Experimental and theoretical nuclear astrophysics: the quest for the origin of the elements*

William A. Fowler

W. K. Kellogg Radiation Laboratory, California Institute of Technology, Pasadena, California 91125

Ad astra per aspera et per ludum

I. INTRODUCTION

We live on planet Earth warmed by the rays of a nearby star we call the sun. The energy in those rays of sunlight comes initially from the nuclear fusion of hydrogen into helium deep in the solar interior. Eddington told us this in 1920, and Hans Bethe developed the detailed nuclear processes involved in the fusion in 1939. For this he was awarded the Nobel Prize in Physics in 1967.

All life on Earth, including our own, depends on sunlight and thus on nuclear processes in the solar interior. But the sun did not produce the chemical elements which are found in the Earth and in our bodies. The first two elements and their stable isotopes, hydrogen and helium, emerged from the first few minutes of the early high-temperature, high-density stage of the expanding universe, the so-called "big bang." A small amount of lithium, the third element in the Periodic Table, was also produced in the big bang, but the remainder of the lithium and all of beryllium, element four, and boron, element five, are thought to have been produced by the spallation of still heavier elements by the cosmic radiation in the interstellar medium between stars. These elements are in general very rare, in keeping with this explanation of their origin, as reviewed in detail by Audouze and Reeves (1982).

Where did the heavier elements originate? The generally accepted answer is that all of the heavier elements from carbon, element six, up to long-lived radioactive uranium, element ninety-two, were produced by nuclear processes in the interior of stars in our own Galaxy. The stars we see at the present time in what we call the *Milky Way* are located in a spiral arm of our Galaxy. In Sweden you call it *Vinter Gatan,* the Winter Street. We see with our eyes only a small fraction of the one hundred billion stars in the Galaxy. Astronomers cover almost the full range of the electromagnetic spectrum and thus can observe many more Galactic stars and even individual stars in other galaxies.

The stars which synthesized the heavy elements in the solar system were formed or born, evolved or aged, and eventually ejected the ashes of their nuclear fires into the interstellar medium over the lifetime of the Galaxy *before* the solar system itself formed four and one-half billion years ago.

The lifetime of the Galaxy is thought to be more than ten billion years but less than twenty billion years. In any case the Galaxy is much older than the solar system. The ejection of the nuclear ashes or newly formed elements took place by slow mass loss during the old age of the star, called the *giant* stage of stellar evolution, or during the relatively frequent outbursts which astronomers call *novae,* or during the final spectacular stellar explosions called *supernovae.* Supernovae can be considered to be the death of stars, but the white dwarfs or neutron stars or black holes that they leave behind may represent a form of stellar purgatory.

In any case the sun and the Earth and all the other planets in the solar system condensed under gravitational and rotational forces from a gaseous solar nebula in the interstellar medium consisting of "big bang" hydrogen and helium mixed with the heavier elements synthesized in earlier generations of Galactic stars. All of this is illustrated in Fig. 1.

This idea can be generalized to successive generations of stars in the Galaxy with the result that the heavy-element content of the interstellar medium and the stars which form it increases with time. The oldest stars in the Galactic halo, that is, those we believe to have formed first, are found to have heavy-element abundances less than one percent of the heavy-element abundance of the solar system. The oldest stars in the Galactic disk have approximately ten percent. Only the less massive stars

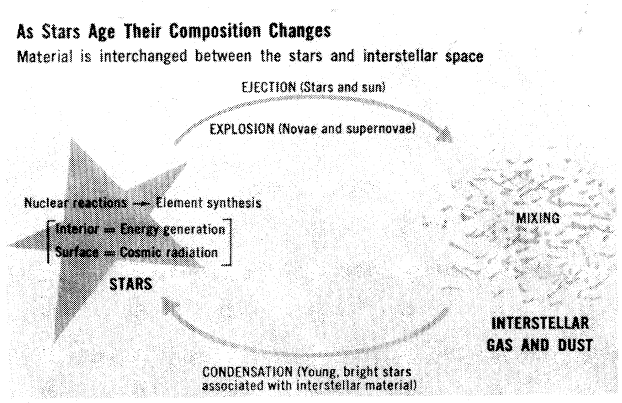

FIG. 1. Synthesis of the elements in stars.

*This lecture was delivered December 8, 1983, on the occasion of the presentation of the 1983 Nobel Prize in Physics.

among those first formed can have survived to the present as so-called Population II stars. Their small concentration of heavy elements may have been produced in a still earlier but more massive generation of stars, Population III, which rapidly exhausted their nuclear fuels and survived for only a very short lifetime. Stars formed in the disk of the Galaxy over its lifetime are referred to as Population I stars.

We speak of this element building as nucleosynthesis in stars. It can be generalized to other galaxies such as our twin, the Andromeda Nebula, and so this mechanism can be said to be a universal one. Astronomical observations of other galaxies have contributed much to our understanding of nucleosynthesis in stars.

We refer to the basic physics of energy generation and element synthesis in stars as Nuclear Astrophysics. It is a benign application of nuclear physics, in contrast to reactors and bombs. For the nuclear physicist this contrast is a personal and professional paradox. However, there is one thing of which I am certain. The science which explains the origin of sunlight must not be used to raise a dust cloud which will black out that sunlight from our planet.

As for all physics, the field of Nuclear Astrophysics involves experimental and theoretical activities on the part of its practitioners, and hence the first part of the title of this lecture. This lecture will emphasize nuclear experimental results and the theoretical analysis of those results almost, but not entirely, to the exclusion of other theoretical aspects. It will not in any way do justice to the observational activities of astronomers and cosmochemists, which are necessary to complete the cycle: experiment, theory, observation. Nor will it do justice to the calculations by many theoretical astrophysicists of the results of nucleosynthesis of the elements and their isotopes under astrophysical conditions during the many stages of stellar evolution.

My deepest personal interest is in experimental data, in the analysis of the data, and in the proper use of the data in theoretical stellar models. I continue to be encouraged in this regard by this one-hundred-and-ten-year-old quotation from Mark Twain:

> There is something fascinating about science. One gets such wholesale returns of conjecture out of such a trifling investment of fact.
>
> *Life on the Mississippi,* 1874

For me Twain's remark is a challenge to the experimentalist. The experimentalist must try to eliminate the word "trifling" through his endeavors in uncovering the facts of nature.

Experimental research and theoretical research are often very hard work. Fortunately this is lightened by the fun of doing physics and in obtaining results which bring a personal feeling of intellectual satisfaction. To my mind the hard work and the resulting intellectual fun transcend in a way the benefits which may accrue to society through subsequent technological applications. Please understand—I do not belittle these applications, but I am unable to overlook the fact that they are a two-edged sword. My subject matter resulted from the hard work of a nuclear astrophysicist which when successful brought him joy and satisfaction. It was hard work but it was fun. Thus I have chosen the subtitle for this lecture—"*Ad astra per aspera et per ludum,*" which can be freely translated "To the stars through hard work and fun." This is in keeping with my paraphrase of the biblical quotation from Matthew, "Man shall not live by work alone."

With that in the record let us next ask what are the goals of Nuclear Astrophysics? First of all, Nuclear Astrophysics attempts to understand energy generation in the sun and other stars at all stages of stellar evolution. Energy generation by nuclear processes requires the transmutation of nuclei into new nuclei with lower mass. The small decrease in mass is multiplied by the velocity of light squared, as Einstein taught us, and a relatively large amount of energy is released.

Thus the first goal is closely related to the second goal, that attempts to understand the nuclear processes which produced, under various astrophysical circumstances, the relative abundances of the elements and their isotopes in nature; whence the second part of the title of this lecture. Figure 2 shows a schematic curve of atomic abundances as a function of atomic weight. The data for this curve were first systematized from a plethora of terrestrial,

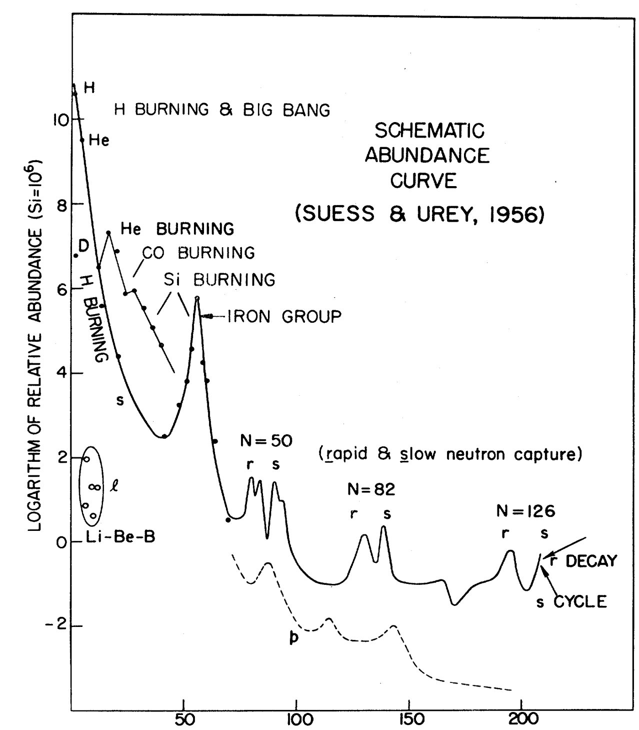

FIG. 2. Schematic curve of atomic abundances relative to $Si=10^6$ vs atomic weight for the sun and similar main sequence stars.

meteoritic, solar, and stellar data by Hans Suess and Harold Urey (1959), and the available data have been periodically updated by A. G. W. Cameron (1982). Major contributions to the experimental measurement of atomic transition rates needed to determine solar and stellar abundances have been made by my colleague, Ward Whaling (1982). The articles by Cameron (1982) and Whaling (1982) occur in a book, *Essays in Nuclear Astrophysics,* which reviews the field up to 1982. In the words of one of America's baseball immortals, Casey Stengel, "You can always look it up."

The curve in Fig. 2 is frequently referred to as "universal" or "cosmic," but in reality it primarily represents relative atomic abundances in the solar system and in main sequence stars similar in mass and age to the sun. In current usage the curve is described succinctly as "solar." It is beyond the scope of this lecture to elaborate on the difficult, beautiful research in astronomy and cosmochemistry which determined this curve. How this curve serves as a goal can be simply put. In the sequel it will be noticed that calculations of atomic abundances produced under astronomical circumstances at various postulated stellar sites are almost invariably reduced to ratios relative to "solar" abundances.

II. EARLY RESEARCH ON ELEMENT SYNTHESIS

George Gamow and his collaborators, R. A. Alpher and R. C. Herman (1950), attempted to synthesize all of the elements during the big bang using a nonequilibrium theory involving neutron (n) capture with gamma-ray (γ) emission and electron (e) beta decay by successively heavier nuclei. The synthesis proceeded in steps of one mass unit at a time, since the neutron has approximately unit mass on the mass scale used in all the physical sciences. As they emphasized, this theory meets grave difficulties beyond mass 4 (^{4}He) because no stable nuclei exist at atomic mass 5 and 8. Enrico Fermi and Anthony Turkevich attempted valiantly to bridge these "mass gaps" without success and permitted Alpher and Herman to publish the results of their attempts. Seventeen years later Wagoner, Fowler, and Hoyle (1967), armed with nuclear reaction data accumulated over the intervening years, succeeded only in producing ^{7}Li at a mass fraction of at most 10^{-8} compared to hydrogen plus helium for acceptable model universes. All heavier elements totaled less than 10^{-11} by mass. Wagoner, Fowler, and Hoyle (1967) did succeed in producing ^{2}D, ^{3}He, ^{4}He, and ^{7}Li in amounts in reasonable agreement with observations at the time. More recent observations and calculations are frequently used to place constraints on models of the expanding universe and in general favor open models in which the expansion continues indefinitely.

It was in connection with the "mass gaps" that the W. K. Kellogg Radiation Laboratory first became involved, albeit unwittingly, in astrophysical and cosmological phenomena. Before proceeding, it is appropriate at this point to discuss briefly the origins of the Kellogg Radiation Laboratory, where I have worked for 50 years. The laboratory was designed and the construction supervised by Charles Christian Lauritsen in 1930 through 1931. Robert Andrews Millikan, the head of Caltech, acquired the necessary funds from Will Keith Kellogg, the American "corn flakes king." The Laboratory was built to study the physics of 1-MeV x rays and the application of those x rays in the treatment of cancer. In 1932 Cockcroft and Walton discovered that nuclei could be disintegrated by protons (p), the nuclei of the light hydrogen atom ^{1}H, accelerated to energies well under 1 MeV. Lauritsen immediately converted one of his x-ray tubes into a positive ion accelerator (they were powered by alternating current transformers!) and began research in nuclear physics. Robert Oppenheimer and Richard Tolman were instrumental in convincing Millikan that Lauritsen was doing the right thing. Oppenheimer played an active role in the theoretical interpretation of the experimental results obtained in the Kellogg Laboratory in the early crucial years.

Lauritsen supervised my doctoral research from 1933—1936, and I worked closely with him until his death. It was he who taught me that physics was both hard work and fun. He was a native of Denmark and was an accomplished violinist as well as physicist, architect, and engineer. He loved the works of Carl Michael Bellman, the famous Swedish poet-musician of the 18th century, and played and sang Bellman for his students. It is well known that many of Bellman's works were drinking songs. That made it all the better.

We must now return to the first involvement of the Kellogg Radiation Laboratory in the mass gap at mass 5. In 1939, in Kellogg, Hans Staub and William Stephens (1939) detected resonance scattering by ^{4}He of neutrons with orbital angular momentum equal to one in units of $\hbar$ (p-wave) and energy somewhat less than 1 MeV, as shown in Fig. 3. This confirmed previous reaction studies by Williams, Shepherd, and Haxby (1937) and showed that

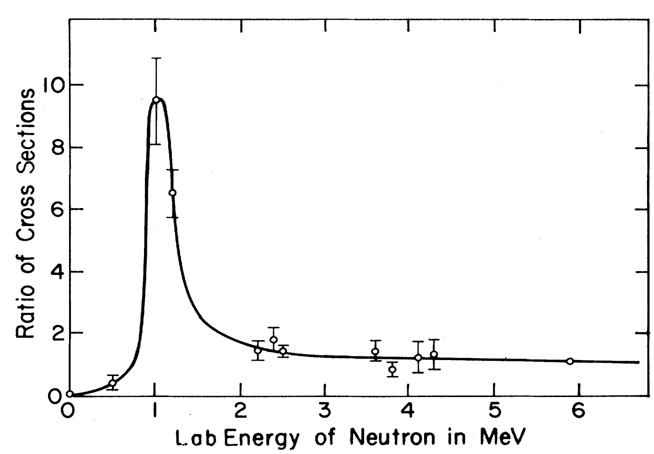

FIG. 3. The ratio of the backward scattering cross section of helium to hydrogen as a function of the laboratory energy in MeV of the incident neutron.

the ground state of ^{5}He is unstable. As fast as ^{5}He is made it disintegrates! The same was later shown to be true for ^{5}Li, the other candidate nucleus at mass 5. The Pauli exclusion principle dictates for fermions that the third neutron in ^{5}He must have at least unit angular momentum and not zero, as permitted for the first two neutrons with antiparallel spins. The attractive nuclear force cannot match the outward centrifugal force in classical terminology. Still later, in the Kellogg Radiation Laboratory, Tollestrup, Fowler, and Lauritsen (1949) confirmed, with improved precision, the discovery of Hemmendinger (1948,1949) that the ground state of ^{8}Be is unstable. They found the energy of the ^{8}Be breakup to be 89 ± 5 keV, compared to the currently accepted value of 91.89 ± 0.05 keV! The Pauli exclusion principle is again at work in the instability of ^{8}Be. As fast as ^{8}Be is made it disintegrates into two ^{4}He nuclei. The latter may be bosons, but they consist of fermions. The mass gaps at 5 and 8 spelled the doom of Gamow's hopes that all nuclear species could be produced in the big bang, one unit of mass at a time.

The eventual commitment of the Kellogg Radiation Laboratory to Nuclear Astrophysics came about in 1939, when Bethe (1939; see also Bethe, 1967) brought forward the operation of the CN cycle as one mode of the fusion of hydrogen into helium in stars (since oxygen has been found to be involved, the cycle is now known as the CNO cycle). Charles Lauritsen, his son Thomas Lauritsen, and I were measuring the cross sections of the proton bombardment of the isotopes of carbon and nitrogen which constitute the CN cycle. Bethe's paper (1939) told us that we were studying in the laboratory processes which are occurring in the sun and other stars. It made a lasting impression on us. World War II intervened, but in 1946 on returning the laboratory to nuclear experimental research, Lauritsen decided to continue in low-energy, *classical* nuclear physics, with emphasis in the study of nuclear reactions thought to take place in stars. In this he was strongly supported by Ira Bowen, a Caltech Professor of Physics who had just been appointed Director of the Mt. Wilson Observatory, by Lee DuBridge, the new President of Caltech, by Carl Anderson, Nobel Prize winner 1936, and by Jesse Greenstein, newly appointed to establish research in astronomy at Caltech. In Kellogg, Lauritsen did not follow the fashionable trend to higher and higher energies which has continued to this day. He did support Robert Bacher and others in establishing high-energy physics at Caltech.

