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SPECTROPHOTOMETRY OF THE ORION NEBULA II. THE ABUNDANCE OF ELEMENTS

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ABSTRACT

Using a recent spectrophotometric study of the Orion Nebula, the chemical composition of this object is derived. The electron temperature is also given. The abundance by mass is X = 0.713; Y = 0.26 ± 0.02; Z = 0. 027 ± 0.004, which is very close to the so-called primordial abundance.

SUMARIO

Con los datos de un estudio espectrofotométrico reciente sobre la Nebulosa de Orion, la composición química de este objeto ha sido derivada. La temperatura electrónica también es presentada. La abundancia por masa es: X = 0.713; Y = 0.26 ± 0.02; Z = 0. 027 ± 0.004 lo cual es muy cercana a la abundancia primordial, derivada con la observación de neutrinos solares.

Introduction

Photoelectric spectrophotometry of the densest parts of the Orion Nebula has been recently presented by the author (1967; hereafter referred to as Paper I). The spectral range covered involves a good number of lines, which lie between λ3737 and λ10830, with intensities sufficiently high as to permit accurate photometric measurements. If one accepts that the Balmer and Paschen decrements for the lower member of the series are accurately described by the Case B recombination theory, one can then safely conclude that the intensities of the lines, given in Paper I, are properly corrected for interstellar reddening. Consequently, the spectral lines observed can be employed in the determination of the ionic concentration of several elements.

The problem of abundance determination in a gaseous nebula is in principle simpler than the corresponding one in stellar atmospheres. The ionic concentration in such a low density medium is directly proportional to the line intensity. If the relevant atomic parameters are known, the constant of proportionality can be computed. For some elements one can expect better abundance determinations than those obtained through analysis of stellar spectra; one can give as an example the helium to hydrogen ratio. Unfortunately, for some of the elements, which produce forbidden line radiation, the relevant collisional cross sections are not known with sufficient precision. In some cases, approximates must be made introducing thus uncertainties. As an additional difficulty, the radiation field in the nebula can not be determined with accuracy; therefore the ionization equilibrium equation, needed to weight the contribution of the various stages of ionization, is not known explicitely. Some elements, in certain stages of ionization do not show emission lines in the observable region. In those cases, the total abundance has to be computed byindirect means, which might also introduce errors (Bowen and Wyse 1939). The density fluctuations in the Orion Nebula are also associated with temperature fluctuations. These inhomogeneities necessarily introduce stratification: for instance, the OIII forbidden lines are produced predominantly in the hot regions and the SII lines in the dense zones. It must be concluded that there is an averaging, along the line of sight, in the measurement of the intensities of the lines, which is different for each stage of ionization.

The work presented here is based on the spectrophotometry of region 1, of Paper I. This zone is located in the densest part of the nebula and constitutes a single very bright condensation. Since the area covered by the diaphagm is a circle of 14 seconds of arc in diameter, there are not large fluctuations in the directions perpendicular to the line of sight, difficulty which is present in other investigations.

II. The observations

The photoelectric observations, used in this work. have already been described in Paper I. The line intensities listed there do not go fainter than the intensity 2 in the list published by Flather and Osterbrock (1960). In fact, a limit for the lowest intensity to be measured photoelectrically was set accordingly. The lines fainter than that limit were measured using photographic methods. In addition to the spectral resolution improvement (the photoelectric resolution was set at 10 Å), a much higher definition of the observed areas is attained by photographic means. This higher precision is required if one desires to have an estimate of the amplitude of temperature fluctuations, as indicated by the [OIII] line intensity ratios.

Most of the spectrograms were obtained using the Newtonian B-spectrograph. This instrument was used with the 100-inch telescope and with the 60-inch as well, at the Mount Wilson Observatory. The 88 B grating, with 600 lines/mm. in combination with the 3-inch Schmidt camera, gave the dispersion of 85 A/mm. The Cassegrain spectrograph of the 40-inch telescope at Tonantzintla was used in a small fraction of the program. The f/1.4 camera, with a 400 lines/mm grating, yielded spectrograms with dispersion of 125 A/mm. in the blue.

The measurement of the photographic line intensities were obtained by the usual tracing and calibration procedures. These processes were partly carried out with the recording microdensitometer at the University of Mexico. The automatic microphotometer at the California Isstitute of Technology was also used.

