METALLICITY IN THE GALACTIC CENTER: THE QUINTUPLET CLUSTER

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ABSTRACT

We present a measurement of metallicity in the Galactic center Quintuplet cluster made using quantitative infrared spectroscopy of two luminous blue variables (LBVs). The analysis employs line-blanketed non-LTE wind/atmosphere models fit to high-resolution near-infrared spectra containing lines of H, He I, Si II, Mg II, and Fe II. We are able to break the H/He ratio versus mass-loss rate degeneracy found in other LBVs and to obtain robust estimates of the He content of both objects. Our results indicate solar iron abundance and roughly twice solar abundance in the α-elements. These results are discussed within the framework of recent measurements of oxygen and carbon composition in the nearby Arches cluster and iron abundances in red giants and supergiants within the central 30 pc of the Galaxy. The relatively large enrichment of α-elements with respect to iron is consistent with a history of more nucleosynthesis in high-mass stars than the Galactic disk.

Key words: Galaxy: abundances – Galaxy: center – infrared: stars – stars: abundances – stars: individual (Pistol Star, FMM362)

Online-only material: color figures

1. INTRODUCTION

Elements heavier than hydrogen and helium (metals) are primarily created by nucleosynthesis in stars. Metals are important ingredients in many astrophysical processes such as radiative cooling, in mass loss during star formation and at all stages of stellar evolution. They also play a fundamental role in stellar evolution through their influence on stellar opacities, and represent a historical record of galactic chemical enrichment via stellar winds and supernovae ejecta.

In the Galaxy metal abundance increases with decreasing galactocentric radius, as seen in stars and gas (Afflerbach et al. 1997; Rudolph et al. 2006; Maciel & Quireza 1999; Fuhrmann 1998; Rolleston et al. 2000; Smartt et al. 2001; Luck et al. 2006). Other galaxies show a similar trend, having highest metal abundances in their nuclei (Urbaneja et al. 2005; Kennicutt et al. 2003).

Previous work (Frogel et al. 1999; Feltzing & Gilmore 2000; Carr et al. 2000; Ramirez et al. 1997, 1999, 2000) on the Galactic center (GC) has indicated roughly solar stellar metal abundances, whereas the analyses of interstellar emission lines (Shields & Ferland 1994; Maeda et al. 2000) have suggested considerably higher abundances. It is not clear why the stellar and gas-phase measurements should differ so greatly.

The GC contains three dense and massive star clusters that have recently formed in the inner 50 pc, the Arches, Quintuplet, and Central clusters. Using quantitative spectral analysis, Najarro et al. (2004, Paper I) determined that the WNL stars in the very young (2–2.5 Myr) Arches cluster have roughly solar metallicities. Being more evolved (∼ 4 Myr), the Quintuplet cluster (Glass et al. 1987, 1990; Nagata et al. 1990; Okuda et al. 1990; Moneti et al. 1994) contains a variety of massive stars, including WN, WC, WN9/Ofpe, luminous blue variables (LBVs) and less evolved blue supergiants (Figer et al. 1995, 1999a, 1999b). Two LBVs in it are known, the Pistol Star (Moneti et al. 1994; Cotera et al. 1994; Figer et al. 1995, 1998, 1999c), and FMM362 (Figer et al. 1999b; Geballe et al. 2000), each having an infrared spectrum rich in metal lines of Fe II, Si II, and Mg II.

In this paper, we use quantitative infrared spectroscopy of the two Quintuplet LBVs to make direct determinations of metallicity in those stars. We also use the derived α-elements versus Fe ratio to address the dominance of massive stars on the IMF in this region.

2. OBSERVATIONAL DATA

The data were obtained at UKIRT6 using CGS4. The Pistol Star was observed in 1996 April (Rc: R ∼ 3000), 1997 July (L; R ∼ 16000), and 1998 April (H; R ∼ 5000 and K R ∼ 3000). Likewise, spectra for FMM362 were obtained in April (L) and May 1999 (H and K), using CGS4 in medium resolution mode (R ∼ 5000–6500). The slit width was 0′6 for all observations. We used the photometric measurements of Figer et al. (1998) for the Pistol Star, to scale the reduced spectra. For FMM362, given its photometric variability, we adopted the average value, K = 7.30, obtained by Glass et al. (1999) for the epoch closest to our spectroscopic observations. This value agrees, within the 0.26 standard deviation derived by Glass et al. (1999), with the K = 7.50 value adopted by Geballe et al. (2000) from flux-calibrated spectra. We assume the same extinction for both objects and adopt the value of A_K = 3.2 derived by Figer et al. (1998) for the Pistol Star. The reader is referred to these papers for a detailed discussion on the reduction of the observed spectra and photometry.