Although Bethe in 1939 and others still earlier had previously discussed energy generation by nuclear processes in stars, the grand concept of nucleosynthesis in stars was first definitely established by Fred Hoyle (1946,1954).[1] In two classic papers the basic ideas of the concept were presented within the framework of stellar structure and evolution, with the use of the then known nuclear data.

Again the Kellogg Laboratory played a role. Before his second paper Hoyle was puzzled by the slow rate of the formation of ^{12}C nuclei from the fusion ($3\alpha\rightarrow{}^{12}$C) of three alpha particles (α) or ^{4}He nuclei in red giant stars. Hoyle was puzzled because his own work with Schwarzschild (Hoyle and Schwarzschild, 1955) and previous work of Sandage and Schwarzschild (1952; in particular, see last paragraph on p. 475) had convinced him that helium burning through $3\alpha\rightarrow{}^{12}$C should commence in red giants just above 10^8 K rather than at 2×10^8 K as required by the reaction rate calculation of Salpeter (1952a). Salpeter made his calculation while a visitor at the Kellogg Laboratory during the summer of 1951 and used the Kellogg value (Tollestrup *et al.*, 1949) for the energy of ^{8}Be in excess of two ^{4}He to determine the resonant rate for the process ($2\alpha\leftrightarrow{}^8$Be), which takes into account both the formation and decay of the ^{8}Be. However, in calculating the next step, $^8\mathrm{Be}+\alpha\rightarrow{}^{12}\mathrm{Ca}+\gamma$, Salpeter had treated the radiative fusion as nonresonant.

Hoyle realized that this step would be speeded up by many orders of magnitude, thus reducing the temperatures for its onset, if there existed an excited state of ^{12}C with energy 0.3 MeV in excess of $^8\mathrm{Be}+\alpha$ at rest and with the angular momentum and parity $(0^+,1^-,2^+,3^-,\ldots)$ dictated by the selection rules for these quantities. Hoyle came to the Kellogg Laboratory early in 1953 and questioned the staff about the possible existence of his proposed excited state. To make a long story short, Ward Whaling and his visiting associates and graduate students (Dunbar *et al.*, 1953) decided to go into the laboratory and search for the state using the $^{14}\mathrm{N}(d,\alpha)\,^{12}\mathrm{C}$ reaction. They found it to be located almost exactly where Hoyle had predicted. It is now known to be at 7.654-MeV excitation in ^{12}C or 0.2875 MeV above $^8\mathrm{Be}+\alpha$ and 0.3794 MeV above 3α. Cook, Fowler, Lauritsen, and Lauritsen (1957) then produced the state in the decay of radioactive ^{12}B and showed it could break up into 3α and thus by reciprocity could be formed from 3α. They argued that the spin and parity of the state must be 0^+, as is now known to be the case.

The $3\alpha\rightarrow{}^{12}$C fusion in red giants jumps the mass gaps at 5 and 8. This process could never occur under big bang conditions. By the time the ^{4}He was produced in the early expanding universe the subsequent density and temperature were too low for the helium fusion to carbon to occur. In contrast, in red giants, after hydrogen conversion to helium during the main sequence stage, gravitational contraction of the helium core raises the density and temperature to values where helium fusion is ignited. Hoyle and Whaling showed that conditions in red giant stars are just right.

Fusion processes can be referred to as nuclear burning in the same way we speak of chemical burning. Helium burning in red giants succeeds hydrogen burning in main sequence stars and is in turn succeeded by carbon, neon, oxygen, and silicon burning to reach to the elements near iron and somewhat beyond in the Periodic Table. With these nuclei of intermediate mass as seeds, subsequent processes similar to Gamow's involving neutron capture

[1]For a discussion of earlier ideas, including suggestions and retractions, see Chap. XII of Chandrasekhar (1939).

at a slow rate (s process) or at a rapid rate (r process), continued the synthesis beyond ^{209}Bi, the last stable nucleus, up through short-lived radioactive nuclei to long-lived ^{232}Th, ^{235}U, and ^{238}U, the parents of the natural radioactive series. This last requires the r process, which actually builds beyond mass 238 to radioactive nuclei which decay back to ^{232}Th, ^{235}U, and ^{238}U rapidly at the cessation of the process.

The need for two neutron capture processes was provided by Suess and Urey (1956). With the adroit use of relative isotopic abundances for elements with several isotopes, they demonstrated the existence of the double peaks (r and s) in Fig. 2. It was immediately clear that these peaks were associated with neutron shell filling at the magic neutron numbers $N=50$, 82, and 126 in the nuclear shell model of Hans Jensen and Maria Goeppert-Mayer, who won the Nobel Prize in Physics just twenty years ago.

In the s process the nuclei involved have low capture cross sections at shell closure and thus large abundances to maintain the s-process flow. In the r process it is the proton-deficient radioactive progenitors of the stable nuclei which are involved. Low capture cross sections and small beta-decay rates at shell closure lead to large abundances, but after subsequent radioactive decay these large abundances appear at lower A values than for the s process since Z is less and thus $A=N+Z$ is less. In Hoyle's classic papers (1946,1954) stellar nucleosynthesis up to the iron-group elements was attained by charged-particle reactions. Rapidly rising Coulomb barriers for charged particles curtailed further synthesis. Suess and Urey (1956) made the breakthrough which led to the extension of nucleosynthesis in stars by neutrons unhindered by Coulomb barriers all the way to ^{238}U.

The complete run of the synthesis of the elements in stars was incorporated into a paper by Burbidge, Burbidge, Fowler, and Hoyle (1957), commonly referred to as B^{2}FH (see also Hoyle, Fowler, Burbidge, and Burbidge, 1956), and was independently developed by Cameron (1957). Notable contributions to the astronomical aspects of the problem were made by Jesse Greenstein (1954,1982) and by many other observational astronomers. The discovery of technicium lines in S-stars by Merrill (1952) was of key importance. Since that time Nuclear Astrophysics has developed into a full-fledged scientific activity, including the exciting discoveries of isotopic anomalies in meteorites by my colleagues Gerald Wasserburg, Dimitri Papanastassiou, and Samuel Epstein and many other cosmochemists. What follows will highlight a few of the many experiments and theoretical researches under way at the present time or carried out in the past few years. This account will emphasize research activities in the Kellogg Laboratory because they are closest to my interest and knowledge. However, copious references to the work of other laboratories and institutions are cited in the hope that the reader will obtain a broad view of current experimental and theoretical studies in Nuclear Astrophysics.

This account cannot discuss the details of the nucleosynthesis of all the elements and their isotopes, which would, for a given nuclear species, involve discussing all the reactions producing that nucleus and all those which destroy it. The reader will find some of these details for ^{12}C, ^{16}O, and ^{55}Mn.

It will be noted that the measured cross sections for the reactions are customarily very small at the lowest energies of measurement, for $^{12}\mathrm{C}(\alpha,\gamma)^{16}\mathrm{O}$ even less than one nanobarn (10^{-33} cm^2) near 1.4 MeV. This means that experimental Nuclear Astrophysics requires accelerators with large currents of well focused, monoenergetic ion beams, thin targets of high purity and stability, detectors of high sensitivity and energy resolution, and experimentalists with great tolerance for the long running times required and with patience in accumulating data of statistical significance. Classical Rutherfordian measurements of nuclear cross sections are required in experimental Nuclear Astrophysics, and the results are in turn essential to our understanding of the physics of nuclei.

A comment on nuclear reaction notation is necessary at this point. In the reaction $^{12}\mathrm{C}(\alpha,\gamma)^{16}\mathrm{O}$ discussed in the previous paragraph ^{12}C is the laboratory target nucleus, α is the incident nucleus (^{4}He) accelerated in the laboratory, γ is the photon produced and detected in the laboratory, and ^{16}O is the residual nucleus which can also be detected if it is desirable to do so. If ^{12}C is accelerated against a gas target of ^{4}He and the ^{16}O products are detected but not the gamma rays, then the laboratory notation is $^4\mathrm{He}(^{12}\mathrm{O},^{16}\mathrm{O})\gamma$. The stars could not care less. In stars all the particles are moving, and only the center-of-momentum system is important for the determination of stellar reaction rates. In $^{12}\mathrm{C}(\alpha,n)^{15}\mathrm{O}(e^+\nu)^{15}\mathrm{N}$, n is the neutron promptly produced and detected and e^+ is the positron which can also be detected in the beta decay of ^{15}O. The neutrino emitted with the positron is designated by ν.

As an aside at this point I am proud to recall that I first spoke to the Royal Swedish Academy of Sciences on "Nuclear Reactions in Stars" on January 26, 1955. It does not seem so long ago, and some of you in the audience heard that talk!

III. STELLAR REACTION RATES FROM LABORATORY CROSS SECTIONS

Thermonuclear reaction rates in stars are customarily expressed as $N_A\langle\sigma v\rangle$ reactions per sec per (mol cm^{-3}), where $N_A=6.022\times10^{23}$ mol^{-1} is Avogadro's number and $\langle\sigma v\rangle$ is the Maxwell-Boltzmann average as a function of temperature for the product of the reaction cross section, σ, in cm^2, and the relative velocity of the reactants, v in cm sec^{-1}. Multiplication of $\langle\sigma v\rangle$ by the product of the number densities per cm^3 of the two reactants is necessary to obtain rates in reactions per sec per cm^3. N_A is incorporated so that mass fractions can be used, as described in detail in Fowler, Caughlan, and Zimmerman (1967,1975; see also Harris *et al.*, 1983 and Caughlan, 1984). These authors also describe procedures for reactions involving more than two reactants and give analytical expressions for reactions mainly involving γ, e, n, p, and α with nuclei having atomic mass number $A \leq 30$. Bose-Einstein statistics for γ have been necessarily incorporated, but the extension to Fermi-Dirac statistics for degenerate e, n, and p and the extension to Bose-Einstein statistics for α are not included. Factors for calculating reverse reaction rates are given.

Early work on the evaluation of stellar rates from experimental laboratory cross sections was reviewed in Bethe's Nobel Lecture (1967). Fowler, Caughlan, and Zimmerman (1967) have provided detailed numerical and analytical procedures for converting laboratory cross sections into stellar reaction rates. It is first of all necessary to accommodate the rapid variation of the nuclear cross sections at low energies which are relevant in astrophysical circumstances. For neutron-induced reactions this is accomplished by defining a cross-section S factor equal to the cross section (σ) multiplied by the interaction velocity (v) in order to eliminate the usual v^{-1} singularity in the cross section at low velocities and low energies.

For reactions induced by charged particles such as protons, alpha particles, or the heavier ^{12}C,^{16}O, . . . , nuclei it is necessary to accommodate the decrease by many orders of magnitude from the lowest laboratory measurements to the energies of astrophysical relevance. This is done in the way first suggested by E. E. Salpeter (1952b,1955) and emphasized by Bethe (1967). Table I shows how a relatively slowly varying S factor can be defined by eliminating the rapidly varying term in the Gamow penetration factor governing transmission through the Coulomb barrier. The cross section is usually expressed in barns (10^{-24} cm^2) and the energy in MeV (1.602×10^{-6} erg), so the S factor is expressed in MeV b, although keV b is sometimes used. In Table I, the two charge numbers and the reduced mass in atomic mass units of the interacting nuclei are designated by Z_0, Z_1, and A. Table II then shows how stellar reaction rates can be calculated as an average over the Maxwell-Boltzmann distribution for both nonresonant and resonant cross sections. In Table II the effective stellar reaction energy is given numerically by $E_0 = 0.122(Z_0^2 Z_1^2 A)^{1/3} T_9^{2/3}$ MeV, where T_9 is the temperature in units of 10^9 K. Expressions for reaction rates derived from theoretical statistical model calculations are given by Woosley, Fowler, Holmes, and Zimmerman (1978).

It is true that the extrapolation from the cross sections measured at the lowest laboratory energies to the cross sections at the effective stellar energy can often involve a decrease by many orders of magnitude. However, the elimination of the Gamow penetration factor, which causes this decrease, is based on the solution of the Schrödinger equation for the Coulomb wave functions, in which one can have considerable confidence. The main uncertainty lies in the variation of the S factor with energy, which depends primarily on the value chosen for the radius at which formation of a compound nucleus between two interacting nuclei or nucleons occurs, as discussed long ago in B^{2}FH (1957). The radii used by my colleagues and me in recent work are given in Woosley, Fowler, Holmes, and Zimmerman (1978). There is, in ad-

TABLE I.

DEFINITION OF THE S FACTOR (BETHE, 1967)
AS A FUNCTION OF REACTION ENERGY (E)

$$\sigma(E) = \pi\lambda\!\!\!^{-}{}^2 \times P \times \text{INTRINSIC NUCLEAR FACTOR}$$

$$\pi\lambda\!\!\!^{-}{}^2 \propto E^{-1} \qquad \lambda\!\!\!^{-} = \text{DE BROGLIE WAVELENGTH}/2\pi$$

$$P(E) = \text{GAMOW PENETRATION FACTOR}$$

$$\propto \exp(-E_G^{1/2}/E^{1/2}) \qquad E_G \approx Z_0^2 Z_1^2 A \quad \text{MeV}$$

$$S(E) \equiv E\sigma(E)\ \exp(+E_G^{1/2}/E^{1/2})$$

$$S(E) \begin{cases} \text{PERMITS MORE PRECISE EXTRAPOLATION FROM} \\ \text{LOWEST ENERGY MEASUREMENTS IN LABORATORY} \\ \text{TO VERY LOW EFFECTIVE STELLAR ENERGIES} \end{cases}$$

TABLE II.

STELLAR REACTION RATES AS FUNCTIONS OF TEMPERATURE(T)
$\langle\sigma v\rangle_{MB} = f(T) \propto T^{-3/2}\int S(E)\exp(-E_G^{1/2}/E^{1/2} - E/kT)dE$
MB $\equiv$ AVERAGE OVER MAXWELL-BOLTZMANN DISTRIBUTION
MAXIMUM IN INTEGRAND OCCURS AT E_r AND AT
$E_o \equiv$ EFFECTIVE STELLAR REACTION ENERGY $\propto E_G^{1/3}T^{2/3}$
NONRESONANT RATE
$\langle\sigma v\rangle_{nr} \propto S(E_o)T^{-2/3}\exp(-3E_o/kT) \qquad E_o/kT \propto T^{-1/3}$
RESONANT RATE
$\langle\sigma v\rangle_{r} \propto S(E_r)T^{-3/2}\exp(-E_r/kT)$
E_r = ENERGY AT RESONANCE

dition, the uncertainty in the *intrinsic nuclear factor* of Table I, which can only be eliminated by recourse to laboratory experiments. The effect of a resonance in the compound nucleus just below or just above the threshold for a given reaction can often be ascertained by determination of the properties of the resonance in other reactions in which it is involved and which are easier to study.

IV. HYDROGEN BURNING IN MAIN SEQUENCE STARS AND THE SOLAR NEUTRINO PROBLEM

Hydrogen burning in main sequence stars has contributed at the present time only about 20% more helium than that which resulted from the big bang. However, hydrogen burning in the sun has posed a problem for many years. In 1938 Bethe and Critchfield (1938) proposed the proton-proton or *pp* chain as one mechanism for hydrogen burning in stars. From many cross-section measurements in Kellogg and elsewhere, it is now known to be the mechanism which operates in the sun rather than the CNO cycle.

Our knowledge of the weak nuclear interaction (beta decay, neutrino emission and absorption, etc.) tells us that two neutrinos are emitted when four hydrogen nuclei are converted into helium nuclei. Detailed elaboration of the *pp* chain by Fowler (1958) and Cameron (1958) showed that a small fraction of these neutrinos, those from the decay of ^{7}Be and ^{8}B, should be energetic enough to be detectable through interaction with the nucleus ^{37}Cl to form radioactive ^{37}Ar, a method of neutrino detection suggested by Pontecorvo (1946) and Alvarez (1948). Raymond Davis (1983) and his collaborators have attempted for more than 25 years to detect these energetic neutrinos, employing a 380 000-liter tank of perchloroethylene ($C_2{}^{35}Cl_3{}^{37}Cl_1$) located one mile deep in the Homestake Gold Mine in Lead, South Dakota. They find only about one-quarter of the number expected on the basis of the model-dependent calculations of Bahcall *et al.* (1982)

Something is wrong—either the standard solar models are incorrect, the relevant nuclear cross sections are in error, or the electron-type neutrinos produced in the sun are converted in part into undetectable muon neutrinos or taon neutrinos on the way from the sun to the Earth. There indeed have been controversies about the nuclear cross sections, which have been for the most part resolved, as reviewed in Robertson *et al.* (1983) and Osborne *et al.* (1982) and Skelton and Kavanagh (1984).

