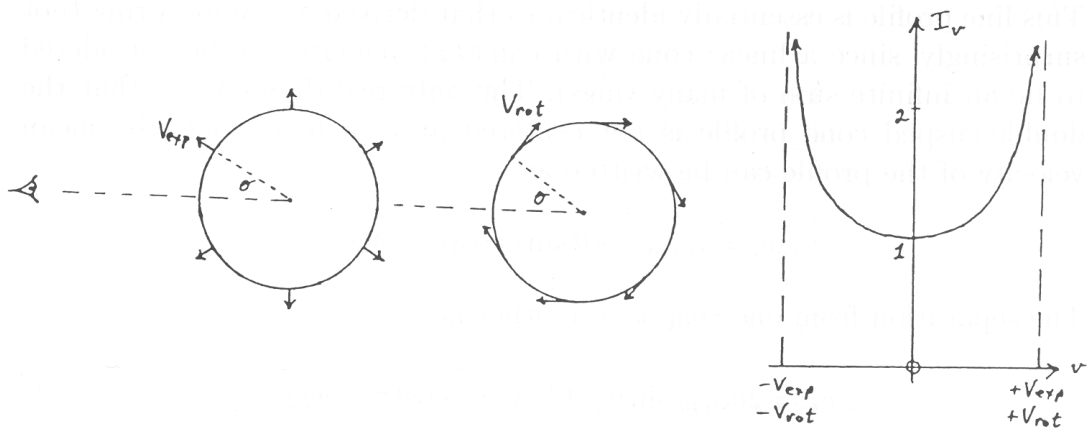


Figure I.1 – Schéma d'un anneau émetteur en expansion (ou rotation uniforme) et le profil à double pic correspondant. (Tiré de Moffat et al. (1998))

## Annexe II

### Mesures des vitesses radiales et de l'excès d'émission pour WR 140

Tableau II.I: Vitesses radiales des composantes WR et O de WR 140. Les erreurs typiques ( $\sigma$ ) sont déduites de la dispersion des écarts entre les mesures et l'orbite théorique (RV-Fit) pour chaque observatoire. Les vitesses radiales sont en  $\text{km s}^{-1}$ .

JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54813.2605	35.1	-4.26	-101.6	-5.08	3.7	OHP
54816.3396	30.3	-10.54	-100.2	-0.12	3.7	OHP
54818.3191	40.5	-1.40	-100.2	2.37	3.7	OHP
54819.3371	37.9	-4.50	-103.3	0.50	3.7	OHP
54820.2146	39.6	-3.22	-107.9	-2.94	3.7	OHP
54821.2162	42.9	-0.46	-109.3	-3.08	3.7	OHP
54822.2497	42.5	-1.37	-112.0	-4.35	3.7	OHP
54823.2120	41.4	-3.03	-112.5	-3.64	3.7	OHP
54828.2995	49.4	2.33	-114.2	1.14	3.7	OHP
54829.2444	51.3	3.79	-117.6	-1.25	3.7	OHP
54834.2488	44.2	-5.01	-126.5	-6.05	3.7	OHP
54835.2263	43.8	-5.53	-123.6	-2.79	3.7	OHP
54836.2219	49.7	0.43	-120.9	-0.10	3.7	OHP
54843.2282	51.1	6.24	-109.6	0.35	3.7	OHP
54844.2272	40.7	-2.75	-105.6	0.77	3.7	OHP
54845.2257	44.0	2.23	-104.4	-2.00	3.7	OHP
54845.2319	42.0	0.20	-106.1	-3.70	3.7	OHP
54846.2475	48.0	8.08	-93.1	4.75	3.7	OHP
54846.2568	47.2	7.28	-92.3	5.53	3.7	OHP
54847.2226	45.3	7.36	-91.4	1.63	3.7	OHP
54852.2358	23.9	-2.76	-66.3	-0.93	3.7	OHP
54853.2322	34.5	10.08	-53.8	5.93	3.7	OHP
54855.2416	24.9	4.90	-49.1	-0.10	3.7	OHP
54857.2429	25.8	9.80	-34.5	4.58	3.7	OHP
54860.2381	13.7	3.01	-19.5	6.80	3.7	OHP

JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54860.2423	12.7	1.96	-18.7	7.55	3.7	OHP
54861.2456	15.8	6.61	-14.8	7.68	3.7	OHP
54861.2496	13.3	4.17	-16.6	5.87	3.7	OHP
54862.2499	6.7	-1.07	-17.0	2.05	3.7	OHP
54862.2548	5.0	-2.74	-16.5	2.50	3.7	OHP
54873.7413	-3.8	-0.90	7.0	0.01	3.7	OHP
54874.7349	-5.1	-1.63	9.5	1.04	3.7	OHP
54875.7415	-1.5	2.46	9.7	-0.04	3.7	OHP
54877.7303	-7.9	-3.00	12.3	0.22	3.7	OHP
54879.7247	-8.9	-3.17	13.3	-0.83	3.7	OHP
54880.7078	-8.7	-2.55	15.5	0.53	3.7	OHP
54881.7095	-9.4	-2.88	14.2	-1.70	3.7	OHP
54884.7141	-3.7	3.76	17.6	-0.59	3.7	OHP
54885.7229	-11.3	-3.61	20.0	1.19	3.7	OHP
54886.6986	-5.8	2.10	20.3	0.89	3.7	OHP
54888.7010	-8.3	0.12	21.0	0.36	3.7	OHP
54889.7124	-4.6	4.07	18.0	-3.08	3.7	OHP
54890.6953	-15.4	-6.60	19.6	-2.02	3.7	OHP
54893.7053	-9.3	0.10	24.1	1.12	3.7	OHP
54897.6988	-11.9	-1.92	24.9	0.57	3.7	OHP
54897.7015	-8.2	1.72	24.2	-0.22	3.7	OHP
54897.7043	-10.6	-0.63	24.4	0.05	3.7	OHP
54898.7011	-16.5	-6.45	22.9	-1.79	3.7	OHP
54898.7042	-14.3	-4.18	25.4	0.76	3.7	OHP
54898.7073	-16.2	-6.13	24.6	-0.13	3.7	OHP
54900.6852	-14.0	-3.66	21.7	-3.53	3.7	OHP
54900.6883	-13.6	-3.28	22.1	-3.15	3.7	OHP
54900.6914	-18.3	-7.99	20.1	-5.17	3.7	OHP
54901.6822	-8.3	2.15	24.5	-1.02	3.7	OHP
54901.6853	-7.2	3.26	24.5	-0.98	3.7	OHP
54901.6884	-5.0	5.42	26.8	1.29	3.7	OHP
54903.6958	-9.2	1.37	26.5	0.49	3.7	OHP
54905.6836	-12.3	-1.50	23.5	-2.97	3.7	OHP
54906.6655	-3.0	7.88	24.5	-2.14	3.7	OHP
54911.6871	-12.4	-1.20	25.5	-2.04	3.7	OHP



JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54912.6389	-16.1	-4.86	25.6	-2.03	3.7	OHP
54913.6545	-9.4	1.92	28.8	1.03	3.7	OHP
55018.7888	-25.8	-14.04	25.1	-3.67	8.5	OMM
55019.7976	-23.2	-11.44	19.4	-9.33	8.5	OMM
55021.7251	1.8	13.51	32.6	3.92	8.5	OMM
55022.5769	-0.3	11.43	31.0	2.31	8.5	OMM
55023.5802	-12.1	-0.40	26.3	-2.34	8.5	OMM
55026.7784	1.6	13.25	34.1	5.60	8.5	OMM
55028.5613	-	-	27.1	-1.38	8.5	OMM
55033.5434	-19.4	-7.79	17.7	-10.65	8.5	OMM
55035.7731	-10.9	0.60	28.7	0.39	8.5	OMM
55037.7481	-6.0	5.54	32.5	4.28	8.5	OMM
55038.6906	-13.4	-1.85	25.4	-2.85	8.5	OMM
55043.5629	-11.7	-0.28	28.3	0.20	8.5	OMM
55047.5379	-12.7	-1.26	20.3	-7.65	8.5	OMM
55048.5356	-	-	45.0	17.06	8.5	OMM
55049.7346	-34.4	-23.05	16.2	-11.74	8.5	OMM
55050.7467	-6.9	4.45	39.1	11.22	8.5	OMM
55051.7337	-13.0	-1.67	30.1	2.23	8.5	OMM
55052.7340	1.7	13.00	30.2	2.39	8.5	OMM
54578.8834	7.5	-4.06	-26.5	1.79	6.5	DAO
54580.8731	7.6	-4.07	-22.3	6.16	6.5	DAO
54614.7872	19.3	6.40	-36.0	-4.31	6.5	DAO
54615.7836	16.0	3.04	-37.1	-5.31	6.5	DAO
54642.9241	10.7	-3.57	-37.8	-2.94	6.5	DAO
54643.9182	15.6	1.33	-39.2	-4.16	6.5	DAO
54664.7166	25.0	9.59	-36.2	1.58	6.5	DAO
54665.7094	19.0	3.52	-33.3	4.63	6.5	DAO
54675.8039	13.3	-2.79	-34.6	4.86	6.5	DAO
54701.7292	1.0	-16.96	-40.4	3.70	6.5	DAO
54722.7395	9.0	-10.88	-48.4	0.29	6.5	DAO
54723.7921	31.8	11.79	-50.4	-1.51	6.5	DAO
54724.7203	26.7	6.66	-53.9	-4.78	6.5	DAO
54817.3515	39.4	-1.98	-97.0	4.40	7.5	MONS
54818.3541	32.2	-9.68	-106.2	-3.63	7.5	MONS

JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54819.3206	34.3	-8.03	-108.8	-4.97	7.5	MONS
54820.2973	39.9	-2.96	-115.3	-10.26	7.5	MONS
54821.3117	43.8	0.33	-101.9	4.45	7.5	MONS
54822.3031	40.4	-3.58	-108.7	-1.00	7.5	MONS
54824.3186	44.4	-0.59	-	-	7.5	MONS
54827.3022	41.7	-4.83	-118.0	-3.88	7.5	MONS
54834.3241	40.0	-9.16	-	-	7.5	MONS
54835.3351	40.1	-9.16	-99.9	20.93	7.5	MONS
54836.3579	44.2	-5.09	-117.3	3.50	7.5	MONS
54844.3302	39.0	-4.31	-110.0	-3.95	7.5	MONS
54846.3243	44.6	4.82	-85.3	12.16	7.5	MONS
54852.3204	24.8	-1.64	-71.8	-6.98	7.5	MONS
54853.3189	19.2	-5.02	-67.5	-8.24	7.5	MONS
54855.3129	23.4	3.59	-59.4	-10.81	7.5	MONS
54856.3128	25.2	7.38	-48.0	-4.41	7.5	MONS
54858.7387	8.4	-4.83	-28.9	3.54	7.5	MONS
54859.7453	18.9	7.33	-19.5	8.74	7.5	MONS
54860.7312	18.4	8.42	-30.2	-5.75	7.5	MONS
54861.7234	34.0	25.50	-	-	7.5	MONS
54865.7312	8.0	4.50	0.9	9.54	7.5	MONS
54871.7333	-2.6	-0.96	-13.5	-17.38	7.5	MONS
54873.7436	-4.8	-1.93	7.0	-0.01	7.5	MONS
54874.7614	0.3	3.80	3.5	-4.97	7.5	MONS
54876.7646	2.0	6.50	-4.4	-15.39	7.5	MONS
54878.7469	-12.7	-7.37	9.5	-3.66	7.5	MONS
54887.6717	1.4	9.53	28.1	8.06	7.5	MONS
54890.6937	-5.7	3.16	20.1	-1.52	7.5	MONS
54891.6738	-6.9	2.12	29.5	7.42	7.5	MONS
54892.6705	-9.1	0.11	32.9	10.40	7.5	MONS
54900.6975	-14.4	-4.08	-	-	7.5	MONS
54907.6390	-11.3	-0.31	30.0	3.22	7.5	MONS
54912.6621	-12.9	-1.58	38.1	10.47	7.5	MONS
54838.2352	40.4	-8.55	-100.8	19.12	11.0	BS
54840.2001	48.0	0.13	-111.0	6.40	11.0	BS
54841.2043	41.7	-5.38	-101.2	14.25	11.0	BS

JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54842.2012	49.7	3.58	-105.9	7.11	11.0	BS
54843.2125	45.7	0.83	-95.4	14.58	11.0	BS
54845.2213	32.6	-9.20	-102.9	-0.45	11.0	BS
54852.2053	22.7	-4.08	-81.3	-15.77	11.0	BS
54857.2025	27.0	11.02	-66.7	-27.40	11.0	BS
54861.2139	20.9	11.66	-	-	11.0	BS
54861.2140	-	-	-40.5	-17.85	11.0	BS
54445.2470	-	-	-29.6	-10.11	9.0	THO
54446.2376	-	-	10.9	30.46	9.0	THO
54464.3585	-8.3	-16.70	-9.9	10.60	9.0	THO
54472.2292	-	-	-32.7	-11.77	9.0	THO
54482.2175	-8.9	-17.72	-16.0	5.54	9.0	THO
54495.6969	3.5	-5.63	-17.3	5.02	9.0	THO
54572.5640	4.9	-6.41	-43.9	-16.12	9.0	THO
54600.4890	0.8	-11.55	-31.1	-0.84	9.0	THO
54748.3681	19.5	-3.34	-58.5	-2.53	9.0	THO
54795.2708	24.7	-7.74	-80.3	-0.69	9.0	THO
54800.2354	29.8	-4.36	-84.3	-0.69	9.0	THO
54802.3446	38.7	3.84	-83.6	1.81	9.0	THO
54803.2519	31.6	-3.57	-85.0	1.20	9.0	THO
54810.2058	28.5	-9.54	-97.1	-4.02	9.0	THO
54811.2458	42.8	4.32	-118.7	-24.48	9.0	THO
54826.2450	55.2	9.14	-115.3	-2.51	9.0	THO
54827.2396	45.9	-0.59	-111.6	2.44	9.0	THO
54828.2719	47.3	0.28	-106.6	8.68	9.0	THO
54829.2202	-	-	-107.0	9.33	9.0	THO
54830.2334	52.2	4.25	-113.6	3.80	9.0	THO
54831.2206	-	-	-119.3	-0.94	9.0	THO
54833.2167	57.3	8.38	-112.2	7.72	9.0	THO
54837.2858	-	-	-123.2	-2.69	9.0	THO
54844.2330	39.6	-3.81	-106.9	-0.52	9.0	THO
54845.2511	43.5	1.74	-107.3	-4.95	9.0	THO
54851.2443	32.3	3.36	-81.6	-10.63	9.0	THO
54854.2303	24.7	2.56	-62.3	-8.04	9.0	THO
54855.2592	34.3	14.35	-52.8	-3.91	9.0	THO

JD-2400000	RV (O)	RV-Fit	RV (WR)	RV-Fit	$\sigma$	Obs.
54856.2249	26.6	8.67	-45.1	-1.16	9.0	THO
54857.2431	28.7	12.75	-39.0	0.06	9.0	THO
54866.6723	-	-	-3.3	3.00	9.0	THO
54870.2430	-0.8	-0.31	7.0	5.71	9.0	THO
54874.2506	0.3	3.48	13.6	5.82	9.0	THO
54882.6374	-	-	14.7	-1.93	9.0	THO
54892.6583	1.9	11.12	25.6	3.13	9.0	THO
54895.5894	-6.6	3.02	27.8	4.18	9.0	THO

Tableau II.II: Mesures des vitesses radiales, largeurs, et largeurs équivalentes de l'excès d'émission de la raie C III 5696 pour les données OHP. Les mesures sont en  $\text{km s}^{-1}$ .

JD-2400000	RV	$\sigma$	FW	$\sigma$	EW	$\sigma$
54819.3371	-1302	292	1935	819	2.38	0.34
54820.2146	-1221	338	2105	866	1.65	0.28
54821.2162	-1005	214	2386	506	3.30	0.40
54822.2497	-1303	276	1951	770	2.67	0.36
54823.2120	-1074	213	2166	533	3.46	0.41
54828.2995	-823	183	2613	400	3.84	0.43
54829.2444	-1003	209	2131	509	3.24	0.40
54834.2488	-316	120	2867	257	5.81	0.53
54835.2263	-256	119	2931	253	5.92	0.54
54836.2219	-230	115	2917	246	6.14	0.55
54843.2282	318	95	2956	202	9.78	0.69
54844.2272	395	100	3005	210	9.63	0.69
54845.2257	468	97	2804	207	9.82	0.70
54845.2319	450	95	2815	203	10.06	0.70
54846.2475	671	112	2697	237	8.84	0.66
54846.2568	626	110	2740	233	8.77	0.66
54847.2226	579	98	2827	208	10.95	0.74
54852.2358	858	121	2936	253	10.34	0.71
54853.2322	1303	161	2024	438	7.89	0.62