6 The United Kingdom Infrared Telescope (UKIRT) is operated by the Joint Astronomy Centre on behalf of the Particle Physics and Astronomy.
3. MODELS

To model the LBVs and estimate their physical parameters, we have used CMFGEN, the iterative, non-LTE line blanketing method presented by Hillier & Miller (1998) which solves the radiative transfer equation in the comoving frame and in spherical geometry for the expanding atmospheres of early-type stars. The model is prescribed by the stellar radius, $R_*$, the stellar luminosity, $L_*$, the mass-loss rate, $M_*$, the velocity field, $v(r)$ (defined by $V_\infty$ and $\beta$), the volume filling factor characterizing the clumping of the stellar wind, $f(r)$ (see Section 4.2), and elemental abundances. Hillier & Miller (1998, 1999) present a detailed discussion of the code. For the present analysis, we have assumed the atmosphere to be composed of H, He, C, N, O, Mg, Si, S, Fe, and Ni. Given the parameter domain the LBVs are located, the $\tau = 2/3$ radius is located close or above the sound speed, and therefore the assumed hydrostatic structure plays no role. Thus, no spectroscopic information about the mass of the object can be obtained. The atomic data sources are described in detail Hillier et al. (2001). Here, we focus on the model atoms used for our abundance determinations Fe ii, Mg ii, and Si ii. Using the superlevel formalism (NS/NI, number of superlevels versus number of levels in the full atom; e.g., Hillier & Miller 1998), we chose 233 $f_0$ for Mg ii, 37/50 (up to 119400 cm$^{-1}$) for Fe ii, 37/50 (up to 119400 cm$^{-1}$) for Mg ii, and 37/50 (up to 125000 cm$^{-1}$) for Si ii. The choice of the appropriate packing has been extensively tested in Najarro (2001). We will revise the importance of this issue for the case of Mg in Section 4.4.

Observational constraints are provided by the H, K, and L band spectra of the stars and the reddened K magnitudes from Figer et al. (1998), Geballe et al. (2000) and Glass et al. (1999). As in Paper I, a distance of 8 kpc has been assumed. The validity of our technique has been demonstrated in Najarro et al. (1999) and Najarro (2001) by calibrating our method against stars with similar spectral type such as P Cygni and HDE 316285 for which not only infrared but also optical and UV spectra are available.

4. RESULTS

Table 1 gives the derived stellar parameters for both LBVs, and Figures 1 and 2 show model fits to the relevant lines in the stars. Theoretical spectra have been convolved with the instrumental resolution. We note that given the large number of parameters involved in the analysis, it is unaffordable to perform a full systematic error analysis in the whole parameter domain. We rather proceed by estimating the range of values for the main stellar parameters which provide acceptable fits to the observed spectra. Once those ranges are set, we derive the corresponding abundances and their errors. From Table 1, it can be seen that the Pistol Star and FMM362 have very similar properties, with the exceptions of the Pistol Star’s significantly higher wind density (evidenced by the its stronger spectral lines) and its higher He content. The latter may denote a slightly advanced evolutionary stage for the Pistol Star (see below). Given the general resemblance of the spectra of the objects, we discuss them together.

4.1. Main Diagnostic Lines and Stellar Properties

Several spectral diagnostics constrain our estimates of the stellar temperature, and thus the ionization structure, in particular the He i (5–4) components near 4.05 $\mu$m. If helium is predominantly singly ionized, even for the most favorable case with (minimum) cosmic helium abundance, the observed ratio of H to He i lines exceeds the expected values by large factors.

| Parameter | Pistol | FMM362 |
|-----------|--------|--------|
| $L_*$ (10$^8 L_\odot$) | 1.60 | 1.77 |
| $R_*$ (R$_\odot$) | 306 | 350 |
| $T_{eff}$ (10$^4$ K) | 1.18 | 1.13 |
| H/He | 1.5 | 2.8 |
| Fe/Fe$^*$ | 1.1 (0.78) | 1.1 (0.78) |
| Mg/Mg$^*$ | 2.2 | 1.5 |
| Si/Si$^*$ | 1.8 | 2.1 |
| $\beta$ | 3.0 | 1.3 |
| $V_\infty$ (km s$^{-1}$) | 105 | 170 |
| $D_{atom} = \log (M V_\infty (R/R_\odot)^{1/2})$ | 29.39 | 29.38 |
| $M_{Edd}$ (M$_\odot$) | 22.5 | 30.5 |
| CL1 | 0.08 | 0.08 |
| CL2 | 2.5 | 2.00 |
| CL3 | 2.00 | – |

Notes. $\beta$ is the exponent describing the velocity field, $D_{atom}$ is the modified wind momentum (Kudritzki & Puls 2000) and clumping parameter, the CL are defined in Equation (1). $M_{Edd}$ are the Eddington masses. H/He is the ratio by number, and other abundances are relative to solar after Grevesse & Noels (1993). The Fe abundance relative to the value in Anders & Grevesse (1989) is given in parentheses (see text).