It is generally agreed that the next step is to build a detector which will detect the much larger flux of low-energy neutrinos from the sun through neutrino absorption by the nucleus ^{71}Ga to form radioactive ^{71}Ge. This will require 30—50 tons of gallium, at a cost (for 50 tons) of approximately 25 million dollars or 200 million krona. An international effort is being made to obtain the necessary amount of gallium. We are back at square one in Nuclear Astrophysics. Until the solar neutrino problem is resolved, the basic principles underlying the operation of nuclear processes in stars are in question. A gallium detector should go a long way toward resolving the problem.

The chlorine detector must be maintained in low-level operation until the chlorine and gallium detectors can be operated at full level simultaneously. Otherwise endless conjecture concerning time variations in the solar neutrino flux will ensue. Moreover, the results of the gallium observations may uncover information that has been overlooked in the past chlorine observations.

The CNO cycle operates at the higher temperatures which occur during hydrogen burning in main sequence stars somewhat more massive than the sun. This is the case because the CNO cycle reaction rates rise more rapidly with temperature than do those of the *pp* chain. The cycle is important because ^{13}C, ^{14}N, ^{15}N, ^{17}O, and ^{18}O are produced from ^{12}C and ^{16}O as seeds. The role of these nuclei as sources of neutrons during helium burning is discussed in Sec. V.

V. THE SYNTHESIS OF ^{12}C AND ^{16}O AND NEUTRON PRODUCTION IN HELIUM BURNING

The human body is 65% oxygen by mass and 18% carbon, with the remainder mostly hydrogen. Oxygen (0.85%) and carbon (0.39%) are the most abundant elements heavier than helium in the sun and similar main sequence stars. It is little wonder that the determination of the ratio ^{12}C/^{16}O produced in helium burning is a problem of paramount importance in Nuclear Astrophysics. This ratio depends in a fairly complicated manner on the density, temperature, and duration of helium burning, but it depends directly on the relative rates of the $3\alpha \rightarrow {}^{12}$C process and the ^{12}C$(\alpha,\gamma)^{16}$O process. If $3\alpha \rightarrow {}^{12}$C is much faster than ^{12}C$(\alpha,\gamma)^{16}$O, then no ^{16}O is produced in helium burning. If the reverse is true, then no ^{12}C is produced. For the most part the subsequent reaction ^{16}O$(\alpha,\gamma)^{20}$Ne is slow enough to be neglected.

There is general agreement about the rate of the $3\alpha \rightarrow {}^{12}$C process, as reviewed by Barnes (1982). However there is a lively controversy at the present time about the laboratory cross section for ^{12}C$(\alpha,\gamma)^{16}$O and about its theoretical extrapolation to the low energies at which the reaction effectively operates. The situation is depicted in Figs. 4, 5, and 6, taken with some modification from Langanke and Koonin (1983), Dyer and Barnes (1974), and Kettner *et al.* (1982). The Caltech data obtained in the Kellogg Laboratory is shown as the experimental points in Fig. 4, taken from Dyer and Barnes (1974), who compared their results with theoretical calculations by Koonin, Tombrello, and Fox (1974). The Münster data are shown as the experimental points in Fig. 5, taken from

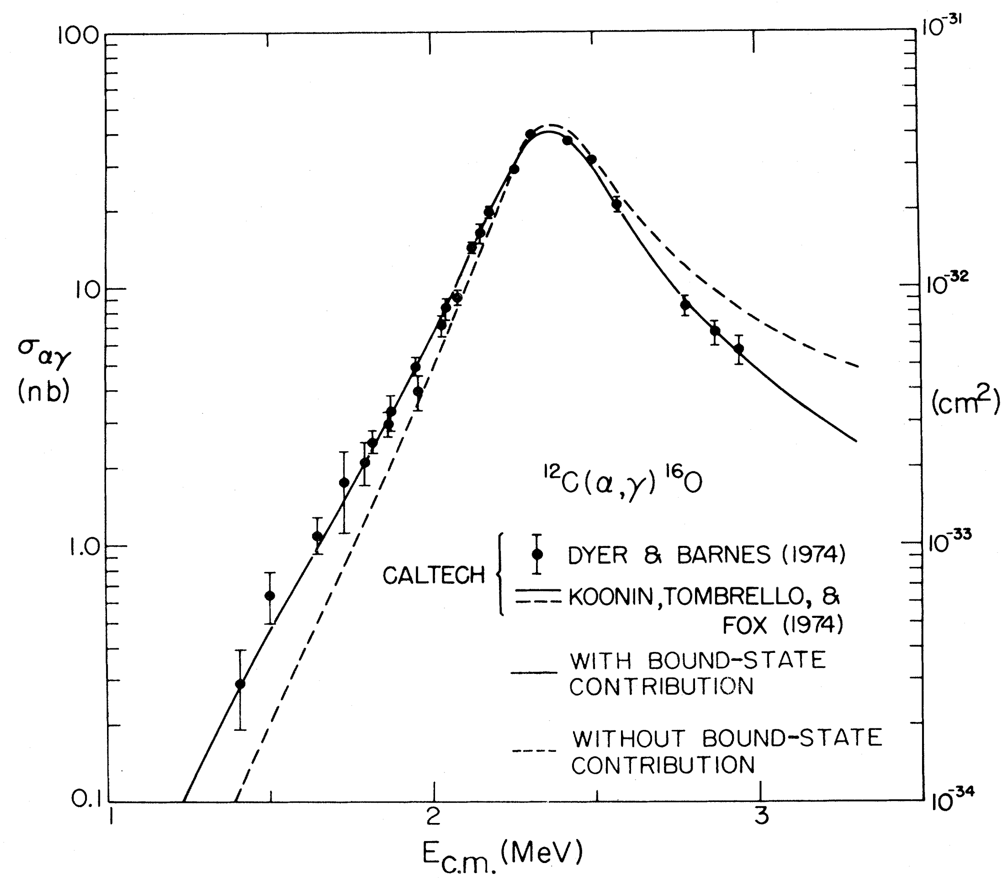

FIG. 4. The cross section in nanobarns (nb) vs center-of-momentum energy in MeV for ^{12}C$(\alpha,\gamma)^{16}$O, measured by Dyer and Barnes (1974) and compared with theoretical calculations by Koonin, Tombrello, and Fox (1974).

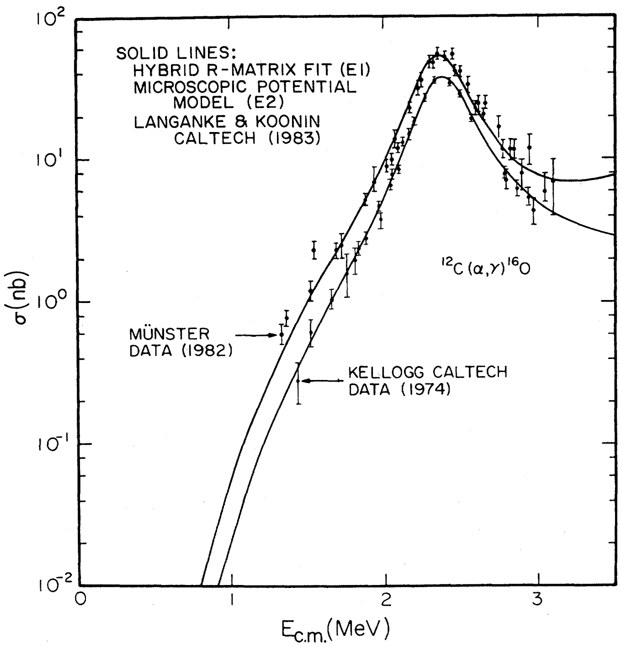

FIG. 5. The cross section in nanobarns (nb) vs center-of-momentum energy in MeV for $^{12}C(\alpha,\gamma)^{16}O$. The Münster data were obtained by Kettner *et al.* (1982), and the Kellogg Caltech data were obtained by Dyer and Barnes (1974). The solid lines are theoretical calculations made by Langanke and Koonin (1983).

Kettner *et al.* (1982) in comparison with the data of Dyer and Barnes (1974). The theoretical curves which yield the best fit to the two sets of data are from Langanke and Koonin (1983, and private communication).

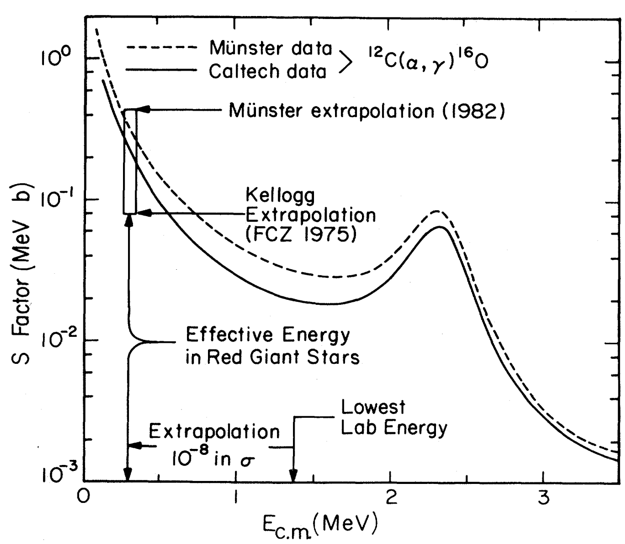

FIG. 6. The cross-section factor S in MeV b, vs center-of-momentum energy in MeV for $^{12}C(\alpha,\gamma)^{16}O$. The dashed and solid curves are the theoretical extrapolations of the Münster and Kellogg Caltech data, respectively, by Langanke and Koonin (1983).

The crux of the situation is made evident in Fig. 6, which shows the extrapolations of the Caltech and Münster cross-section factors from the lowest measured laboratory energies ($\sim$1.4 MeV) to the effective energy $\sim$0.3 MeV, at $T=1.8\times10^8$ K, a representative temperature for helium burning in red giant stars. The extrapolation in cross sections covers a range of 10^{-8}! The rise in the cross-section factor is due to the contributions of two bound states in the ^{16}O nucleus just below the $^{12}C(\alpha,\gamma)^{16}O$ threshold, as clearly indicated in Fig. 4. It is these contributions plus differences in the laboratory data which produce the current uncertainty in the extrapolated S factor. Note that Langanke and Koonin (1983) increase the 1975 extrapolation of the Caltech data by Fowler, Caughlan, and Zimmerman (1975) by a factor of 2.7 and lower the 1982 extrapolation of the Münster data by 23%. There remains a factor of 1.6 between their extrapolation of the Münster data and of the Caltech data. There is a lesson in all of this. The semiempirical extrapolation of their data by the experimentalists, Dyer and Barnes (1974), was only 30% lower than that of Langanke and Koonin (1983) and their quoted uncertainty extended to the value of Langanke and Koonin (1983). Caughlan *et al.* (1984) will tabulate the analysis of the Caltech data by Langanke and Koonin (1983).

With so much riding on the outcome, it will come as no surprise that both laboratories are engaged in extending their measurements to lower energies with higher precision. In the discussion of quasistatic silicon burning in what follows, it will be found that the abundances produced in that stage of nucleosynthesis depend in part on the ratio of ^{12}C to ^{16}O produced in helium burning, and that the different extrapolations shown in Fig. 6 are in the range crucial to the ultimate outcome of silicon burning. These remarks do not apply to explosive nucleosynthesis.

Recently the ratio of ^{12}C to ^{16}O produced under the special conditions of helium flashes during the asymptotic giant phase of stellar evolution has become of great interest. The hot blue star PG 1159-035 has been found to undergo nonradial pulsations with periods of 460 and 540 sec and others not yet accurately determined. The star is obviously highly evolved, having lost its hydrogen atmosphere, leaving only a hot dwarf of about 0.6 solar masses behind. Theoretical analysis of the pulsations by Starrfield *et al.* (1983) and Becker (1983, private communication) requires substantial amounts of oxygen in the pulsation-driving regions, where the oxygen is alternately ionized and deionized. Carbon is completely ionized in these regions and only diminishes the pulsation amplitude. It is not yet clear that sufficient oxygen is produced in helium flashes, which certainly involve $3\alpha\rightarrow{}^{12}C$ but may not last long enough for $^{12}C(\alpha,\gamma)^{16}O$ to be involved. The problem may not lie in the nuclear reaction rates, according to Starrfield *et al.* (1983) and Becker (1983). We shall see!

In what follows in this paper β^+ decay is designated by $(e^+\nu)$, since both a positron (e^+) and a neutrino (ν) are *emitted.* Similarly β^- decay will be designated by $(e^-\bar{\nu})$, since both an electron (e^-) and antineutrino $(\bar{\nu})$ are *emit-*

ted. Electron capture (often indicated by ε) will be designated by (e^-,ν), the comma indicating that an electron is captured and a neutrino emitted. The notations $(e^+,\bar{\nu})$, (ν,e^-), and $(\bar{\nu},e^+)$ should now be obvious.

Neutrons are produced when helium burning occurs under circumstances in which the CNO cycle has been operative in the previous hydrogen burning. When the cycle does not go to completion, copious quantities of ^{13}C are produced in the sequence of reactions $^{12}\mathrm{C}(p,\gamma)\,^{13}\mathrm{Ne}(e^+v)\,^{13}\mathrm{C}$. In subsequent helium burning, neutrons are produced by $^{13}\mathrm{C}(\alpha,n)\,^{16}\mathrm{O}$. When the cycle goes to completion the main product ($>95\%$) is ^{14}N. In subsequent helium burning, ^{18}O and ^{22}Ne are produced in the sequence of reactions $^{14}\mathrm{N}(\alpha,\gamma)\,^{18}\mathrm{F}(e^+v)\,^{18}\mathrm{O}(\alpha,\gamma)\,^{22}\mathrm{Ne}$, and these nuclei in turn produce neutrons through $^{18}\mathrm{O}(\alpha,n)\,^{21}\mathrm{Ne}(\alpha,n)\,^{24}\mathrm{Mg}$ and $^{22}\mathrm{Ne}(\alpha,n)\,^{25}\mathrm{Mg}$. However, the astrophysical circumstances and sites under which the neutrons produce heavy elements through the *s* process and the *r* process are, even today, matters of some controversy and much study (see Sec. XI).

VI. CARBON, NEON, OXYGEN, AND SILICON BURNING

The advanced burning processes discussed in this section involve the network of reactions shown in Fig. 7. Because of the high temperature at which this network can operate, radioactive nuclei can live long enough to serve as live reaction targets. In addition, excited states of even the stable nuclei are populated and also serve as targets. The determination of the nuclear cross sections and stellar rates of the approximately 1000 reactions in the network has involved and will continue to involve extensive experimental and theoretical effort.

The following discussion applies to massive enough stars such that electron degeneracy does not set in as nuclear evolution proceeds through the various burning stages discussed in this section. In less massive stars electron degeneracy can terminate further nuclear evolution at certain stages, with catastrophic results leading to the disruption of the stellar system. The reader will find Fig. 8, especially 8(a), instructive in following the discussion in this section. Figure 8 is taken from Woosley and Weaver (1982); a much more detailed recent version is shown in Fig. 9 from Weaver, Woosley, and Fuller (1983). Figure 8(a) applies to the preexplosive stage of a young (Population I) star of 25 solar masses and shows the result of various nuclear burnings in the following mass zones: (1) $>10M_\odot$, convective envelope with the results of some CNO burning; (2) $7-10M_\odot$, products mainly of H burning; (3) $6.5-7M_\odot$, products of He burning; (4) $1.9-6.5M_\odot$, products of C burning; (5) $1.8-1.9M_\odot$, products of Ne burning; (6) $1.5-1.8M_\odot$, products of O burning; (7) $<1.5M_\odot$, the products of Si burning in the partially neutronized core are not shown in detail, but consist mainly of ^{54}Fe as well as substantial amounts of other neutron-rich nuclei such as ^{48}Ca, ^{50}Ti, ^{54}Cr, and ^{58}Fe. ^{54}Fe, ^{48}Ca, and ^{50}Ti have $N=28$, for which a neutron subshell is closed. Both Figs. 8(a) and 8(b) have been

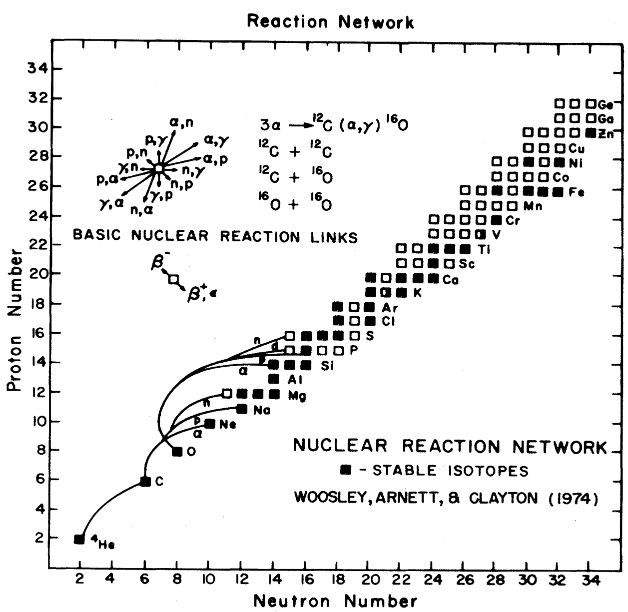

FIG. 7. The reaction network for nucleosynthesis involving the most important stable and radioactive nuclei with $N=2-34$ and $Z=2-32$. Stable nuclei are indicated by solid squares. Radioactive nuclei are indicated by open squares.