TABLE 1

Line wavelength	*Transition*	*Theory*	*Observed*
*5876	$3^3D - 2^3P$	10	10
4471	$4^3D - 2^3P$	4.2	4.1
4388	$5^1D - 2^1P$	0.63	0.5
4026	$5^3D - 2^3P$	2.3	2.4

* normalized to 10

The spectrogram reduction is a lenghty process; however, the work was simplified by the use of the results obtained with the photoelectric measurements. Table 1 gives a comparison of the intensity of some HeI lines observed photoelectrically, with those predicted by the pure recombination theory (Seaton 1960). As it can be noticed, the agreement seems to be excellent. Consequently, it can be safely assumed that the line intensities of HeI, in the nebular spectrum, can be accurately described by the theoretical recombination intensities. Therefore, the line intensities of other elements can be measured with respect to a nearby HeI line of about the same intensity. The accuracy attained obviously increases when the reference He I line is close to the line to be measured. The intensity of the important line of [OIII] at λ4363, for instance, can be easily measured with respect to λ4388 with the errors being reduced to a minimum. On the other hand, the intensities of the N_1 and N_2 lines can be obtained in terms of H_β with good precision, since the three lines are of comparable intensity. Because the intensity of any HeI line can be obtained in units of H_β, the $4363/N_1+N_2$ ratio is readily computed.

The relevant lines produced by other elements were measured in the same way, although in some cases the quality of the intensity determinations is not as good as in the case of λ4363. The nebular reddening correction was included; when it was significant, the values found for region 1

TABLE 2

Slit	$I(4363/(H_\beta) \times 100$	r	T_e (°K)
a	1.1	185	10100
b	1.8	215	9700
c	1.5	246	9300
d	2.0	245	9300
e	1.8	224	9600
f	1.6	210	9700
g	1.5	224	9600
h	1.4	240	9400
i	1.3	285	8900
j	0.9	175	10200
k	1.0	190	10000
l	1.2	170	10300
m	1.7	280	9000
n	1.8	270	9100
p	1.9	225	9600
q	1.0	190	10000
n	1.8	270	9100
p	1.9	225	9600
s	1.4	230	9500
t	1.3	225	9600

were always used. Since the wavelength intervals are small no appreciable error is introduced with this procedure.

Several regions were observed in order to determine the electron temperature. Figure 1, of Paper I, shows with letters the location of the regions studied.

III. The electron temperature

Any intensity ratio of the forbidden lines arising from term 3 to those originating from term 2 would provide a relation between the electron density and the temperature. Due to the stratification problem, the same element lines should be used for both determinations, that of density and of electron temperature. Although that is true for most of the ions producing forbidden lines, the OIII ion is a fortunate exception; for the electron densities prevailing in the Orion Nebula, the calculation of the temperature using the OIII lines, is independent of density. The ratio r, defined by I (5007 +4959)/I (4363), was obtained from several points in nebula. From these values, one can derive the electron temperature in those regions, using the well known formulae. The new target area parameters of OIII computed by Billings et al. (1967) have been employed. Table 2 gives the results.

Although one can notice that there are variations in the electron temperature, the values obtained agree satisfactorily with previous similar determinations (Aller and Liller 1959). However, most of past investigations refer only to the brighest region. Peimbert (1967) carried out photoelectric observations in three different regions and his $(N_1 + N_2)/4363$ ratios could be compared with the numbers in Table 2. His observations were obtained using a very long slit (120″) thus, no direct agreement can be expected. Nevertheless, is can be seen that the tendency of the temprature to increase outward, clearly shown in Table 2, is also present in Peimbert's result.

The width and structure of some emission lines indicate the existence of mechanical mass motions besides those of thermal origin. Wilson et al. (1959) adopting a kinetic temperature of 10^4 °K derived, from the line profiles, the r.m.s velocities of mass motion. Their results indicate that the ratio of the mass motion velocity to the velocity of sound is mostly less than one. One has to bear in mind this fact when lower kinetic temperatures are demanded, because higher turbulent velocities for mass motions are consequently required to explain the line widths observed in the Orion Nebula. With the temperatures obtained in this work, the mass motion velocities derived still remain below the speed of sound. Lower temperatures would indicate the existence of supersonic velocities, and then the mass motions should be rapidly damped, dissipating energy.

IV. The abundance determination

a) The Helium Content.

In paper I, the intensity of 6 lines of the triplet system of HeI were presented. However, only some of those lines are suitable for abundance calculations. The reason being the metastability of the 2^3S level, which introduces appreciable changes on the pure recombination spectrum through self-absorption transitions. An overpopulation of that level necessarily leads to self-absorption of those lines, which have the 2^3S state as lower level. Inthis way, some upper triplet levels are populated, from where some other lines, of larger wavelength, may be produced. In conclusion, for the HeI spectrum the effects of self-absorption must be taken into account in addition to the recombination processes.