This indicates that He ii must recombine to He i very close to the photosphere, implying an upper limit of around 13,000 K for the temperatures of these objects. We find a lower limit of 10,000 K for the temperatures, as lower values would require nondetection of the He i components. Also the strengths of the He i 1700 $\mu$m and He i 2.112/3 $\mu$m lines are very sensitive to temperature, so that they appear in emission above 12,500 K and vanish below 10,500 K. These lower limits on the effective temperature are also consistent with the nondetections of the Si ii $3s^2 3p^2 S_{1/2} \rightarrow 3s^2 3p^2 D_{3/2}$ 2.180 $\mu$m and $3s^2 3p^2 S_{1/2} \rightarrow 3s^2 3p^2 D_{3/2}$ 2.209 $\mu$m intercombination lines, as they are expected to appear strongly in absorption as soon as the temperature drops below 10,000 K. Hence, $\Delta T \pm 1500$ K are conservative estimates of uncertainties for the temperatures of the Pistol Star and FMM362 given in Table 1.

To estimate the terminal velocities, we make use of the Fe ii] (semiforbidden Fe ii] z$^2$(F$_y/2$ e$^{-2}$F$_y/2$ 1.688 $\mu$m line (Geballe et al. 2000; Figer et al. 1998) that forms in the outer wind and has a weak oscillator strength (gf $\sim 10^{-5}$). This is because nonnegligible continuum opacity effects at 4 $\mu$m may provide only lower limits if Br $\alpha$ is used. The larger $V_\infty$ derived for FMM362 can be clearly inferred from the width of this Fe ii] line and the obvious overlap at Br between the H i and He i components (see Figure 2). For the Pistol Star (Figure 1), the Br components are fairly well separated.

Our analysis of wind density ($M_*, \beta$) gives values of $\beta$ that agree with those inferred in the literature for other LBVs (Najarro 2001) and B-supergiants (e.g., Crowther et al. 2006). The values are fairly well constrained by the shapes of the hydrogen lines, especially those of Br $\alpha$ and Br $\gamma$ which are inconsistent with the same value of $\beta$ for both objects.

Although the wind density derived for the Pistol Star is much higher than for FMM362, the modified wind momenta $D_{atom} = \log (M V_\infty \sqrt{R/R_\odot})$ (Kudritzki & Puls 2000) of the two LBVs are nearly identical (Table 1). This result is qualitatively consistent with the wind momentum–luminosity relation (Kudritzki & Puls 2000) which predicts the same modified momenta for objects with the same stellar type and

Table 1

\begin{table}
\begin{tabular}{|c|c|c|}
\hline
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\hline
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\hline
\end{tabular}
\end{table}
Figure 1. Model fits (dashed lines) to the observed infrared diagnostic lines (solid lines) of the Pistol Star. The forbidden [Fe\textsc{ii}] line at 1.677\,μm was not included in the models. (A color version of this figure is available in the online journal.)

 luminosity. Note that we have used clumping-corrected values of \(\dot{M}\) to compute the modified momenta. If unclumped values were assumed, \(D_{\text{mom}}\) would be closer to 30.0. Interestingly, the latter agrees very well with the averaged modified momentum of AG Car at maximum (i.e., at similar \(T_{\text{eff}}\) to our objects) obtained using the values of \(\dot{M}, V_\infty\), and \(R_*\) derived by Stahl et al. (2001) from fits to Hα. Those authors obtained \(\log(\dot{M}) \sim -4.1\) for about this temperature, while our unclumped values are \(\log(\dot{M}) = -4.1\) and \(-4.4\) for the Pistol Star and FMM362, respectively.

Further, using radiation-driven wind models for LBVs, Vink & de Koter (2002) were able to predict the Stahl et al. (2001) \(\dot{M}\) value for AG Car assuming \(V_\infty/N_{\text{esc}} \sim 1.3\) and a current stellar mass of 35\,\(M_\odot\). The same value of \(V_\infty/N_{\text{esc}}\) for the LBVs would imply current stellar masses of 27.5\,\(M_\odot\) for the Pistol Star and 46\,\(M_\odot\) for FMM362. Although these masses should be regarded with caution, they are consistent with the Pistol Star being more evolved than FMM362 (as inferred from their He/H ratios), and hence having lost more mass during its evolution, as indicated by the presence of a nebula around it.

Compared to the results obtained in Figer et al. (1998) by means of nonblanketed models, the new blanketed models provide a significant improvement in our knowledge of the physical properties of these two stars. The degeneracy of the “high” and “low” luminosity \((T_{\text{eff}})\) solutions for the Pistol Star presented by Figer et al. (1998) is broken by the Si\textsc{ii}, Mg\textsc{ii}, and Fe\textsc{ii} lines, which are clearly more consistent with the “low” solution. We derived for this star a luminosity of \(\sim 1.6(10^6)\,L_\odot\), an effective temperature of \(\sim 11,800\,\text{K}\), and an initial mass of 100\,\(M_\odot\). The stellar luminosity is reduced by a factor of two compared with the previous estimate, illustrating the importance of the new generation of line-blanketed models. Below, we discuss in detail the role of two additional stellar properties that are derived using the new models, wind clumping, and elemental abundances.