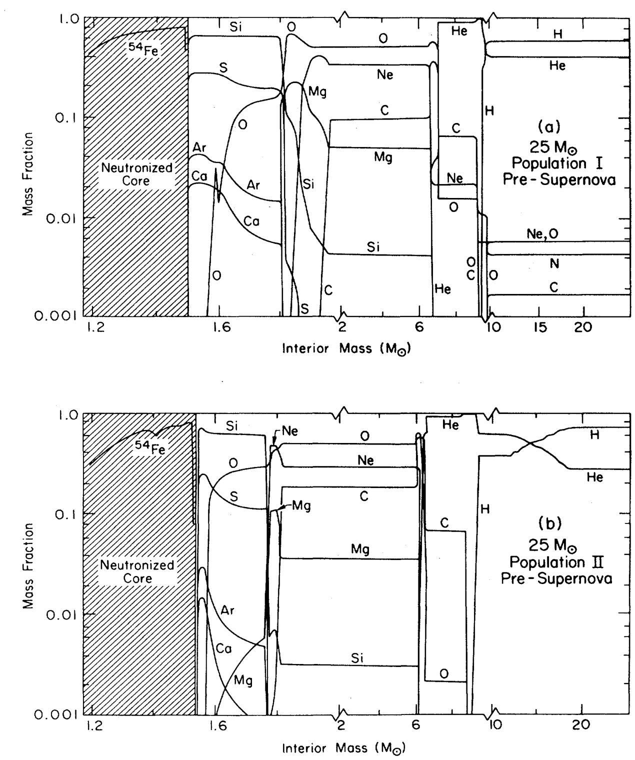

FIG. 8. Presupernova abundances by mass fraction vs increasing interior mass in solar masses $M_\odot$, measured from zero at the stellar center to $25M_\odot$, the total stellar mass, from Woosley and Weaver (1982): (a) Population I star; (b) Population II star.

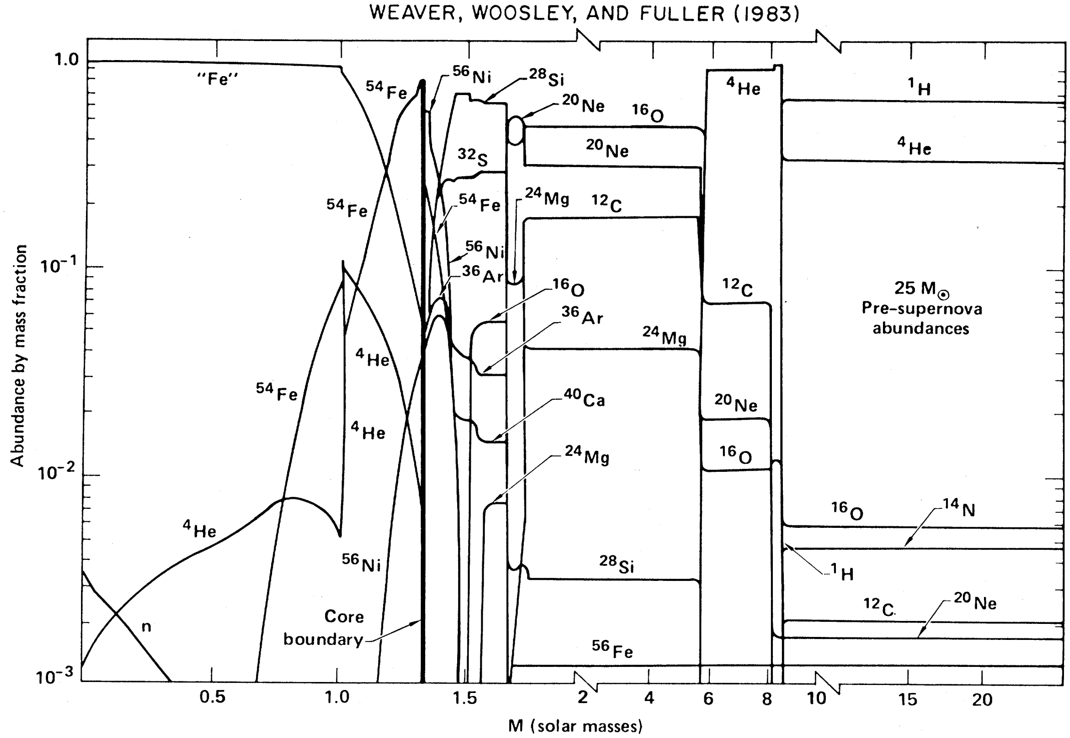

FIG. 9. Presupernova abundances by mass fraction vs increasing interior mass for a Population I star with total mass equal to $25M_\odot$, from Weaver, Woosley, and Fuller (1983).

evaluated shortly after photodisintegration has initiated core collapse, which will then be subsequently sustained by the reduction of the outward pressure through electron capture and the resulting almost complete neutronization of the core.

It must be realized that the various burning stages took place originally over the central regions of the star and finally in a shell surrounding that region. Subsequent stages modify the inner part of the previous burning stage. For example, in the 25 solar mass Population I star of Fig. 8(a), C burning took place in the central 6.5 solar masses of the star, but the inner 1.9 solar masses were modified by subsequent Ne, O, and Si burning.

Helium burning produces a stellar core consisting mainly of ^{12}C and ^{16}O. After core contraction the temperature and density rise until carbon burning through $^{12}C+^{12}C$ fusion is ignited. The S factor for the total reaction rate shown in Fig. 10 has been taken from p. 213 of Barnes (1982) and is based on measurements in a number of laboratories. The extrapolation to the low energies of astrophysical relevance is uncertain, as Fig. 10 makes clear, and more experimental and theoretical studies are urgently needed. At the lowest bombarding energy, 2.4 MeV, the cross section is $\sim 10^{-8}$ b. For a representative burning temperature of 6×10^8 K the effective energy is $E_0=1.7$ MeV and the extrapolated cross section is $\sim 10^{-13}$ b. The main product of carbon burning is ^{20}Ne, produced primarily in the $^{12}C(^{12}C,\alpha)^{20}Ne$ reaction. The reactions $^{12}C(^{12}C,p)^{23}Na$ and $^{12}C(^{12}C,n)^{23}Mg(e^+v)^{23}Na$ also occur, as well as many secondary reactions such as $^{23}Na(p,\alpha)^{20}Ne$. When the ^{12}C is exhausted, ^{20}Ne and ^{16}O are the major remaining constituents. As the temperature rises, from further gravitational contraction, the ^{20}Ne is destroyed by photodisintegration, $^{20}Ne(\gamma,\alpha)^{16}O$. This occurs because the alpha particle in ^{20}Ne is bound to its closed-shell partner, ^{16}O, by only 4.731 MeV. In ^{16}O, for example, the binding of an alpha particle is 7.162 MeV.

The next stage is oxygen burning through $^{16}O+^{16}O$ fusion. The S factor for the total reaction rate is shown in Fig. 11 and is based entirely on data obtained in the Kellogg Laboratory at Caltech. The work of Hulke, Rolfs, and Trautvetter (1980) using gamma-ray detection is in fair agreement with the gamma-ray measurements at Caltech. As in the case of $^{12}C+^{12}C$, the extrapolation to the low energies of astrophysical relevance is uncertain, although only one of many possible extrapolations is shown in Fig. 11. The main product of oxygen burning is ^{28}Si, through the primary reaction $^{16}O(^{16}O,\alpha)^{28}Si$ and a number of secondary reactions. Under some conditions neutron-induced reactions lead to the synthesis of significant quantities of ^{30}Si. Oxygen burning can result in nuclei with a small but important excess of neutrons over protons.

The onset of Si burning signals a marked change in the nature of the *fusion process.* The Coulomb barrier between two ^{28}Si nuclei is too great for fusion to produce the compound nucleus ^{56}Ni directly at the ambient temperatures ($T_9=3-5$) and densities ($\rho=10^5-10^9$ g cm^{-3}). However, the ^{28}Si and subsequent products are easily pho-

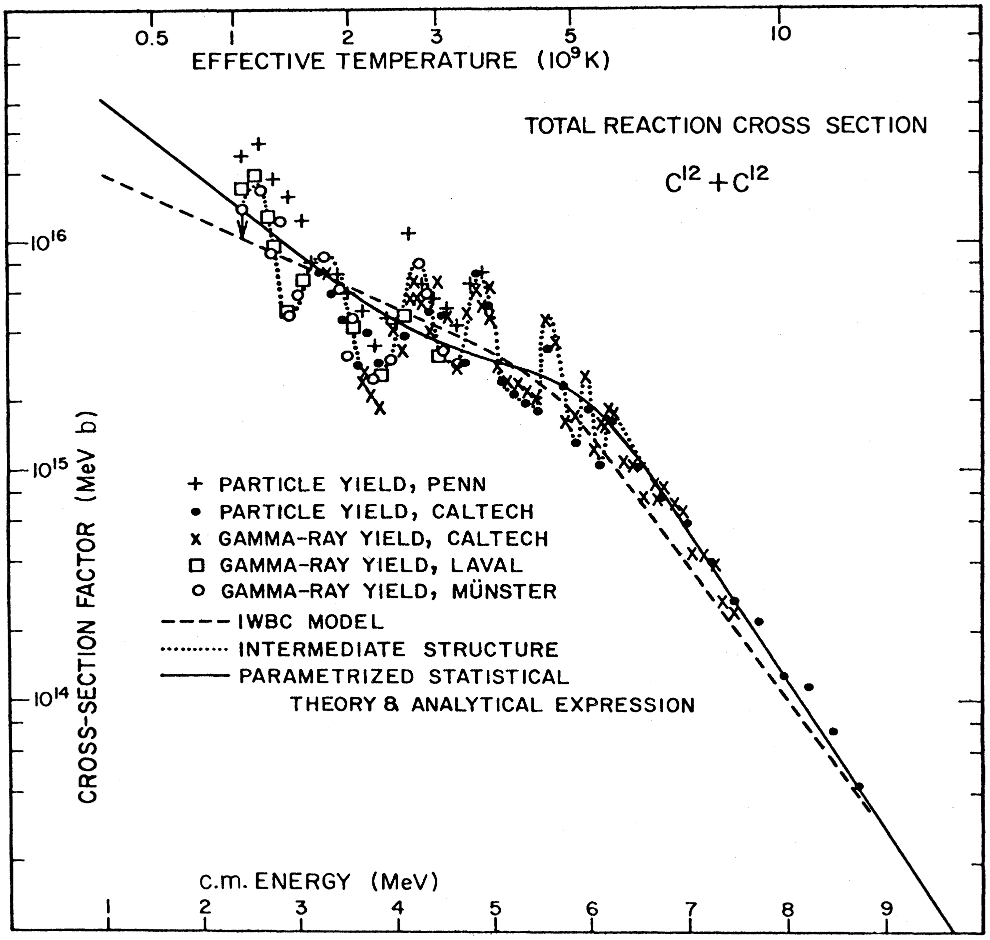

FIG. 10. The total cross-section factor in MeV b vs center-of-momentum energy in MeV for the fusion of ^{12}C and ^{12}C. The experimental data from several laboratories are shown, along with schematic intermediate structure in the dotted curve. Two parametrized adjustments to the data, ignoring intermediate structure, are shown in the dashed and solid curves.

todisintegrated by (γ,α), (γ,n), and (γ,p) reactions. As Si burning proceeds, more and more ^{28}Si is reduced to nucleons and alpha particles which can be captured by the remaining ^{28}Si nuclei to build through the network in Fig. 7 up to the iron-group nuclei. The main product in explosive Si burning is ^{56}Ni, which transforms eventually through two beta decays to ^{56}Fe.

In quasistatic Si burning the weak interactions are fast enough that ^{54}Fe, with two more neutrons than protons, is the main product. Because of the important role played by alpha particles (α) and because of the inexorable trend to equilibrium (e) involving nuclei near mass 56, which have the largest binding energies per nucleon of all nuclear species, B^2FH (1957) broke down what is now called Si burning, into their α process and e process. Quasiequilibrium calculations for Si burning were made by Bodansky, Clayton, and Fowler (1968) who cite the original papers in which the basic ideas of Si burning were developed. Modern computers permit detailed network flow calculations to be made, as discussed in Woosley and Weaver (1982) and Weaver *et al.* (1983).

The extensive laboratory studies of Si-burning reactions are reviewed in Barnes (1982). Figures 12 and 13, adapted from Zyskind *et al.* (1978), show the laboratory excitation curves for $^{54}Cr(p,n)^{54}Mn$ and $^{54}Cr(p,\gamma)^{55}Mn$ as examples. The neutrons produced in the first of these reactions will increase the number of neutrons available in Si burning but will not contribute directly to the synthesis of ^{55}Mn as does the second reaction. In fact, above its threshold at 2.158 MeV the (p,n) reaction competes strongly with the (p,γ) reaction, which is of primary interest, and produces the pronounced *competition cusp* in the excitation curve in Fig. 13. Competition in the disintegration of the compound nucleus produced in nuclear reactions was stressed very early by Niels Bohr, so perhaps the cusps should be called *Bohr cusps.* They arise from the same basic cause but are not the long known *Wigner cusps.* It will be clear from Fig. 13 that the rate of the $^{54}Cr(p,\gamma)^{55}Mn$ reaction at very high temperatures will be an order of magnitude lower because of the cusp than would otherwise be the case.

The element manganese has only one isotope, ^{55}Mn.

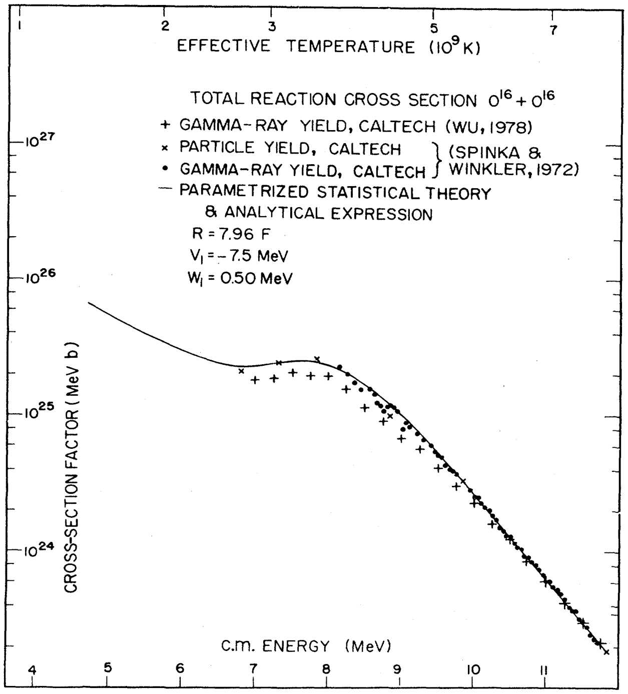

FIG. 11. The total cross-section factor in MeV b vs center-of-momentum energy in MeV for the fusion of $^{16}O + {}^{16}O$. The experimental data from several measurements at Caltech are shown and compared with a parametrized theoretical adjustment in the solid curve.