The He/H abundance ratio can be determined using the strongest lines of HeI, which are little effected by self-absorption. Block (1967), using a 14 level He atom, carried out self-absorption computations and determined the relative intensities of some lines in the triplet system, as a function of the optical depth at the λ3889 line. From his data, and from the spectrophotometry of the 6 helium lines observed, one concludes that the intensity of the λ5876 line is increased over the recombination value, by about 7 per cent. The $5876/H_\beta$ intensity ratio is 0.098 for region 1, that is, 0.091 due to pure recombination. Using the recombination coefficients as computed by Seaton (1959) and by Burgess and Seaton (1960), one derives

$$\frac{N(He^+)}{N(H^+)} = 0.97\,\frac{I(5876)}{I(H_\beta)} = 0.088 \qquad (1)$$

and also

$$\frac{N(He^+)}{N(H^+)} = 2.39\,\frac{I(4471)}{I(H_\beta)} = 0.102$$

giving a weight of 2 to the first ratio, the mean is 0.093 ± 0.005.

b) The Heavier Elements.

The various lines, to be used for the determination of heavier element abundances arise mostly from transitions among the three terms of the ground state configuration namely p^2, p^3 or p^4. Let the three terms be denoted by numbers 1, 2 and 3, in order of increasing energy. The population of each term depends on collisional cross-sections and optical transition probabilities. If the latter quantities are provided, the population can be computed as a function of electron temperature and density in the following manner.

The electron density for the region considered is close to 10^4; this relatively low value permits us to make certain simplifications to the problem of abundance determinations from forbidden line emission. With such conditions, one can neglect the existence of term 3. An atom of just two terms needs to be considered. The emission per unit volume of any line can be expressed simply by:

$$E = N_2 A_{21} h\nu_{21} = N_1 p\, q_{12} h\nu_{21} \qquad (3)$$

where p is the factor, which takes into account the depopulation rate of term 2, both due to the second kind collisions and to radiative transitions from that term. The collisional excitation rate q_{12} is given by:

$$q_{12} = 9.6 \times 10^{-8} \frac{N_e \Omega\,(1.2)}{\omega} e^{-\frac{1.6}{\lambda}} \qquad (4)$$

here Ω (1,2) is the collisional strength and ω is the statistical weight of the term; λ is expressed in microns. Since the H_β-emission per unit volume is equal to $\alpha_{4,2}\ (Te)\ N(H^+) N_e\ h\nu_\beta$, where the effective recombination coefficient $\alpha_{4,2}\ (Te)$ can be obtained from the computations done by Seaton (1959). Therefore, for an ion T, which produces the emission line (s) at λ with an intensity $r(\lambda)$, in units of $I\ (H_\beta)$, the relative abundance is given by

$$\frac{N(T)}{N(H^+)} = 6.5 \times 10^{-7} \frac{\lambda}{p\Omega} r(\lambda)\, \omega\, e^{\frac{1.6}{\lambda}} \qquad (5)$$

TABLE 3

ION	λ	$r(\lambda) \times 100$	Ω (1,2)	$N(T)/N(H) \times 10^5$
C II	4267	0.43	rec	90.0
N II	6583	51.0	3.13	1.2
O II	3726-29	110	1.43	10.0
O III	5007	360	2.39	18.0
Ne III	3868	20	1.26	6.0
Si II	3856	0.10	rec	5.0
S II	6716-30	4.0	1.87	.02
S III	9069	32	4.97	1.3
A III	7135	8.0	4.75	0.5
A IV	4740	0.15	1.43	.004
Fe II	4287	0.20	1.00	.08
Fe III	4658	0.90	0.27	.60

The relevant data used in the calculations are given in Table 3, where the permitted lines, produced by recombination, λ4267 of CII and λ3856 of SiII, are also included.

The collisional strengths are not known for FeIII nor for FeII; some estimates must be made of those quantities. For FeIII, the relevant line λ4658 arises from the transition $a^3F_4 - a^5D_4$. This transition occurs by both magnetic-dipole and electric quadrupole radiation. It involves a spin change collision and the excitation can only occur by electron exchange, that is through the process:

$$A + 1 \rightarrow B + 1' \qquad (6)$$

where 1 and 1′ are the angular momentum quantum numbers of the free electron before and after the collision. The conservation theorem due to Mohr, Peierls and Plasczek, which states that

$$\sum_i \frac{\Omega\,(i,j)}{\omega_i} = (21 + 1) \qquad (7)$$

is specially useful for setting an upper limit to the collision strengths. Moreover van Regemorter (1962) has shown that Ω (i,j) is approximately proportional to the corresponding quadrupole transition strengths, which in turn are directly proportional to the spontaneous transition probabilities. Garstang (1957) has carried out extensive computations for the FeII ion, and the A_q, s are provided by his work. There are 12 lines arising between the $^3F_4$ and $^5D_4$ terms; consequently, through the corresponding optical transition probabilities, one can weight the various Ω's. By assuming that the main contribution to the collisions is produced by *s* and *p* electrons, the value of Ω, given in Table 3, was derived. This number sets a lower limit on the abundance of FeIII. As the ionization potential of FeII is 16.2 ev., one can expect only small contributions from this stage of ionization, as compared with FeIII. Consequently, the value of Ω is not very significant and has been arbitrarily set at unity.