4.2. Clumping

Clumping is normally invoked in stellar winds to explain inconsistencies arising between \(\rho\) (density) and \(\rho^2\) diagnostics. For a given mass loss, clumping causes an enhancement of \(\rho^2\) processes while leaving unaltered those which depend linearly on \(\rho\). Further, if mass-loss rate and clumping are scaled without changing the \(\dot{M}/f^{0.5}\) ratio, the \(\rho\)-dependent diagnostics vary, while the recombination lines profiles \((\propto \rho^2)\) remain basically unaltered.

To investigate the clumping, we introduce the following clumping law:

\[
f = CL_1 + (1 - CL_1)e^{V/CL_2} + (CL_4 - CL_1)e^{(V - V_\infty)/CL_3},\tag{1}
\]

where CL\(_1\) and CL\(_4\) are volume filling factors, and CL\(_2\) and CL\(_3\)
are velocity terms defining locations in the stellar wind where the clumping structure changes. CL₁ sets the maximum degree of clumping reached in the stellar wind (provided CL₄ > CL₁), while CL₂ determines the velocity of the onset of clumping. CL₃ and CL₄ control the clumping structure in the outer wind. Hence, when the wind velocity approaches $V_{\infty}$, so that $(V - V_{\infty}) \leq CL₃$, clumping starts to migrate from CL₁ toward CL₄. If CL₄ is set to unity, the wind will be unclumped in the outermost region. Such a behavior was already suggested by Nugis et al. (1998) and was utilized by Figer et al. (2002) and Najarro et al. (2004) for the analysis of the WNL stars in the Arches cluster. Recently, Puls et al. (2006) have also found similar behavior from Hα and radio studies of OB stars with dense winds. Furthermore, our clumping parametrization is consistent with results from hydrodynamical calculations by Runacres & Owocki (2002).

From Equation (1), we note that if CL₃ and therefore CL₄ is not considered (CL₃ → 0), we recover the simpler variation proposed by Hillier & Miller (1999). To avoid entering free parameters heaven, we set CL₄ = 1 in all of our investigations, aiming to get an appropriate amount of leverage on the amount of nonconstant clumping in the outer wind regions.

**CL₁: Estimating the wind clumpiness.** Figure 3 illustrates the sensitivity to CL₁ of the main diagnostic lines utilized to obtain the degree of clumping in the stellar winds of FMM362. For each value of CL₁ displayed in Figure 3 the mass-loss rate in the model was scaled while keeping $\dot{M}/\dot{M}_{\odot}^{0.5}$ constant as described previously. Although there are some lines that follow this scaling quite well (Brα, and also Brγ and some metal lines not displayed in the figure), the He i lines and weak H i lines react quite sensitively to the absolute degree of wind clumping. It can be seen that the He i lines only provide an upper limit to CL₁ (a bit lower than 0.1) and do not react to lower values, but a unique value for CL₁ can be selected by some H i lines. The Hα (H i(14–6)) line displays the highest sensitivity to clumping. Unfortunately, the wavelength interval surrounding this line was not observed with sufficiently high signal-to-noise ratio (S/N) in FMM362 and there is a large uncertainty in the continuum value, which is critical for estimating CL₁. We could make full use of the Hα line to determine the clumping only for the Pistol Star, where this line is relatively stronger in emission (see Figure 1). Nevertheless, Figure 3 shows that CL₁ lies between 0.1 and 0.05. It must be stressed that only with a well-determined
clumping may we address the He/H abundance issue (see Section 4.3).

**CL\(_2\) and CL\(_3\): Mapping the clumping structure.** The upper panels of Figure 4 illustrate the behavior of the clumping structure for different sets of CL\(_2\) and CL\(_3\) values, while the lower panels display the influence of such behavior on diagnostic lines in the spectrum of the Pistol Star. It is evident that for some spectral lines, e.g., He\(_i\) 2.112 µm, the behavior of the profiles with clumping is far from being monotonic. Furthermore, not only do lines of different ions react differently to clumping, but also lines within the same ion, e.g., H\(_i\), behave differently. For example, Fe\(_{ii}\) does not respond in the same way to changes in CL\(_2\) and CL\(_3\). Figure 4 shows the great potential of the different IR lines to constrain the clumped structure of the stellar wind and demands the following detailed discussion.

The general impact of clumping on line profiles that was described at the beginning of this section will occur provided the ionization equilibrium is on the “safe” side. We consider the “safe” region to be where the population of the next ionization stage clearly dominates over the one the line belongs to (i.e., H\(_{ii}\) ≫ H\(_i\) for the hydrogen lines). Noting, however, that ionization depends linearly on density whereas recombination is proportional to \(\rho^2\), a “changing” ionization situation may occur, where two adjacent ionization stages have similar populations. In such a case clumping, which enhances recombination, will cause a net reduction of the mean ionization. This will result in weaker lines.