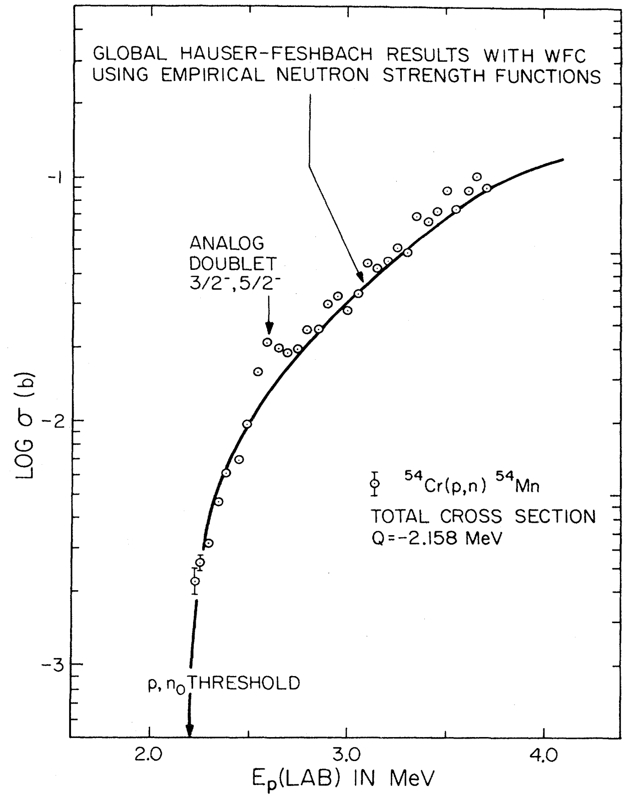

FIG. 12. The total cross section in b integrated over all outgoing angles vs laboratory proton energy in MeV for the reaction $^{54}Cr(p,n)\,^{54}Mn$. The data of Zyskind *et al.* (1978) are compared with unnormalized global Hauser-Feshbach calculations made by Woosley *et al.* (1978).

The manganese in nature is produced in quasistatic Si burning, most probably through the $^{54}Cr(p,\gamma)\,^{55}Mn$ reaction just discussed in the previous paragraph. The reaction network extends to ^{54}Cr and then on through ^{55}Mn. $^{51}V(\alpha,\gamma)\,^{55}Mn$ and $^{52}V(\alpha,n)\,^{55}Mn$ may also contribute, especially in explosive Si burning. The overall synthesis of ^{55}Mn involves a balance in its production and destruction. In quasistatic Si burning the reactions which destroy ^{56}Mn are most probably $^{55}Mn(p,\gamma)\,^{56}Fe$ and $^{55}Mn(p,n)\,^{55}Fe$, which are discussed and illustrated in Mitchell and Sargood (1983). $^{55}Mn(\alpha,\gamma)\,^{59}Co$, $^{55}Mn(\alpha,p)\,^{58}Fe$, and $^{55}Mn(\alpha,n)\,^{58}Co$ may also destroy some ^{55}Mn in explosive Si burning. In the figures discussed in Sec. VIII it will be noted that calculations of the overall synthesis of ^{55}Mn yield values in fairly close agreement with the abundance of this nucleus in the solar system. Unfortunately, the same cannot be said about many other nuclei.

The laboratory measurements on Si-burning reactions have covered only about 20% of the reactions in the network of Fig. 7 involving stable nuclei as targets. Direct measurements on short-lived radioactive nuclei and the excited states of all nuclei are impossible at the present time. In this connection the production of radioactive ion beams holds great promise for the future. Richard Boyd (1981) and Haight *et al.* (1983) have pioneered in the development of this technique. Figures 14 and 15 show the beam transport system developed by Haight *et al.* (1983), which has produced accelerated beams of ^7Be and ^{13}N and successfully determined the cross section of the reaction $^2H(^7Be,^8B)n$ to be 59 ± 11 mb for 16.9-MeV ^7Be ions. The equivalent center-of-momentum energy for the $^7Be(d,n)\,^8B$ reaction is 3.8 MeV. It is my view that continued development and application of radioactive ion-beam techniques could bring the most exciting results in laboratory Nuclear Astrophysics in the next decade. For example, the rate of the $^{13}N(p,\gamma)\,^{14}O$ reaction, which will be studied as $^1H(^{13}N,\gamma)\,^{14}O$, is crucial to the operation of the so-called fast CN cycle.

In any case it has been clear for some time that experimental results on Si-burning reactions must be systematized and supplemented by comprehensive theory. Fortunately theoretical average cross sections will suffice in many cases. This is because the stellar reaction rates integrate the cross sections over the Maxwell-Boltzmann distribution. For most Si-burning reactions resonances in the cross section are closely spaced and even overlapping, and the integration covers a wide enough range of energies that the detailed structure in the cross sections is automatically averaged out. The statistical model of nuclear reactions developed by Hauser and Feshbach (1952), which yields average cross sections, is ideal for the purpose. Accordingly Holmes, Woosley, Fowler, and Zim-

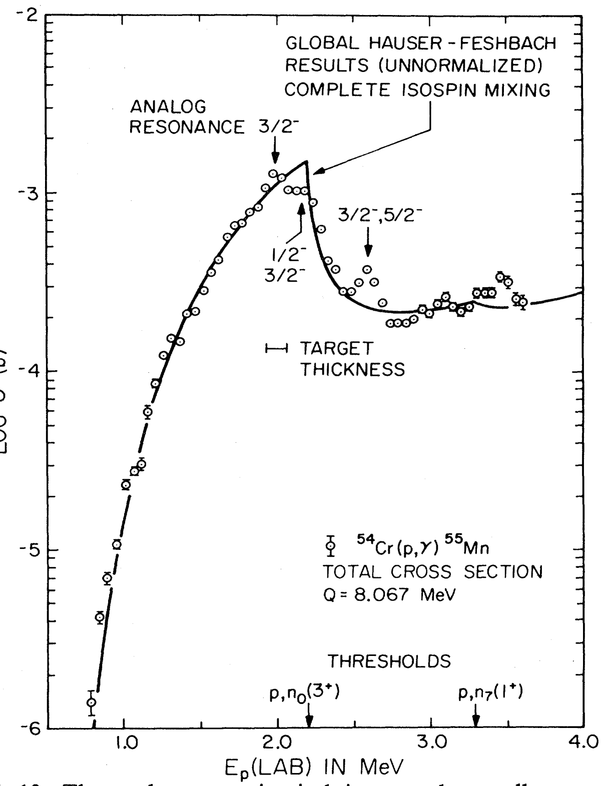

FIG. 13. The total cross section in b integrated over all outgoing angles vs laboratory proton energy in MeV for the reaction $^{54}Cr(p,\gamma)^{55}Mn$. The data of Zyskind *et al.* (1978) are compared with unnormalized global Hauser-Feshbach calculations made by Woosley *et al.* (1978).

merman (1976) undertook the task of developing a *global, parametrized* Hauser-Feshbach theory and computer program for use in Nuclear Astrophysics. The contribution of this group to the Nuclear Data Tables (Woosley *et al.*, 1978) is an extension of this work. The free parameters are the radius, depth, and compensating reflection factor of the blackbody, square-well equivalent of the Woods-Saxon potential characteristic of the interaction between n, p, and α with nuclei having $Z \geq 8$. Two free parameters must also be incorporated to adjust the intensity of electric and magnetic dipole transitions for gamma radiation. Weak interaction rates must also be specified, and

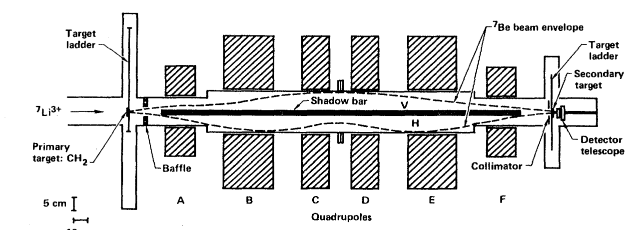

FIG. 14. Radioactive beam transport system developed by Haight *et al.* (1983).

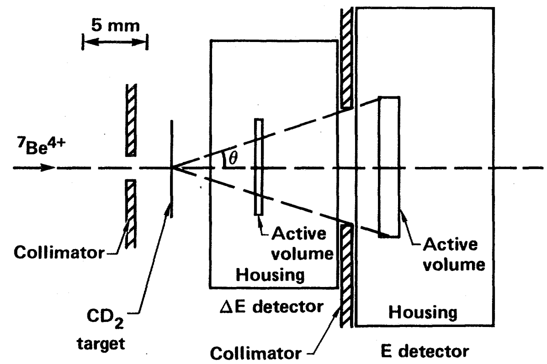

FIG. 15. Detail of the target and detector in the radioactive beam transport system developed by Haight *et al.* (1983).

ways and means for doing this will be discussed later in Sec. VII.

The parameters originally chosen for n, p, and α reactions were taken from earlier work of Michaud and Fowler (1970) who depended heavily on studies by Vogt (1969). These parameters and those chosen for electromagnetic and weak interactions have survived comparison of the theory with a plethora of laboratory measurements. More sophisticated programs have been developed which use experimental neutron strength functions instead of that from the equivalent square well or which use realistic Woods-Saxon potentials for all interactions, as done by Mann (1976). In addition marked improvement in the correspondence between theory and experiment is found when width-fluctuation corrections are made as described in Zyskind *et al.* (1980).

It is well known that the free parameters can always be adjusted to fit the cross sections and reaction rates of any one particular nuclear reaction. This is not done in a *global* program. The parameters are in principle determined by the best least-squares fit to all reactions for which experimental results are available. For example, see the figure, p. 307, in Holmes *et al.* (1976). It is on this basis that some confidence can be had in predictions in those cases where experimental results are unavailable.

The original program (Holmes *et al.*, 1976; Woosley *et al.*, 1978) has produced reaction rates either in numerical or analytical form as a function of temperature. Ready comparison with integrations of laboratory cross sections for target ground states are possible. Using the same *global* parameters which apply to reactions involving the ground states of stable nuclei, the theoretical program calculates rates for the ground states of radioactive nuclei and for the excited states of both stable and radioactive nuclei. Summing over the statistically weighted contributions of the ground and known excited states or theoretical level density functions yields the stellar reaction rate for the equilibrated statistical population of the nuclear states. After summing, division by the partition function of the target nucleus is necessary. Analytical parametrized expressions for the partition functions of nuclei with $8 \leq Z \leq 36$ are given in Table IIA of Woosley

TABLE III. Statistical model calculations vs measurements (I). Ratio of reaction rate (ground state of target) from Woosley, Fowler, Holmes, and Zimmerman (1978) to reaction rates from experimental yield measurements (1970—1982) at Bombay, Caltech, Colorado, Kentucky, Melbourne, and Toronto.

Reaction	$T_9=T/10^9$ K 1	2	3	4	5
^{23}Na$(p,n)^{23}$Mg	1.4	1.2	1.1	1.1	1.0
^{25}Mg$(p,\gamma)^{26}$Al	1.2	1.1	1.0	0.9	0.8
^{25}Mg$(p,n)^{25}$Al	1.1	1.0	0.9	0.8	0.8
^{27}Al$(p,\gamma)^{28}$Si	3.7	2.1	1.5	1.3	1.1
^{27}Al$(p,n)^{27}$Si	1.8	1.4	1.3	1.3	1.2
	0.9	0.9	0.9	1.0	1.0
^{28}Si$(p,\gamma)^{29}$P		1.2	1.3	1.2	0.9
^{29}Si$(p,\gamma)^{30}$P		1.0	1.6	1.6	1.5
^{39}K$(p,\gamma)^{40}$Ca	15	4.5	3.0	2.6	2.5
^{41}K$(p,\gamma)^{42}$Ca	0.5	0.5	0.5	0.4	0.4
^{41}K$(p,n)^{41}$Ca	0.8	1.0	1.1	1.2	1.3
^{40}Ca$(p,\gamma)^{41}$Sc				0.1	0.2
^{42}Ca$(p,\gamma)^{43}$Sc	1.3	1.4	1.4	1.4	1.3
	0.8	1.1	1.3	1.4	1.4

et al. (1978) as a function of temperature over the range $0 \le T \le 10^{10}$ K.

Sargood (1982,1983) has compared experimental results from a number of laboratories for protons and alpha particles reacting with 80 target nuclei, which are, of course, in their ground states, with the theoretical predictions of Woosley *et al.* (1978). Ratios of statistical model calculations to laboratory measurements for 12 cases are shown in Table III for temperatures in the range from 1 to 5×10^9 K. The double entry for ^{27}Al$(p,n)^{27}$Si signifies ratios of theory to measurements made in two different laboratories. It is fair to note that the theoretical calculations match the experimental results within 50% with a few marked exceptions. In American vernacular "You win some and you lose some." For the rather light targets in Table III, espcially at low temperature, the global mean rates can be in error whenever more and stronger resonances or fewer and weaker resonances than expected on average occur in the excitation curve of the reaction at low energies.

Sargood (1982,1983) has also compared the ratio of the stellar rate of a reaction with target nuclei in a thermal distribution of ground and excited states with the rate for all target nuclei in their ground state. The latter is of course determined from laboratory measurements. A number of cases are tabulated for $T=5\times10^9$ K in Table IV. In many cases, notably for reactions producing gamma rays, the ratio of stellar to laboratory rates is close to

TABLE IV. Stellar/laboratory reaction rates. $\langle\sigma v\rangle^*/\langle\sigma v\rangle^0$. Temperature $=5\times10^9$ K. Data from Sargood (1983) and Woosley, Fowler, Holmes, and Zimmerman (1978).

Target nucleus	Reaction (n,γ)	(n,p)	(n,α)	(p,γ)	(p,n)	(p,α)	(α,γ)	(α,n)	(α,p)
^{20}Ne	0.959	12.2	4.98	0.954	34.1	6.86	0.907	4.90	1.29
^{21}Ne	0.808	6.15	1.13	0.818	1.78	1.95	0.943	0.985	1.37
^{22}Ne	0.917	159	22.1	0.895	5.11	2.72	0.968	0.996	2.46
^{23}Na	0.897	4.95	9.70	0.890	2.27	0.944	0.826	1.30	0.918
^{24}Mg	0.939	20.4	7.30	0.924	120	15.0	0.835	4.70	1.04
^{25}Mg	0.905	5.05	3.18	0.862	3.48	5.02	0.958	0.973	1.10
^{26}Mg	0.968	71.4	53.8	0.958	8.05	4.92	0.974	1.00	1.41
^{27}Al	0.934	4.12	10.9	0.913	3.22	1.14	0.905	1.13	0.972
^{28}Si	0.976	6.51	7.26	0.950	140	23.5	0.933	3.55	1.02
^{29}Si	0.943	8.67	3.34	0.907	3.18	50.1	0.927	0.964	1.18
^{30}Si	0.989	89.4	28.6	0.982	2.99	6.63	0.973	1.01	1.09
^{31}P	0.972	2.63	18.4	0.901	3.77	1.11	0.969	1.70	0.978
^{32}S	0.988	2.33	1.57	0.980	90.1	7.35	0.975	3.79	1.00
^{33}S	0.943	1.46	1.06	0.920	4.73	3.24	0.916	0.995	1.01
^{34}S	1.00	25.8	13.1	0.979	8.02	2.02	0.964	1.05	1.02
^{36}S	0.998	428	95.9	1.00	1.00	1.02	0.995	1.00	1.68
^{35}Cl	0.972	1.19	3.06	0.948	4.48	1.05	0.945	1.23	0.992
^{37}Cl	0.994	26.0	13.7	0.987	1.00	1.00	0.985	1.00	0.995

unity. In other cases the ratios can be high by several orders of magnitude. This can occur for a number of reasons. It frequently occurs when the ground state can interact only through partial waves of high angular momentum, resulting in small penetration factors and thus small cross sections and rates. This makes clear a basic assumption in the prediction of stellar rates: a statistical theory which does well predicting ground-state results is assumed to do equally well in predicting excited-state results. This assumption is frequently not valid. Bahcall and Fowler (1969) have shown that in a few cases laboratory measurements on inelastic scattering involving excited states can be used indirectly to determine reaction cross sections for those states.

Ward and Fowler (1980) have investigated in detail the circumstances under which long-lived isomeric states do not come into equilibrium with ground states. When this occurs it is necessary to incorporate into network calculations the stellar rates for both the isomeric and ground state. An example of great interest is the nucleus ^{26}Al. The ground state has spin and parity $J^{\pi}=5^{+}$ and isospin $T=0$, and has a mean lifetime for positron emission to ^{26}Mg of 10^6 years. The isomeric state at 0.228 MeV has $J^{\pi}=0^{+}, T=1$, and mean lifetime 9.2 sec. Ward and Fowler (1980) show that the isomeric state effectively does not come into equilibrium with the ground state for $T<4\times10^{8}$ K. At these low temperatures both the isomeric state and the ground state of ^{26}Al must be included in the network of Fig. 7.[2]

VII. ASTROPHYSICAL WEAK-INTERACTION RATES

Weak nuclear interactions play an important role in astrophysical processes in conjunction with the strong nuclear interactions, as indicated in Fig. 7. Only through the weak interaction can the overall proton number and neutron number of nuclear matter change during stellar evolution, collapse, and explosion. The formation of a neutron star requires that protons in ordinary stellar matter capture electrons. Gravitational collapse of a Type-II supernova core is retarded as long as electrons remain to exert outward pressure.