The contribution from ions, which do not show any lines, needs to be estimated using the procedure established by Bowen and Wyse (1939). Actually, this method is described by Aller and Liller (1959) for the Orion Nebula itself. A similar empirical ionization curve was obtained using the relative abundance data for OII, OIII, SII, SIII, AIII and AIV.

The lines of CII and SiII require the knowledge of the effective recombination coefficients. The value of these quantities can be estimated from the tabulation of those coefficients given by Allen (1963). The following values were adopted as upper limits (which in turn would set lower limits for the abundances):

$$\alpha\ (4^2F - 3^2D) = 1.1 \times 10^{-13} \qquad \text{for CII}$$

$$\alpha\ (4^2P - 3^2D) = 0.3 \times 10^{-13} \qquad \text{for SiII}$$

the data derived from the latter numbers refers obviously to the abundances of CIII and SiIII, respectively.

After taking in account all the significant ionization stages, the total abundances can be computed. Table 4 shows the results obtained for the chemical composition of the Orion Nebula. There is a remarkable agreement with the abundance derived from the stars in the vicinity. Magnesium has thus been tentatively included using the spectral analysis of 22 Ori (Aller, 1960). Table 4 also gives the corresponding solar abundance as derived by Aller et al. (1960).

The major discrepancy in the comparison, shown in Table 4, is that of the carbon content, which is about 4 times higher than for 22 Ori. Aller and Liller (1959) using a different method of computation derived about the same value, as an upper limit. Another uncertain quantity is the relative abundance of NeII. According to the ionization curve the most abundant stage is NeII, which is not observable. The probable error of the result is of the order of 30%.

TABLE 4

Element	*Orion Neb.*	*Sun*	22 *Ori*
H	10^6	10^6	10^6
He	93000	——	——
C	1000	540	540
N	90	100	60
O	320	900	400
Mg	90	25	115
Ne	750	——	——
Si	90	32	85
S	16	20	25
A	6	——	——
Fe	7	3.2	——

V. Discussion

Although the Orion Nebula is an extremely young object as compared with the Sun, it is of interest to compare the chemical composition derived here with the so-called primordial composition obtained for the Sun. From the data given in Table 4, the abundance by mass can be computed, with the results:

$$X = 0.711 \ ; \ Y = 0.262 \pm 0.020 \ ; \ Z = 0.027 \pm 0.004$$

The solar-neutrino detection experiment has provided a powerful tool for testing the accuracy of theoretical models of the internal structure of the Sun. More specifically, the neutrino flux observed from the Sun is a function of the initial solar composition. With a set accurate models of the present Sun, Bahcall and Shaviv (1968) have obtained the primordial composition by mass:

$$X = 0.713 \; ; \; Y = 0.26 \pm 0.02 \; ; \; Z = 0.027 \pm 0.004$$

Gaustad (1964) has also derived $X = 0.72$; $Y = 0.26$ $Z = 0.02$ for the Sun using the abundance data of *He, C, N,* and *Ne* from cosmic rays produced by solar flares. However the Ne/O ratio quoted by Gaustad is 0.1, while the corresponding value for Orion is much larger (about 2.3). For the $He/C + N + O$) ratio, Sears (1964) has derived a value of 64 as compared with 66 for this work. These numbers are not inconsistent with the rocket observations of 54 ± 6, obtained by Gaustad. The conclusion to be derived from the above comparisons is that the helium content of the interstellar medium has not increased in the past $(4.5 \pm 0.2) \times 10^9$ years.

However, the experiments of Davis et al. (1968), designed to detect solar neutrinos based upon the neutrino capture reaction Cl^{37} (ν, e^-) A^{37} have given important conclusions regarding the heavy element abundance. Bahcall et al. (1968), using the observed neutrino-capture rate, conclude that the value of Z is about 0.015 and the primordial helium content is 0.22 ± 0.03 by mass.

If these results are correct the Orion Nebula, a typical representative of the present-day interstellar matter, has a chemical composition very close to the primordial one and that the stars, which produce helium and heavy elements, have not returned much of the former to space, but they have certainly duplicated the primordial abundance of heavier elements.

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