Finally, in the infrared, via bound-free and free-free processes \(\propto \rho^2\), not only the lines but also the continuum will depend on clumping, resulting in high sensitivity of the continuum-rectified line profiles to CL\(_1\) and CL\(_2\).

One may, therefore, distinguish between lines formed on the “safe” region and those arising from the “changing” region.

Within the parameter domain of the two LBVs studied here, we find that H\(_i\), Si\(_{ii}\), and Mg\(_{ii}\) lines and also the Fe\(_{ii}\) photospheric lines are formed in “safe” regions, while He\(_i\) and Fe\(_{ii}\) lines arise from “changing” regions.

Increasing the clumping (decreasing the CL\(_1\) value), or alternatively decreasing the velocity at which clumping sets in (decreasing CL\(_2\)) results in stronger Si\(_{ii}\) and Mg\(_{ii}\) lines, as shown in Figure 4(left). The strong H\(_i\) lines are formed further out than the Si\(_{ii}\) and Mg\(_{ii}\) lines and their strengths should in principle show no sensitivity to CL\(_2\). However, the continuum is clearly affected by clumping. Thus, the stronger the clumping (lower CL\(_2\)), the stronger the continuum and the weaker the resulting line-to-continuum ratio, as clearly shown by B\(_{\gamma}\) and B\(_{\alpha}\) in Figure 4(left). However, weaker H\(_i\) lines such as H\(_{14}\) or B\(_{10}\) and B\(_{11}\) form much closer to the photosphere and tend to brighten with increasing clumping. The weak Fe\(_{ii}\) lines formed close to the photosphere react in basically the same way as the continuum and thus the normalized spectra of them show no changes with changing clumping. Their near independency allows these lines to be used as Fe abundance indicators (see below). The Fe\(_{ii}\) lines, formed even beyond the H\(_i\) lines, are affected by two competing processes. On one hand, increasing the extent of the clumped region (decreasing CL\(_2\)) results in a reduction of the Fe\(_{iii}\)/Fe\(_{ii}\) ratio in the wind. Since Fe\(_{iii}\) remains the dominant ionization stage in the Fe\(_{ii}\) line-formation zone, this change will cause a slight increase in the strengths of the Fe\(_{ii}\) lines. On the other hand, as the continuum increases with increasing clumping, the line-to-continuum ratio decreases. Thus, the two processes counter-balance (see Figure 4, left). Finally, the He\(_i\) lines, which form close to the photosphere, show weak continuum dependencies, but high sensitivities to ionization/recombination. Thus, starting with the model with the highest CL\(_2\) values, the He\(_i\) lines are not affected by clumping, but the stellar parameters produce strong ionization in the inner parts.
resulting in overly strong emission (He\textsc{i} 1.700 \mu m) and line filling (He\textsc{i} 2.112 \mu m). However, as clumping is enhanced in the line-formation zone, recombination starts to dominate over ionization and the He\textsc{i} line emission weakens, the lines are no longer as filled, and start to appear in absorption.

Regarding clumping in the outer parts of the wind, it can be seen on the right side of Figure 4 that only the strong H\textsc{i} and Fe\textsc{ii} lines react to CL 3. Note that Br\gamma which forms further out than Be\gamma is more sensitive to clumping and the observed ratio of the two profiles may be used to determine CL3. For winds of significantly lower density, these lines will form further in and show little or no dependence on CL3 (e.g., FMM362). The Fe\textsc{ii} lines are more sensitive to CL3, primarily due to the coupling of the Fe ionization structure with that of hydrogen through charge-exchange reactions in the outer wind zones where H\textsc{ii} starts to recombine. Due to the difference in H and Fe abundances, a small and hardly noticeable change in the ionization of hydrogen will be amplified in the Fe\textsc{iii}/Fe\textsc{ii}
Figure 5. Breakdown of the H/He degeneracy in FMM362. Models with H/He ratios ranging from 5.0 to 0.75 provide identical H\textsc{i} and strong He\textsc{i} line profiles, but weaker He\textsc{i} profiles can be used to determine the He abundance (see text). (A color version of this figure is available in the online journal.)

Because of the lower wind density of the Quintuplet LBVs and the sensitivities of some of the infrared lines to the stellar parameters, we are able to break the H/He degeneracy and obtain robust estimates of their He content. Due to the high degree of clumping found in both objects, the $\tau = 2/3$ radius, where $T_{\text{eff}}$ is defined, is reached at considerably lower velocities than for classical LBVs. Thus, wind speeds roughly between half and one-third of the sound speed are found in the Pistol Star and FMM362, while classical LBVs have wind speeds well above the speed of sound (Hillier et al. 1998; Najarro 2001). This enables quasi-photospheric absorption lines to form. The He\textsc{i} 2.112 $\mu$m line is the key in breaking the degeneracy. This is shown for FMM362 in Figure 5, which contains model spectra computed for H/He ratios ranging from H/He = 5.0 to 0.75, with mass-loss rates and metal abundances scaled and the other stellar parameters fine-tuned to reproduce the observed profiles of other lines. The figure shows that while identical H and He\textsc{i} (He\textsc{i} 1.700 $\mu$m) line profiles (also for the rest of hydrogen and metal lines) are obtained for all H/He ratios considered, the absorption depths of the He\textsc{i} 2.112 $\mu$m and He\textsc{i} 2.15 $\mu$m lines react sensitively to the He abundance. Both lines show that the best H/He value must lie between 3.33 and 2.25, and we find a most likely value of 2.8 (see Table 1). Similar behavior was found for the Pistol Star, where we obtain H/He = 1.5.