Many years of theoretical and experimental work on weak-interaction rates in the Kellogg Laboratory and elsewhere have culminated in the calculation and tabulation by Fuller, Fowler, and Newman (1980,1982a,1982b) of electron and positron emission rates and continuum electron and positron capture rates, as well as the associated neutrino energy-loss rates for free nucleons and 226 nuclei with mass numbers between $A=21$ and 60. Extension to higher and lower values for A is now underway.

These calculations depended heavily on experimental determinations in Kellogg by Wilson, Kavanagh, and Mann (1980) of Gamow-Teller elements for 87 discrete transitions in intermediate-mass nuclei. The majority of the experimental matrix elements for both Fermi and Gamow-Teller discrete transitions as well as nuclear level data were taken from the exhaustive tabulation of Lederer and Shirley (1978). Unmeasured matrix elements for allowed transitions were assigned a mean value, as described in Fuller, Fowler, and Newman (1982a). These mean values were $|M_F|^2=0.062$ and $|M_{\rm GT}|^2=0.039$, corresponding to $\log ft=5$, where f is the phase-space factor and t is the half-life for the transition. Nuclear physicists traditionally think in terms of $\log ft$ values in connection with weak-interaction rates.

Simple shell model arguments were employed to estimate Gamow-Teller sum rules and collective-state resonance excitation energies. These estimates have been shown to be fair approximations for $T^{<}$ nuclei and $T^{>}$ nuclei by recent high-resolution measurements on p,n reactions and $^3T,^3$He reactions by Goodman *et al.* (1980) and Ajzenberg-Selove *et al.* (1984), respectively. Here $T^{<}$, with $T\equiv|N-Z|$, represents, for example, ^{56}Fe with $T=2$ in $^{56}{\rm Fe}(e^{-},\nu)^{56}{\rm Mn}$ or $^{56}{\rm Fe}(n,p)^{56}{\rm Mn}$. Similarly $T^{>}$ designates ^{56}Mn with $T=3$. The work described by Fuller, Fowler, and Newman (1980,1982a,1982b) emphasizes the great need for addition results for $T^{>}$ nuclei using the n,p reaction as well as the $^3T,^3$He reaction from which matrix elements for electron capture can be obtained.

Moment method shell model calculations of Gamow-Teller strength functions have been performed by S. D. Bloom and G. M. Fuller (1984) with the Lawrence Livermore National Laboratory's vector shell model code for the ground states and first excited states of ^{56}Fe, ^{60}Fe, and ^{64}Fe. These detailed calculations confirm the general trends in Gamow-Teller strength distributions used in the approximations of Fuller, Fowler, and Newman.

The discrete-state contribution to the rates, dominated by experimental information on the Fermi transitions, determines the weak nuclear rates in the regime of temperature and densities characteristic of the quasistatic phases of presupernova stellar evolution. At the higher temperatures and densities characteristic of the supernova collapse phase, which is of such great current interest, as discussed in detail in Brown, Bethe, and Baym (1982), the electron-capture rates are dominated by the Gamow-Teller collective resonance contribution.

The detailed nature and the difficulty of the theoretical aspects of the combined atomic, nuclear, plasma, and hydrodynamic physics problems in Type-II supernova implosion and explosion were brought home to us by Hans Bethe during his stay in our laboratory as a Caltech Fairchild Scholar early in 1982. His visit plus long-distance interaction with his collaborators resulted in the preparation of two seminal papers, Bethe, Yahil, and Brown (1982) and Bethe, Brown, Cooperstein, and Wilson (1983).

Current ideas on the nuclear equation of state predict that early in the collapse of the iron core of a massive star

[2]For recent experimental data on the production of ^{26}Al, through $^{25}{\rm Mg}(p,\gamma)^{26}{\rm Al}$, see Champagne *et al.* (1983). For recent experimental data on the destruction of ^{26}Al, through $^{26}{\rm Al}(p,\gamma)^{27}{\rm Si}$, see Buchmann *et al.* (1984).

the nuclei present will become so neutron rich that allowed electron capture on protons in the nuclei is blocked. Allowed electron capture, for which $\Delta l = 0$, is not permitted when neutrons have filled the subshells having orbital angular momentum l equal to that of the subshells occupied by the protons.

This neutron shell blocking phenomenon and several unblocking mechanisms operative at high temperature and density, including forbidden electron capture, have been studied in terms of the simple shell model by Fuller (1982). Though the unblocking mechanisms are sensitive to details of the equation of state, typical conditions result in a considerable reduction of the electron-capture rates on heavy nuclei, leading to significant dependence on electron capture by the small number of free protons and a decrease in the overall neutronization rate.

The results of one-zone collapse calculations which have been made by Fuller (1982) suggest that the effect of neutron shell blocking is to produce a larger core lepton fraction (leptons per baryon) at neutrino trapping. In keeping with the Chandrasekhar relation that core mass is proportional to the square of the lepton fraction, this leads to a larger *final* core mass and hence a stronger post-bounce shock. On the other hand, the incorporation of the new electron-capture rates during precollapse Si burning reduces the lepton fraction and leads to a smaller *initial* core mass and thus to a smaller amount of material (*initial* core mass minus *final* core mass) in which the post-bounce shock can be dissipated. The dissipation of the shock is thus reduced. This is discussed in detail in Weaver, Woosley, and Fuller (1983).

Recent work on the weak interaction has concentrated on making the previously calculated reaction rates as efficient as possible for users of the published tables and the computer tapes, which are made available on request. The stellar weak-interaction rates of nuclei are in general very sensitive functions of temperature and density. Their temperature dependence arises from thermal excitation of parent excited states and from the lepton distribution functions in the integrands of the decay and continuum capture phase-space factors.

For electron and positron emission, most of the temperature dependence is due to thermal population of parent excited states at all but the lowest temperatures and highest densities. In general, only a few transitions will contribute to these decay rates, and hence the variation of the rates with temperature is usually not so large that rates cannot be accurately interpolated in temperature and density with the standard grids provided in Fuller, Fowler, and Newman (1980,1982a,1982b). The density dependence of these decay rates is minimal. In the case of electron emission, however, there may be considerable density dependence due to Pauli blocking for electrons where the density is high and the temperature is low. This does not present much of a problem for practical interpolation, since the electron-emission rate is usually very small under these conditions.

The temperature and density dependence of continuum electron and positron capture is a much more serious problem. In addition to temperature sensitivity introduced through thermal population of parent excited states, there are considerable effects from the lepton distribution functions in the integrands of the continuum-capture phase-space factors. This sensitivity of the capture rates means that interpolation in temperature and density on the standard grid to obtain a rate can be difficult, requiring a high-order interpolation routine and a relatively large amount of computer time for an accurate value. This is especially true for electron-capture processes with threshold above zero energy.

We have found that the interpolation problem can be greatly eased by defining a simple continuum-capture phase-space integral, based on the parent-ground-state to daughter-ground-state transition Q value, and then dividing this by the tabulated rates (Fuller, Fowler, and Newman, 1980,1982a,1982b) at each temperature and density grid point to obtain a table of effective ft values; these turn out to be much less dependent on temperature and density. This procedure requires a formulation of the capture phase-space factors which is simple enough to use many times in the inner loop of stellar evolution nucleosynthesis computer programs. Such a formulation in terms of standard Fermi integrals has been found, along with approximations for the requisite Fermi integrals. When the chemical potential (Fermi energy) which appears in the Fermi integrals goes through zero, these approximations have continuous values and continuous derivatives.

We have recently found expressions for the reverse reactions to e^-,e^+ capture (i.e., $\nu,\bar{\nu}$ capture) and for $\nu,\bar{\nu}$ blocking of the direct reactions when $\nu,\bar{\nu}$ states are partially or completely filled. These reverse reactions and the blocking are important during supernova core collapse when neutrinos and antineutrinos eventually become trapped, leading to equilibrium between the two directions of capture. General analytic expressions have been derived and approximated with computer-usable equations. All of these new results described in the previous paragraphs will be published in Fuller, Fowler, and Newman (1984), and new tapes including $\nu,\bar{\nu}$ capture will be made available to users on request.

VIII. CALCULATED ABUNDANCES FOR $A \lesssim 60$ WITH BRIEF COMMENTS ON EXPLOSIVE NUCLEOSYNTHESIS

Armed with the available strong and weak nuclear reaction rates which apply to the advanced stages of stellar evolution, theoretical astrophysicists have attempted to derive the elemental and isotopic abundances produced in quasistatic presupernova nucleosynthesis and in explosive nucleosynthesis occurring during supernova outbursts.

The various stages of preexplosive nucleosynthesis have been discussed in Secs. IV, V, and VI, and it is fair to say that there is reasonably general agreement on nucleosynthesis during these stages. On the other hand, explosive nucleosynthesis is still an unsettled matter, subject to intensive study at the present time, as reviewed, for example, in Woosley, Axelrod, and Weaver (1984).

The abundances produced in explosive nucleosynthesis must of necessity depend on the detailed nature of supernova explosions. Ideas concerning the nature of Type-I and Type-II supernova explosions were published many years ago by Hoyle and Fowler (1960) and Fowler and Hoyle (1964). It was suggested that Type-I supernovae of small mass were precipitated by the onset of explosive carbon burning under conditions of electron degeneracy where pressure is approximately independent of temperature. Carbon burning raises the temperature to the point where the electrons are no longer degenerate, and explosive disruption of the star results. For Type-II supernovae of larger mass it was suggested that Si burning produced iron-group nuclei, which have the maximum binding energies of all nuclei, so that nuclear energy is no longer available. Subsequent photodisintegration and electron capture in the stellar core leads to core implosion and ignition of explosive nucleosynthesis in the infalling inner mantle, which still contains nuclear fuel. These ideas have "survived" but, to say the least, with considerable modification over the years, as indicated in the excellent review by Wheeler (1981). Modern views on Type-II supernovae are given in Weaver *et al.* (1983), Brown *et al.* (1982), Bethe *et al.* (1982), and Bethe *et al.* (1983); modern views on Type-I supernovae are given in Nomoto (1982a,1982b,1984).

We can return to the nuclear abundance problem by reference to Fig. 16, taken from Woosley and Weaver

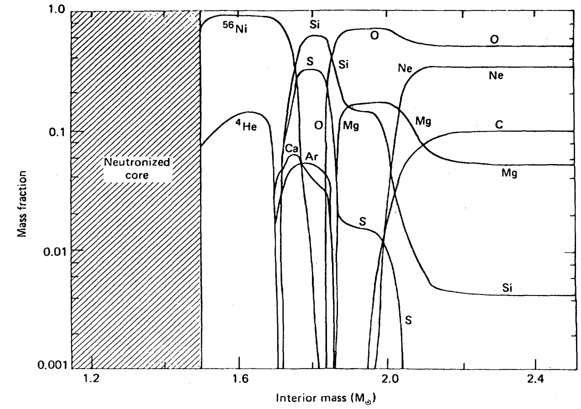

FIG. 16. Final abundances by mass fraction vs increasing interior mass in solar masses $M_\odot$, in Type-II supernova ejecta from a Population I star with total mass equal to $25M_\odot$, from Woosley and Weaver (1982).

(1982), which shows the distribution of the final abundances by mass fraction in the supernovae ejecta of a $25M_\odot$ Population I star. The presupernova distribution is that shown in Fig. 8(a). The modification in the abundances for the mass zones interior to $2.2M_\odot$ is very apparent. The mass exterior to $2.2M_\odot$ is ejected with little or no modification in nuclear abundances. The supernovae explosion was simulated by arbitrarily assuming that on the order of 10^{51} ergs was delivered to the ejected material by the shock generated in the bounce or rebound of

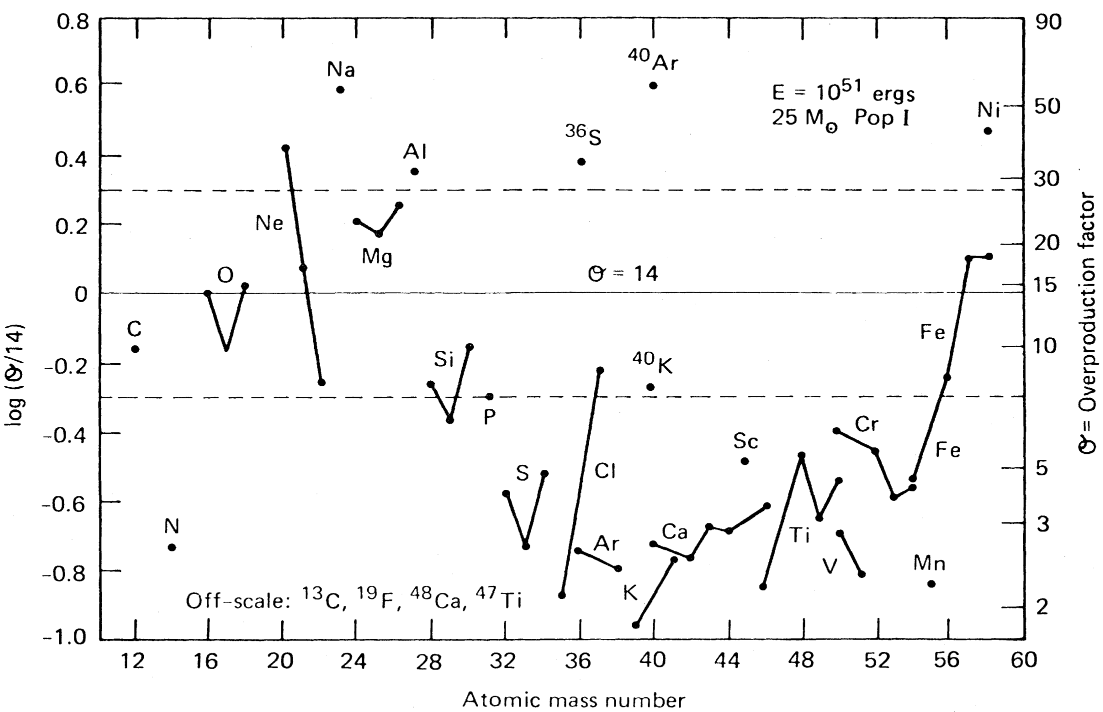

FIG. 17. Overabundance (ϑ) relative to 14 times solar abundances vs atomic mass number for nucleosynthesis resulting from a Type-II, Population I supernova with total mass equal to $25M_\odot$, from Woosley and Weaver (1982).

the collapsing and hardening core.

Integration over the mass zones of Fig. 16 for $1.5M_\odot < M < 2.2M_\odot$ and over those of Fig. 8(a) for $M > 2.2M_\odot$ enabled Woosley and Weaver (1982) to calculate the isotopic abundances ejected into the interstellar medium by their $25M_\odot$ Population I simulated supernova. The results relative to solar abundances (the reader should refer to the last paragraph of Sec. I) are shown in Fig. 17, taken from Woosley and Weaver (1982). The relative ratios are normalized to unity for ^{16}O, for which the overproduction ratio was 14, that is, for each gram of ^{16}O originally in the star, 14 grams were ejected. This overproduction in a single supernova can be expected to have produced the heavy-element abundances in the interstellar medium just prior to formation of the solar system, given the fact that supernovae occur approximately every one hundred years in the Galaxy. The ultimate theoretical calculations will yield a constant overproduction factor of the order of 10.

The results shown in Fig. 17 are disappointing if one expects the ejecta of $25M_\odot$ Population I supernovae to match solar system abundances with a relatively constant overproduction factor. The dip in abundances from sulfur to chromium is readily apparent. Woosley and Weaver (1982) point out that calculations must be made for other stellar masses and properly integrated over the mass distribution for stellar formation, which varies roughly inversely proportional to mass. Woosley, Axelrod, and Weaver (1984) discuss their expectations of the abundances produced in stellar explosions for stars in the mass range $10M_\odot$ to $10^6 M_\odot$. They show that a $200M_\odot$ Population III star produces abundant quantities of sulfur, argon, and calcium, which possibly compensate for the dip in Fig. 17. Population III stars are massive stars in the range $100M_\odot < M < 300M_\odot$, which are thought to have formed from hydrogen and helium early in the history of the Galaxy and evolved very rapidly. Since their heavy-element abundance was zero, they have no counterparts in presently forming Population I stars, as well as no counterparts among old, low-mass Population II stars.