JD-2400000	RV	$\sigma$	FW	$\sigma$	EW	$\sigma$
54857.2429	1411	187	1979	526	6.70	0.57
54860.2381	1561	240	1747	769	4.67	0.48
54860.2423	1545	231	1790	722	5.01	0.50
54861.2456	1451	210	1948	600	5.54	0.52
54861.2496	1484	218	1898	640	5.28	0.51
54862.2499	1448	220	1915	640	5.01	0.50
54862.2548	1421	208	1942	596	5.44	0.52
54873.7413	1438	330	1695	1039	2.12	0.32
54874.7349	1369	296	1844	875	2.47	0.35
54875.7415	1048	241	2575	539	2.92	0.38
54877.7303	448	204	3261	407	2.77	0.37
54879.7247	1351	364	1898	1057	1.62	0.28
54880.7078	1385	437	1520	1457	1.11	0.23
54884.7141	1288	427	1710	1276	1.05	0.22
54890.6953	202	197	3165	411	2.42	0.34



### Annexe III

#### Mesures des vitesses radiales pour WR 12, WR 21, WR 30, WR 31 et WR 47

Tableau III.I: Vitesses radiales de la composante WR de WR 12. Les erreurs typiques ( $\sigma$ ) sont déduites de la dispersion des écarts entre les mesures et l'orbite théorique (RV-Fit). Les vitesses radiales sont en  $\text{km s}^{-1}$ .

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54903.5563	-144.7	-36.4	18.0
54907.5812	2.1	-24.5	18.0
54908.5652	30.2	-31.2	18.0
54909.5517	114.1	22.0	18.0
54910.6056	87.4	-30.7	18.0
54911.5208	122.5	-10.9	18.0
54916.6024	117.0	38.8	18.0
54918.6530	36.7	30.4	18.0
54919.6220	-22.4	7.1	18.0
54920.6204	-37.0	27.5	18.0
54921.6037	-104.5	-9.9	18.0
54922.6396	-74.8	44.8	18.0
54924.5701	-146.3	-4.7	18.0
54925.6572	-144.6	-6.3	18.0
54926.6219	-112.2	13.8	18.0
54927.6143	-121.0	-16.0	18.0
54928.6120	-65.5	11.2	18.0
54930.5986	-7.6	-0.6	18.0
54931.6440	37.7	6.0	18.0
54932.5990	41.6	-23.4	18.0
54933.6252	96.0	-0.3	18.0
54934.6318	100.0	-20.2	18.0
54935.6256	119.6	-15.9	18.0
54936.5964	151.0	9.2	18.0

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54937.6252	158.7	20.4	18.0

Tableau III.II: Vitesses radiales des composantes WR et O de WR 21.

JD-2400000	RV (WR)	RV-Fit	$\sigma$	RV (O)	RV-Fit	$\sigma$
54907.6291	-175.4	-0.3	16.0	66.5	-30.3	24.0
54908.6162	-235.9	-26.3	16.0	78.7	-37.2	24.0
54909.5956	-159.6	-27.3	16.0	94.3	21.2	24.0
54910.6455	0.5	-26.1	16.0	-70.8	-56.1	24.0
54911.5639	158.9	3.1	16.0	-101.4	-15.2	24.0
54916.6407	-220.3	-8.3	16.0	110.1	-7.2	24.0
54918.6942	-27.2	-20.7	16.0	-66.6	-70.2	24.0
54919.6623	116.1	-21.6	16.0	-75.8	0.3	24.0
54920.6585	217.7	6.8	16.0	-108.2	8.4	24.0
54921.6650	162.5	-5.1	16.0	-69.5	23.1	24.0
54922.7070	27.5	2.6	16.0	-18.5	-4.7	24.0
54924.6211	-187.4	19.0	16.0	112.7	-1.5	24.0
54925.7140	-171.7	3.4	16.0	85.0	-11.8	24.0
54926.6780	-52.4	-2.6	16.0	-25.5	-53.0	24.0
54927.6586	110.0	6.6	16.0	-58.1	-0.9	24.0
54928.6644	224.9	22.0	16.0	-60.4	51.8	24.0
54930.6523	58.4	-15.0	16.0	-6.5	34.1	24.0
54931.7054	-82.8	9.2	16.0	65.8	14.9	24.0
54932.6612	-188.6	7.2	16.0	124.3	16.0	24.0
54933.6797	-206.4	-9.5	16.0	106.5	-2.4	24.0
54934.6763	-88.8	-0.0	16.0	39.6	-9.5	24.0
54935.6747	82.6	14.2	16.0	-73.2	-35.4	24.0
54936.6447	219.6	33.9	16.0	-79.1	23.6	24.0
54937.6643	240.8	36.5	16.0	-93.4	19.5	24.0