4.4. Metal Abundances

For the purpose of discussing metal abundances (see Table 2), we adopt the solar composition of Grevesse & Noels (1993). Although their abundances have been recently revised (Asplund et al. 2005; Allende Prieto 2008) (but see also Pinsonneault & Delahaye (2006)), they are the ones used by Iglesias & Rogers (1996) to compute stellar interior opacities and adopted in the
most recent evolutionary models for massive stars with rotation from the Geneva group (Meynet & Maeder 2003, 2005), and the Padova tracks used for cooler, less massive stars (Girardi et al. 2000; Salasnich et al. 2000). Previously published evolutionary tracks for massive stars (Schaller et al. 1992; Meynet et al. 1994) used opacity tables calculated with solar composition (Anders & Grevesse 1989), which differ significantly versus 7.50 in Grevesse & Noels (1993) and very slightly in 1994) used opacity tables calculated with solar composition (Anders & Grevesse 1989), which differ significantly versus Grevesse & Noels (1993) only in Fe (A(Fe/H) = 7.67 versus 7.50 in Grevesse & Noels (1993)) and very slightly in the CNO ratios (A(C/H) = 8.56, A(N/H) = 8.05, A(O/H) = 8.93 in Anders & Grevesse (1989) versus A(C/H) = 8.55, A(N/H) = 7.97, A(O/H) = 8.87 in Grevesse & Noels (1993)). Nevertheless, we have also listed in Table 1, in parentheses, the measured abundances with respect to the solar Fe values from Anders & Grevesse (1989). Si and Mg are the same in all evolutionary models, and have been only slightly revised downward (~0.05 dex) by Asplund et al. (2005). Thus, the reader should note that current discussions found in the literature on the derived α-elements versus Fe ratio may depend critically on the assumed Fe solar abundance.

Iron. Two types of Fe II lines are found in the spectra. The first are the strong semi-forbidden lines, including zF_{5/2}^2-c^4F_{7/2} 1.688 μm and z^4F_{3/2}^2-c^4F_{3/2} 2.089 μm, that form in the outer wind and have small oscillator strengths (gf ~ 10^{-5}). The second are the weak permitted (gf ~ 1) lines connecting higher lying levels, such as the 4e^6G-5p^6F lines near 1.733 μm or 6p^6D-6s^6D at 2.109 μm, that form much closer to the photosphere. The permitted lines are more robust iron abundance indicators, having only weak dependencies on other parameters, such as turbulent velocity. The strengths of the semi-forbidden lines depend on the accuracy of their weak gf values, the mass-loss rate and the run of the iron ionization structure in the outer wind, which is sensitive to the hydrogen ionization structure due to the strong coupling to the Fe/H charge-exchange reactions. Since a change in the run of the clumping factor in the outer wind regions modifies the ratio of recombinations/ionizations in hydrogen, the semi-forbidden lines are diagnostic of the behavior of clumping there. From Figure 1 it can be seen that our model is able to simultaneously reproduce both sets of lines, providing constraints on both clumping and abundance.

We obtain roughly solar iron abundances for both LBVs, with ±0.15 dex as plausible uncertainties (see Figure 6). Our results are similar to A(Fe/H) = 7.59 recently derived by Cunha et al. (2007) from their analysis of a sample of luminous cool stars within 30 pc of the GC. Note in Table 1 that the Fe abundance ratio has significant uncertainty due to the uncertainty in the Fe abundance in the Sun.

Magnesium. The strongest Mg II lines observed in the H and K bands share the 5p^2P level. Those lines with it as the upper level, the 2.13/14 μm and 2.40/41 μm doublets (see Figures 1 and 2) are much stronger than those with it as the lower level (H band lines), revealing that pumping through the resonance 3s^2S-5p^2P line must be a significant populator of the 5p^2P levels. Pumping through the 3s^2S[1/2]-5p^2P[3/2] 1025.968 Å transition is very efficient due to Lyβ fluorescence. This was confirmed in models in which we decoupled the 5p^2P[3/2] and 5p^2P[1/2] levels (see Figure 7), resulting in Mg II 2.13/14 μm ratios much higher than observed.