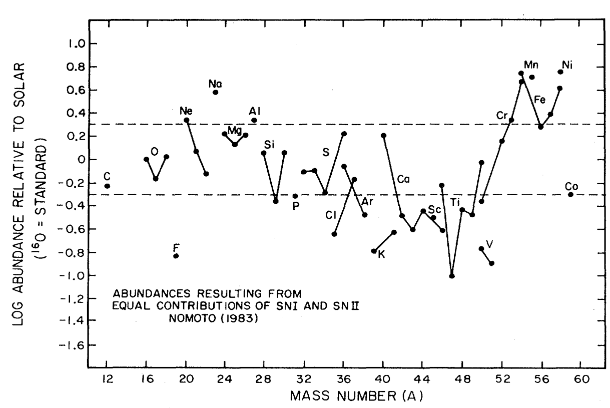

FIG. 18. Abundances relative to solar, with the abundance of ^{16}O taken as standard, produced by equal contributions from typical Type-I and Type-II supernovae, from Nomoto, Thielemann, and Wheeler (1984).

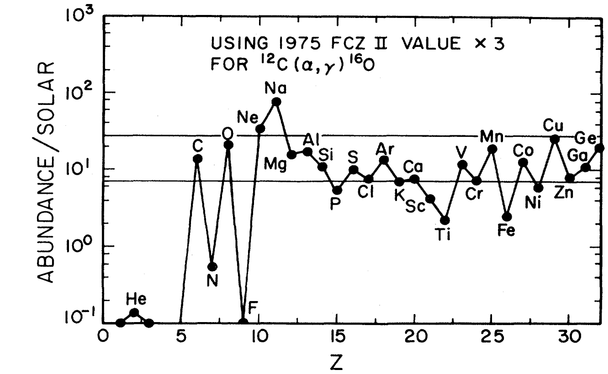

FIG. 19. Overabundance yields relative to solar vs atomic number Z, resulting from the explosion of a Type-II supernova with mass approximately equal to $20M_\odot$, from Arnett and Thielemann (1984). The horizontal lines are a factor of 2 higher and lower than the average overabundances, equal to 14. It is assumed that the presupernova abundances were not modified during the supernova explosion. The reaction rate for $^{12}C(\alpha,\gamma)^{16}O$ of Fowler, Caughlan, and Zimmerman (1967,1975) was multiplied by a factor of 3 in accordance with the theoretical analysis by Langanke and Koonin (1983).

Other authors have suggested a number of solutions to the problem depicted in Fig. 17. Nomoto, Thielemann, and Wheeler (1984) have calculated the abundances produced in carbon deflagration models of Type-I supernovae. By adding equal contributions from Type-I and Type-II supernovae they obtain Fig. 18, which can be considered somewhat more satisfactory than Fig. 17. On the other hand, Arnett and Thielemann (1984) have recalculated quasistatic presupernova nucleosynthesis for $M \approx 20M_\odot$ using a value for the $^{12}C(\alpha,\gamma)^{16}O$ rate equal to three times that given in Fowler, Caughlan, and Zimmerman (1967,1975), Harris *et al.* (1983), and Caughlan (1984). This would seem to be justified by the recent analysis of $^{12}C(\alpha,\gamma)^{16}O$ data by Langanke and Koonin (1983), as discussed in Sec. V. They then assume that explosive nucleosynthesis will not substantially modify their quasistatic abundances and obtain the results shown in Fig. 19. The average overproduction ratio is roughly 14, and deviations are in general within a factor of 2 of this value. However, their assumption of minor modification during explosion and ejection is questionable.

I feel that the results discussed in this section and those obtained by numerous other authors show promise of an eventual satisfactory answer to the question where and how did the elements from carbon to nickel originate. We shall see!

IX. ISOTOPIC ANOMALIES IN METEORITES AND OBSERVATIONAL EVIDENCE FOR ONGOING NUCLEOSYNTHESIS

Almost a decade ago it became clear that nucleosynthesis occurred in the Galaxy up to the time of formation of the solar system, or at least up until several million years before the formation. For slightly over one year it has been clear that nucleosynthesis continues up to the present time, or at least within several million years of the present. The decay of radioactive ^{26}Al ($\bar{\tau}=1.04\times10^6$ years) is the key to these statements, which bring great satisfaction to most experimentalists, theorists, and observers in Nuclear Astrophysics. For the record it must be admitted that the word "clear" is subject to certain reservations in the minds of some investigators, but as a believer, "clear" is clear to me.

Isotopic anomalies in meteorites produced by the decay of short-lived radioactive nuclei were first demonstrated in 1960 by Reynolds (1960), who found large enrichments of ^{129}Xe in the Richardson meteorite. Jeffery and Reynolds (1961) demonstrated in 1961 that the excess ^{129}Xe correlated with ^{127}I in the meteorite, and thus showed that the ^{129}Xe resulted from the decay *in situ* of ^{129}I ($\bar{\tau}=23\times10^6$ years). Quantitative results indicated that $^{129}\mathrm{I}/^{127}\mathrm{I}\approx10^{-4}$ at the time of meteorite formation. On the assumption that ^{129}I and ^{127}I are produced in roughly equal abundances in nucleosynthesis (most probably in the *r* process) over a period of $\sim10^{10}$ years in the Galaxy prior to formation of the solar system, and taking into account that only the ^{129}I produced over a period of the order of its lifetime survives, Wasserburg, Fowler, and Hoyle (1960) suggested that a period of free decay of the order of 10^8 years or more occurred between the last nucleosynthetic event which produced ^{129}I and its incorporation in meteorites in the solar system. There remains evidence for such a period in some cases, notably ^{244}Pu, but probably not in the history of the nucleosynthetic events which produced ^{129}I and other "short"-lived radioactive nuclei such as ^{26}Al and ^{107}Pd ($\bar{\tau}=9.4\times10^6$ years).

The substantiated meteoritic anomalies in ^{26}Mg from ^{26}Al, in ^{107}Ag from ^{107}Pd, in ^{129}Xe from ^{129}I, and in the heavy isotopes of Xe from the fission of ^{244}Pu ($\bar{\tau}=177\times10^6$ years; fission tracks also observed), as well as searches in the future for anomalies in ^{41}K from ^{41}Ca ($\bar{\tau}=0.14\times10^6$ years), in ^{60}Ni from ^{60}Fe ($\bar{\tau}=0.43\times10^6$ years), in ^{53}Cr from ^{53}Mn ($\bar{\tau}=5.3\times10^6$ years), and in ^{142}Nd from ^{146}Sm ($\bar{\tau}=149\times10^6$ years; α decay) are discussed exhaustively by my colleagues Wasserburg and Papanastassiou (1982). They espouse *in situ* decay for the observations to date, but my former student D. D. Clayton (1975,1979,1983,1984; see also Clayton and Hoyle, 1974,1976) argues that the anomalies occur in interstellar grains preserved in the meteorites and originally produced by condensation in the expanding and cooling envelopes of supernovae and novae. Wasserburg and Papanastassiou (1982) write on p. 90, "There is, as yet, no compelling evidence for the presence of preserved presolar grains in the solar system. All of the samples so far investigated appear to have melted or condensed from a gas, and to have chemically reacted to form new phases." With mixed emotions I accept this.

Before turning to some elaboration of the $^{26}\mathrm{Al}/^{26}\mathrm{Mg}$ case it is appropriate to return to a discussion of the free decay interval mentioned above. It is the lack of detectable anomalies in ^{235}U from the decay of ^{247}Cm ($\bar{\tau}=23\times10^6$ years) in meteorites, as shown by Chen and Wasserburg (1981), coupled with the demonstrated occurrence of heavy Xe anomalies from the fission of ^{244}Pu ($\bar{\tau}=117\times10^6$ years), as discussed for example by Burnett, Stapanian, and Jones (1982), which demands a free decay interval of the order of several times 10^8 years. This interval is measured from the "last" *r*-process nucleosynthesis event (supernova?) which produced the actinides, Th, U, Pu, Cm, and beyond, up to the "last" nucleosynthesis events (novae? supernovae with short-run *r* processes?) which produced the short-lived nuclei ^{26}Al, ^{107}Pd, and ^{129}I before the formation of the solar system. The fact that the anomalies produced by these short-lived nuclei relative to normal abundances all are of the order of 10^{-4}, despite a wide range in their mean lifetimes (1.04 to 23×10^6 years), indicates that this anomaly range must be the result of inhomogeneous mixing of exotic materials with much larger quantities of normal solar system materials over a short time, rather than the result of free decay. The challenges presented by this conclusion are manifold. Figure 14 of Wasserburg and Papanastassiou shows the time scale for the formation of dust, rain, and hailstones in the early solar system and for the aggregation into chunks and eventually the terrestrial planets. The solar nebula was almost but not completely mixed when it collapsed to form the solar system. From ^{26}Al it becomes clear that the mixing time down to an inhomogeneity of only one part in 10^3 (see what follows) was on the order of 10^6 years.

Evidence that ^{26}Al was alive in interstellar material in the solar nebula which condensed and aggregated to form the parent body (planet in the asteroid belt?) of the Allende meteorite is shown in Fig. 20, taken with some modification from Lee, Papanastassiou, and Wasserburg

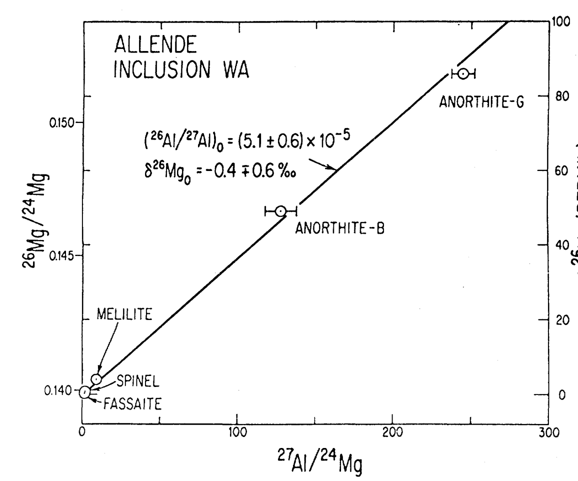

FIG. 20. Evidence for the *in situ* decay of ^{26}Al in various minerals in inclusion WA of the Allende meteorite, from Lee, Papanastassiou, and Wasserburg (1977). The linear relation between $^{26}\mathrm{Mg}/^{24}\mathrm{Mg}$ and $^{27}\mathrm{Al}/^{24}\mathrm{Mg}$ implies that $^{26}\mathrm{Al}/^{27}\mathrm{Al}=(5.1\pm0.6)\times10^{-5}$ at the time of formation of the inclusion, with ^{26}Al considered to react chemically in the same manner as ^{27}Al.

(1977). The Allende meteorite fell near Pueblito de Allende in Mexico on February 8, 1969 and is a carbonaceous chondrite, a type of meteorite thought to contain the most primitive material in the solar system, *unaltered since its original solidification.*

Figure 20 depicts the results for $^{26}Mg/^{24}Mg$ vs $^{27}Al/^{24}Mg$ in different mineral phases (spinel, etc.) from a Ca-Al-rich inclusion called WA obtained from a chondrule found in Allende. It will be clear that excess ^{26}Mg correlates linearly with the amount of ^{27}Al in the mineral phases. Since ^{26}Al is chemically identical with ^{27}Al, it can be inferred that phases rich in ^{27}Al were initially rich in ^{26}Al, which subsequently decayed *in situ* to produce excess ^{26}Mg. ^{26}Al was alive with abundance 5×10^{-5} that of ^{27}Al in one part of the solar nebula when the WA inclusion aggregated during the earliest stages of the formation of the solar system. The unaltered inclusion survived for 4.5 billion years to tell its story. Other inclusions in Allende and other meteorites yield $^{25}Al/^{27}Al$ from zero up to $\sim10^{-3}$, with 10^{-4} a representative value. The reader is referred to the paper of Wasserburg and Papanastassiou (1982) for the rich details of the story and the importance and significance of non-accelerator-based contributions to Nuclear Astrophysics.

Evidence that ^{26}Al is alive in the interstellar medium today is shown in Fig. 21, from Mahoney, Ling, Wheaton, and Jacobson (1984), my colleagues at Caltech's Jet Propulsion Laboratory (JPL). Figure 21 shows the gamma-ray spectrum observed in the range 1760—1824 keV by instruments aboard the High-Energy Astronomical Observatory, HEAO 3, which searched for diffuse gamma-ray emission from the Galactic equatorial plane.

The discrete line at 1809 keV, detected with a significance of nearly five standard deviations, is without doubt due to the transition from the first excited state at 1809 keV in ^{26}Mg to its ground state. Radioactive ^{26}Al decays by $^{26}Al(e^+\nu)^{26}Mg^*(\gamma)^{26}Mg$ to this state and thence to the ground state of ^{26}Mg. This gamma-ray transition shows clearly that ^{26}Al is alive in the interstellar medium in the Galactic equatorial plane today. Given the mean lifetime (1.04×10^6 years) of ^{26}Al, this shows that ^{26}Al has been produced no longer than several million years ago and is probably being produced continuously. It is no great extrapolation to argue that nucleosynthesis in general continues in the Galaxy at the present time. Quantitatively the observations indicate that $^{26}Al/^{27}Al\sim10^{-5}$ in the interstellar medium. This value averages over the Galactic plane interior to the sun at the present time. This average value was probably much the same when the solar system formed, but the variations in $^{26}Al/^{27}Al$ for various meteoritic inclusions show that there were wide variations in the solar nebula about this value ranging from zero to 10^{-3}.

The question immediately arises, what is the site of the synthesis of the ^{26}Al? Since the preparation of the paper by Ward and Fowler (1980) I have been convinced that ^{26}Al could not be by synthesized in supernovae at high temperatures where neutrons are copiously produced because of the expectation of a large cross section for $^{26}Al(n,p)^{26}Mg$. This expectation has been borne out by the measurements on the reverse reaction $^{26}Mg(p,n)^{26}Al$ in the Kellogg Laboratory by Skelton, Kavanagh, and Sargood (1983). Figure 22 is taken from Fig. 1(a) of these authors and shows the great beauty of high-resolution measurements in experimental Nuclear Astrophysics. Until the ^{26}Al targets just recently available can be bom-

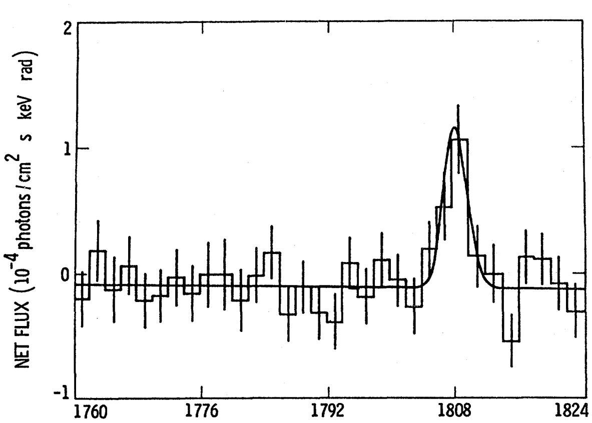

FIG. 21. The High-Energy Astrophysical Observatory (HEAO 3) data on gamma rays in the energy range 1760—1824 keV emitted from the Galactic equatorial plane, from Mahoney *et al.* (1984). The line at 1809 keV is attributed to the decay of radioactive ^{26}Al ($\bar{\tau}=1.04\times10^6$ years) to the excited state of ^{26}Mg at this energy.

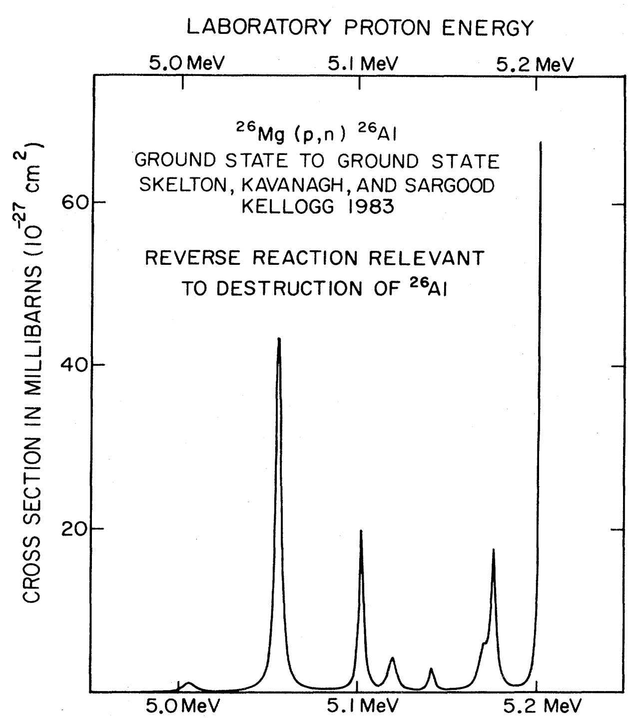

FIG. 22. The cross section in mb vs laboratory proton energy for the ground-state—to—ground-state reaction $^{26}Mg(p,n)^{26}Al$, from Skelton, Kavanagh, and Sargood (1983).

barded with neutrons it is necessary to supplement the laboratory measurements on ^{26}Mg(p,n)^{26}Al, perforce involving the ground state of ^{26}Mg, with theoretical calculations involving excited states (Woosley *et al.*, 1978), in order to calculate the stellar rate for ^{26}Al(n,p)^{26}Mg. There is little doubt that this rate is very large indeed.