Tableau III.III: Vitesses radiales des composantes WR et O de WR 30.

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54907.6837	136.8	14.5	18.0
54908.6678	145.7	-14.0	18.0
54909.6594	184.3	4.1	18.0
54910.6861	161.8	-18.9	18.0
54911.5993	160.3	-3.1	18.0
54916.6979	-72.4	31.1	18.0
54917.6402	-171.5	-26.3	18.0
54918.7316	-145.0	30.4	18.0
54919.7047	-183.9	-0.9	18.0
54920.6942	-162.7	8.3	18.0
54921.7504	-92.6	45.3	18.0
54924.6916	20.0	-3.7	18.0
54926.7547	164.7	31.0	18.0
54927.7018	194.3	28.2	18.0
54928.7150	210.1	27.9	18.0
54930.6968	177.5	22.9	18.0
54932.7083	78.1	18.2	18.0
54933.7357	8.2	10.0	18.0
54934.7282	-67.5	-6.2	18.0
54935.7189	-114.3	-0.2	18.0
54936.7054	-153.9	0.4	18.0
54937.7091	-155.0	23.1	18.0

Tableau III.IV: Vitesses radiales des composantes WR et O de WR 31.

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54907.7567	-10.6	70.4	20.0
54908.7330	-201.4	-20.6	20.0
54909.7360	-35.4	-15.4	20.0
54911.6541	131.3	5.7	20.0

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54916.7820	35.5	-30.6	20.0
54917.7308	-129.3	10.1	20.0
54918.7946	-150.0	-7.4	20.0
54919.7693	82.0	15.0	20.0
54920.7647	151.1	-31.1	20.0
54922.7801	-182.0	-14.9	20.0
54924.7597	109.2	8.5	20.0
54925.7811	205.2	32.6	20.0
54927.7702	-182.1	-3.2	20.0
54928.7733	-93.4	-8.8	20.0
54930.7637	157.6	0.4	20.0
54931.7519	-57.4	-11.6	20.0
54932.7611	-170.1	13.0	20.0
54936.7756	-105.8	-17.4	20.0
54937.7777	-171.7	6.3	20.0

Tableau III.V: Vitesses radiales des composantes WR et O de WR 47.

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54902.7560	111.1	27.0	23.0
54903.6673	-177.4	-30.6	23.0
54906.7533	102.9	-36.4	23.0
54907.7883	272.0	9.0	23.0
54908.7623	126.8	-13.0	23.0
54909.7669	-147.5	-32.7	23.0
54911.6880	-205.7	-25.3	23.0
54915.8271	-96.2	-25.8	23.0
54916.8175	-293.3	-42.1	23.0
54917.7636	-170.8	38.6	23.0
54918.8238	64.0	25.4	23.0
54919.8024	255.4	17.2	23.0
54920.8027	244.7	23.1	23.0

JD-2400000	RV (WR)	RV-Fit	$\sigma$
54921.8172	-10.9	-5.6	23.0
54922.8255	-215.0	11.3	23.0
54924.7949	16.3	48.6	23.0
54925.8195	234.3	26.7	23.0
54926.8363	249.8	3.8	23.0
54927.8065	55.5	-5.0	23.0
54928.8116	-193.1	-8.0	23.0
54930.8108	-73.3	16.4	23.0
54931.7931	153.1	-4.3	23.0
54932.7973	291.4	29.1	23.0
54933.8176	141.0	23.8	23.0
54934.8086	-148.3	-13.9	23.0
54936.8118	-181.0	-35.0	23.0
54937.8039	92.7	-12.4	23.0