The relevance of this process can be easily followed in Figure 7, which displays the behavior of the doublet as a function of the Mg II atom and the turbulent velocity. The latter refers to the fixed Doppler width used in our models to compute the level populations. In the left panel of Figure 7, the levels are considered to be decoupled (i.e., the number of superlevels in the model atom, NS, is set to the total number of levels in the full atom, NF). Because Lyβ lies closer to 3s^2S[1/2]-5p^2P[3/2] (Δν = 72 km s^{-1}) than to 5p^2P[1/2] - 5p^2P[3/2] (Δν = 3.5 km s^{-1}),
Figure 6. Error estimates of Fe (upper panel), Mg (middle panel) and Si (lower panel) abundances. Dashed lines (red) correspond to our best model fitting; the observed (black solid) diagnostic lines of FMM362. Long-dashed (green) and dashed–dotted (blue) lines correspond to models where individual metal abundances have been set to the derived upper and lower estimates respectively.

(A color version of this figure is available in the online journal.)

Due to fluorescence coupling, the Mg II K band lines show a stronger dependence on turbulent velocity than do the H band lines. We estimate about twice solar Mg abundance and an associated uncertainty (see Figure 6) of about $\pm0.25$ dex (due to uncertainties related to the fluorescence contribution).

Silicon. The Si II doublet $5s^2S_{1/2} - 5p^2P_{3/2}$ 1.691 $\mu$m and $5s^2S_{1/2} - 5p^2P_{1/2}$ 1.698 $\mu$m constitutes a powerful diagnostic tool, as it appears in emission for only a very narrow range of stellar temperatures and wind density structures, indicating the presence of amplified non-LTE effects. However, since it forms at the base of the wind, its strong dependence on the details of the velocity field there hinders a precise silicon abundance determination.

Instead, we use the well-behaved recombination line Si II $3s^26g^2G - 3s^25f^2F$ at 1.718 $\mu$m which shows a stronger dependence on the silicon abundance. Once again, a realistic mapping of full- to super-levels in our model atom is required. From our model fits (see Figures 1 and 2), we derive roughly...

3$s^2S_{1/2} - 5p^2P_{1/2}$ ($\Delta v = 114$ km s$^{-1}$), and because in these LBVS the terminal velocities and wind densities determine the line-formation zones, only the $5p^2P_{3/2}$ is pumped through fluorescence. Indeed, it can be seen that as the turbulence velocity is increased, the overlap between Ly$\beta$ and the Mg II line increases, as does the population of the $5p^2P_{3/2}$ level and the strength of the Mg II 2.13 $\mu$m line increases, while the longer wavelength Mg II 2.14 $\mu$m line is unaffected. On the other hand, if the Mg II model atom has both levels combined into a superlevel (NS$\neq$NF; Figure 7, right), the observed ratio is reproduced. Furthermore, increasing the turbulent velocity, and hence the effect of fluorescence, increases the pumping of both levels equally and thus increases the strength of the doublet with a constant ratio between its components. From Figure 7 it can be seen that our assumed collision coefficients connecting the Mg II $5p^2P_{1/2}$ and $5p^2P_{3/2}$ levels may be too low. This comparison illustrates the importance of making the correct choice of model atoms for quantitative spectroscopic analysis.
twice solar abundance (±0.20 dex) for silicon in each LBV, similar to magnesium (see Figure 6).

Other elements. One might expect a number of oxygen lines might be detectable in infrared spectra of LBVs, i.e., strong O i lines at 2.763 μm, 2.893 μm, and 3.098 μm and weaker lines at 1.8243 μm, 3.661 μm, and 3.946 μm. Several of these, but not all, are problematical from ground-based observatories. Unfortunately, our data set only encompasses the O i 1.745 μm line which is blended with a stronger Mg ii line. Thus, we defer an attempt to estimate the oxygen abundance until high resolution observations of unblended lines can be obtained. Determining the oxygen abundances in these objects will provide crucial constraints on their evolutionary status. Models show that when H/He < 1.50 oxygen has reached its maximum depletion within CNO equilibrium, while a significantly higher O content should be present on the stellar surface for H/He values around 3. The results in Table 1 then predict that the Pistol Star and FMM362 have different oxygen abundances. On the other hand, if an LBV has an H/He < 1.50 but is still hydrogen rich, the oxygen abundance determination, expected to be ∼0.04 of the original value, will provide a measure of the metallicity of the natal cloud. High resolution L band spectra of the Pistol Star should be able to address this issue.

The only detected sodium lines are the well known doublet at 2.206/9 μm, from which we obtain a very high abundance, ∼ 20× solar. The strong observed emission of this doublet in the K band spectra of other LBVs has been noted previously by Hillier et al. (1998). Interestingly, our models display only a minor dependence of these lines on clumping. On the other hand, the strengths of the sodium lines might not indicate extraordinary sodium abundance if the lines are produced by fluorescence of circumstellar material, a component that we do not model.

5. DISCUSSION

Our results suggest solar Fe abundances and approximately twice-solar α-element abundances for the Quintuplet LBVs. Presumably, these abundances were the same in the gas that condensed to form these stars and the other stars in the Quintuplet cluster and indeed in the whole of the present-day GC. The results can be discussed in the context of similar measurements of GC objects and with respect to the trend one might expect if the region is an inward extension of the disk or the bulge. In addition, the ratio of Fe to α-elements might be used to decipher the star-formation history in the GC.

Table 2 displays a number of recent determinations of stellar metal abundances in the GC together with the abovementioned three reference patterns for solar abundances. The values derived for Fe abundances in cool stars agree with our result (Carr et al. 2000; Ramirez et al. 1997, 1999, 2000). Cunha et al. (2007) find a very narrow range of Fe abundances clustered around the solar value for a population of cool stars in the central 30 pc. Of particular interest is star VR5-7 from Cunha et al. (2007), sample which is located in the Quintuplet cluster and shows A(Fe/H) = 7.60 and A(Ca/H) = 6.41.

There are relatively few measurements of the α-element abundances ([α/Fe]) in GC stars. Najarro et al. (2004) find solar abundances (as defined in this paper) for hot stars in the Arches cluster based on the oxygen abundance and, to a lesser degree, carbon abundance, and adopting the canonical solar value of A(O/H) = 8.93 (Anders & Grevesse 1989). Those estimates assume that nitrogen has reached its maximum surface abundance value. Evolutionary models indicate that 95% of that value is already attained by the time that H/He < 2 (by number). Najarro et al. (2004) followed the metallicity patterns from the Geneva evolutionary models and assumed no selective enrichment of CNO or α-elements versus Fe, in concluding that the stars in the Arches cluster have solar α-element abundances. However, estimates of solar abundances have varied considerably over the past 15 years (e.g., Allende Prieto 2008; see also Table 2). Thus, depending on the assumed solar CNO composition, the derived nitrogen abundance by Najarro et al. (2004) could imply solar (Anders & Grevesse 1989), 1.2 × solar (Grevesse & Noels 1993) or 2.0 × solar (Asplund et al. 2005) CNO composition.
Recently, Martins et al. (2007, 2008) have analyzed a larger sample of hot stars in the Arches and Central Parsec clusters and find similar results (see Table 2). Interestingly, if one considers only the objects in Martins et al. (2008) with He/H > 0.1 and those with Z(C) < 0.05, i.e., fulfilling the condition to be close enough to Z(N)$_{\text{max}}$, the average value of Z(N) is 1.7.

Geballe et al. (2006) estimate roughly solar oxygen abundance, A(O/H) = 8.91, in IRS 8, an Of supergiant near the central parsec. Cunha et al. (2007) find ⟨A(O/H)⟩ = 9.04 (0.37) and ⟨A(Fe/H)⟩ = 7.59 (0.14) for their sample of cool stars, where the numbers in parentheses are the ratio with respect to the solar value in dex. This implies [O/Fe] = 0.22, i.e., a clear enhancement over the solar ratio. Again, the Cunha et al. (2007) measurements could be interpreted as indicating solar ratios in O over Fe if the solar O abundance in evolutionary models is used. It is crucial to have accurate solar abundances, and that values used in stellar evolution calculations should be consistent with these. An excellent example is attempting to determine whether the possible oxygen enhancement is due to a top-heavy IMF favoring α-elements versus Fe enrichment, or simply an overall CNO and metal enhancement. Thus, taking the CNO abundances for the GC objects from Cunha et al. (2007) and assuming C/N equilibrium values, one can interpret their results either as solar CNO with mildly enhanced (30%) oxygen (Anders & Grevesse 1989) or a clearly supersolar environment with a factor of 1.7 enhancement for C and N and 2.5 for oxygen (Asplund et al. 2005).

Fortunately, there are other α-elements whose adopted solar abundances have suffered basically no major revision. Thus, we believe that the enhanced values obtained in this work for Mg and Si, roughly a factor of two solar, together with the enhancement of Ca found by Cunha et al. (2007), are a strong indication of the enrichment of α-elements compared to Fe.

Our results run counter to the trend in the disk (Rolleston et al. 2000; Smartt et al. 2001; Martín-Hernández et al. 2003), and are more consistent with the values found for the bulge (Frogel et al. 1999; Feltzing & Gilmore 2000). This may imply that the ISM in the disk does not extend inward to the GC, so that material is dragged into the central molecular zone from the bulge rather than from the disk. Another possibility is that the GC stars are forming out of an ISM that has an enrichment history distinctly different from that of the disk. At this point, further studies of the α-elements versus Fe would be useful. Future high S/N and high resolution spectroscopy of the O1 lines in LBVs and K band spectroscopy of WNL stars in the same cluster (F. Najarro et al. 2009, in preparation) will provide two independent measurements of the original oxygen content, and thus set definite constraints on metallicity.

The modest enrichment in α-elements versus Fe that we find in the two Quintuplet LBVs is consistent with a top-heavy IMF in the GC (Figer et al. 1999a). In such a scenario, enhanced yields of α-elements compared to Fe are expected through a higher than average ratio of the number of SNII versus SNIa events (Wheeler et al. 1989; Cunha et al. 2007).

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