Some writers (Clayton, 1975,1979,1983,1984; Clayton and Hoyle, 1974,1976; Mahoney *et al.*, 1982,1984; Arnould *et al.*, 1980) have suggested that ^{26}Al is produced in novae. This is quite reasonable on the basis of nucleosynthesis in novae as discussed in Truran (1982). In current models for novae, hydrogen from a binary companion is accreted by a white dwarf until a thermal runaway involving the fast CN cycle occurs. Similarly a fast MgAl cycle may occur with production of $^{26}\mathrm{Al}/^{27}\mathrm{Al} \gtrsim 1$, as shown in Fig. 9 of Ward and Fowler (1980). The recent experimental measurements cited in Ward and Fowler (1980) substantiate this conclusion. Clayton (1975,1979,1983,1984) argues that the estimated 40 novae occurring annually in the Galactic disk can produce the observed $^{26}\mathrm{Al}/^{27}\mathrm{Al}$ ratio of the order of 10^{-5} on average. He assumes that each nova ejects $10^{-4}M_{\odot}$ of material containing a mass fraction of ^{26}Al equal to 3×10^{-4}.

Another possible source of ^{26}Al is spallation induced by irradiation of protoplanetary material by high-energy protons from the young sun as it settled on the main sequence. This possibility was discussed very early by Fowler, Greenstein, and Hoyle (1962), who also attempted to produce D, Li, Be, and B in this way, requiring such large primary proton and secondary neutron fluxes that many features of the abundance curve in the solar system would have been changed substantially. A more reasonable version of the scenario was presented by Lee (1978) but without notable success. I find it difficult to believe that an early irradiation produced the anomalies in meteorites. The ^{26}Al in the interstellar medium today certainly cannot have been produced in this way.

Anomalies have been found in meteorites in the abundances compared with normal solar system material of the stable isotopes of many elements: O, Ne, Mg, Ca, Ti, Kr, Sr, Xe, Ba, Nd, and Sm. The possibility that the oxygen anomalies are non-nuclear in origin has been raised by Thiemens and Heidenreich (1983), but the anomalies in the remaining elements are generally attributed to nuclear processes.

One example is a neutron-capture/beta-decay $(n\beta)$ process studied by Sandler, Koonin, and Fowler (1982). The seed nuclei consisted of all of the elements from Si to Cr with normal solar system abundances. When this process operates at neutron densities $\approx 10^{7}$ mol cm^{-3} and exposure times of $\approx 10^{3}$ sec, small admixtures ($\lesssim 10^{-4}$) of the exotic material produced are sufficient to account for most of the Ca and Ti isotopic anomalies found in the Allende meteorite inclusion EK-1-4-1 by Niederer, Papanastassiou, and Wasserburg (1979). The anomalies in stable isotope abundances are of the same order as those for short-lived radioactive nuclei and strongly support the view that the solar nebula was inhomogeneous and not completely mixed, with regions containing exotic materials up to 10^{-4} or more of normal material.

Agreement for the ^{46}Ca and ^{49}Ti anomalies in EK-1-4-1 was obtained only by increasing the theoretical Hauser-Feshbach cross sections for ^{46}K(n,γ) and ^{49}Ca(n,γ) by a factor of 10 on the basis of probable thermal resonances just above threshold in the compound nuclei ^{47}K and ^{50}Ca, respectively. In a CERN report which subsequently became available Huck *et al.* (1981) reported an excited state in ^{50}Ca just 0.16 MeV above the ^{49}Ca(n,γ) threshold, which can be produced by *s*-wave capture and fulfills the requirements of Sandler, Koonin, and Fowler (1982).

Sandler *et al.* (1982) suggest that the $\approx 10^{3}$ sec exposure time scale is determined by the mean lifetime of ^{13}N ($\bar{\tau}=862$ sec), produced through ^{12}C(p,γ)^{13}N by a jet of hydrogen suddenly introduced into the helium-burning shell of a red giant star, where a substantial amount of ^{12}C has been produced by the $3\alpha \rightarrow {}^{12}$C process. The beta decay ^{13}N$(e^{+}\nu)$^{13}C is followed by ^{13}C(α,n)^{16}O as the source of the neutrons. All of this is very interesting, if true. More to the point, Sandler *et al.* (1982) predict the anomalies to be expected in the isotopes of chromium. Attempts to measure these anomalies are underway now by Wasserburg and his colleagues. Again, we shall see!

X. OBSERVATIONAL EVIDENCE FOR NUCLEOSYNTHESIS IN SUPERNOVAE

Over the years there has been considerable controversy concerning elemental abundance observations in the optical wavelength range on Galactic supernovae remnants. To my mind the most convincing evidence for nucleosynthesis in supernovae has been provided by Chevalier and Kirshner (1979), who obtained quantitative spectral information for several of the fast-moving knots in the supernova remnant Cassiopeia A (approximately dated 1659, but a supernova event was not observed at that time). The knots are considered to be material ejected from various layers of the original star in a highly asymmetric, nonspherical explosion. In one knot, labeled KB33, the following ratios *relative to solar, designated by brackets,* were observed: $[\mathrm{S/O}]=61$, $[\mathrm{Ar/O}]=55$, $[\mathrm{Ca/O}]=59$. It is abundantly clear that oxygen burning to the silicon-group elements in the layer in which KB33 originated has depleted oxygen and enhanced the silicon-group elements. Other knots and other features designated as filaments show different abundance patterns, albeit, not so easily interpreted. The moral for supernova modelers is that spherically symmetric supernova explosions may be the easiest to calculate, but are not to be taken as realistic. Admittedly they have a good answer: it is expensive enough to compute spherically symmetric models. OK, OK!

Most striking of all has been the payoff from the NASA investment in the High-Energy Astronomy Observatory (HEAO 2), now called the Einstein Observatory. Canizares and Winkler (1981) have analyzed the x-ray spectrum from 0.5 to 1.1 keV observed from Puppis A using the focal plane crystal spectrometer aboard the Einstein Observatory. Their abundance determinations are consistent with the ejection of $3M_\odot$ of oxygen and $1M_\odot$ of neon by a Type-II supernova of $M \geq 25M_\odot$ with subsequent mixing into $150M_\odot$ of interstellar material.

In addition Becker *et al.* (1979) observed the x-ray spectrum in the range 1–4 keV of Tycho Brahe's supernova remnant (1572) shown in Fig. 23. An x-ray spectrum is much simpler than an optical spectrum. For me it is wonderful that satellite observations show the K-level x rays expected from Si, S, Ar, and Ca just where the Handbook of Chemistry and Physics says they ought to be! Such observations are not all that easy in a terrestrial laboratory. Shull (1982,1983) has used a single-velocity, non-ionization-equilibrium model of a supernova blast wave to calculate abundances in Tycho's remnant *relative to solar, designated by brackets,* and finds: [Si]=7.6, [S]=6.5, [Ar]=3.2, and [Ca]=2.6. With considerably greater uncertainty he gives [Mg]=2.0 and [Fe]=2.1. He finds different enhancements in Kepler's remnant (1604) and in Cassiopeia A. One more lesson for the modelers: no two supernovae are alike. Nucleosynthesis in supernovae depends on their initial mass, rotation, mass loss during the red giant stage, the degree of symmetry during explosion, initial heavy-element content, and probably other factors. These details aside it seems clear that supernovae produce enhancements in elemental abundances up to iron and probably beyond. Detection of the much rarer elements beyond iron will require more sensitive x-ray detectors operating at higher energies. The nuclear debris of supernovae eventually enriches the interstellar medium from which succeeding generations of stars are formed. It becomes increasingly clear that novae also enrich the interstellar medium. Sorting out these two contributions poses interesting problems in ongoing research in all aspects of Nuclear Astrophysics.

Explosive silicon burning in the shell just outside a collapsing supernova core primarily produces ^{56}Ni, as shown in Fig. 16. It is generally believed that the initial energy source for the light curves of Type-I supernovae is electron capture by ^{56}Ni ($\bar{\tau}$=8.80 days) to the excited state of ^{56}Co at 1.720 MeV, with subsequent gamma-ray cascades to the ground state. These gamma rays are absorbed and provide energy to the ejected envelope. The subsequent source of energy is the electron capture and positron emission by ^{56}Co ($\bar{\tau}$=114 days) to a number of excited states of ^{56}Fe, with gamma-ray cascades to the stable ground state of ^{56}Fe. Both the positrons and the gamma rays heat the ejected material. If ^{56}Co is an energy source, there should be spectral evidence for cobalt in newly discovered Type-I supernovae, since its lifetime is long enough for detailed observations to be possible after the initial discovery.

The cobalt has been observed! Axelrod (1980) studied the optical spectra of SN1972e obtained by Kirshner and Oke (1975). The spectra obtained at 233, 264, and 376 days after Julian day 2441420, assigned as the initial day of the explosive event, are shown in Fig. 24. Axelrod as-

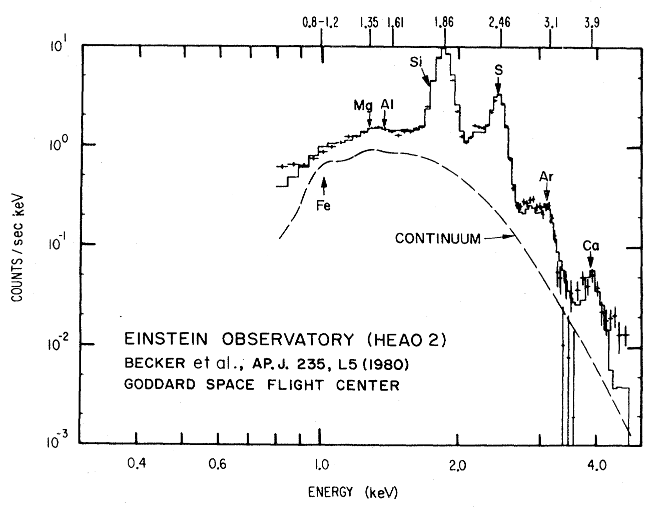

FIG. 23. The Einstein Observatory (HEAO 2) data on the x-ray spectrum of Tycho Brahe's supernova remnant, from Becker *et al.* (1979).

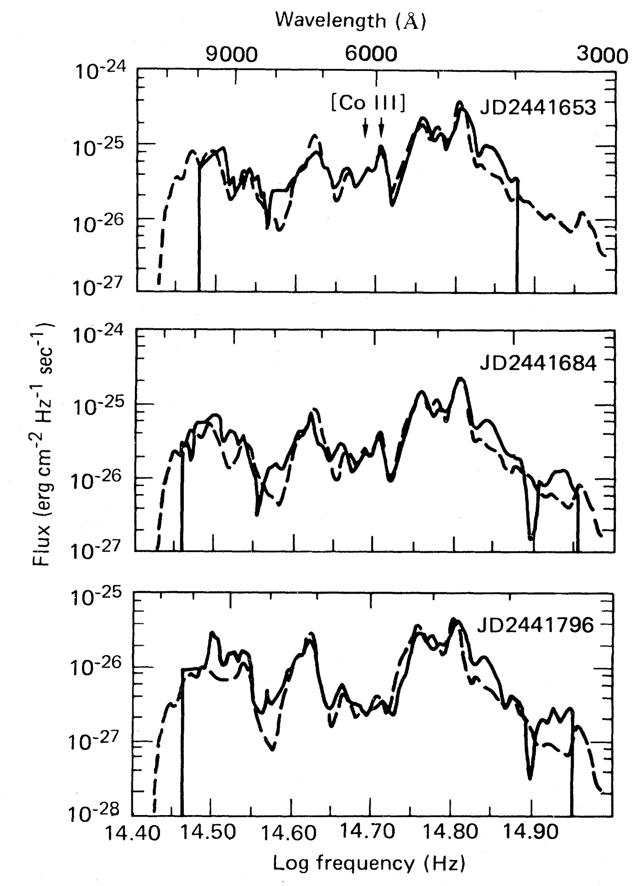

FIG. 24. Analysis by Axelrod (1980) yielding two emission lines from Co III in the observations on SN1972e obtained by Kirshner and Oke (1975). The observations were made 233, 264, and 376 days after (Julian day) JD2441420, assigned as the initial day of the supernova explosion. The mean lifetime of ^{56}Co is 114 days; the Co III lines appear to decay in keeping with their emission from radioactive ^{56}Co.

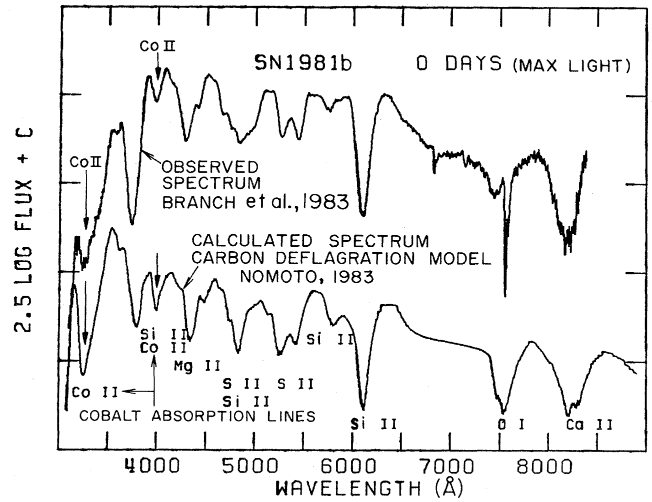

FIG. 25. Top: Analysis by Branch *et al.* (1983) of their absorption spectrum of SN1981*b* at maximum light, showing evidence for Co II absorption features. Bottom: Comparison with the calculated spectrum expected from the carbon deflagration model for Type-I supernovae, according to the calculations of Nomoto (1982a,1982b,1984).

signed the two emission lines near 6000 Å ($\log v = 14.7$) to Co III. They are clearly evident at 233 and 264 days, but are only marginally evident at 376 days ($\sim\bar{\tau}$ later). The lines decay in reasonable agreement with the mean lifetime of ^{56}Co.

Branch *et al.* (1983) have studied absorption spectra during the first hundred days of SN1981*b*. Their results at maximum light are shown in the top curve of Fig. 25. Using the carbon deflagration model for Type-I supernovae of Nomoto (1982a,1982b,1984), Branch (1984) has calculated the spectrum shown in the lower curve of Fig. 25. Deep absorption lines of Co II are clearly evident near 3300 Å and 4000 Å.

It is my conclusion that there is substantial evidence for nucleosynthesis in supernovae of elements produced in oxygen and silicon burning. The role of neutron capture processes in supernovae will be discussed in the next section.

XI. NEUTRON CAPTURE PROCESSES IN NUCLEOSYNTHESIS

In Sec. I the need for two neutron capture processes for nucleosynthesis beyond $A \gtrsim 60$ was discussed in terms of early historical developments in Nuclear Astrophysics. These two processes were designated the *s* process, for neutron capture slow (*s*) compared to electron beta decay, and the *r* process, for neutron capture rapid (*r*) compared to electron beta decay in the process networks.

For a given element the heavier isotopes are frequently bypassed in the *s* process and are produced only in the *r* process; thus the designation *r* only. Lighter isotopes are frequently shielded by more neutron-rich stable isobars in the *r* process and are produced only in the *s* process; thus the designation *s* only. The lightest isotopes are frequently very rare because they are not produced in either the *s* process or the *r* process and are thought to be produced in what is called the *p* process. The *p* process involves positron production and capture, proton capture, neutron photoproduction, and/or (p,n) reactions and will not be discussed further. The reader is referred to Audouze and Vauclair (1980). The results of the *s* process, the *r* process, and the *p* process are frequently illustrated by reference to the ten stable isotopes of tin. The reader is referred to Figs. 10 and 11 of Fowler (1964).

It is fair to say that the *s* process has the clearest phenomenological basis of all processes of nucleosynthesis. This is primarily the result of the correlation of *s*-process abundances first delineated by Seeger, Fowler, and Clayton (1965) with the beautiful series of measurements on neutron capture cross sections in the 1–100 keV range by the Oak Ridge National Laboratory group under Macklin and Gibbons (1965; also see Allen *et al.*, 1971).

This correlation is illustrated in Fig. 26, which shows the product of neutron capture cross sections (σ) at 30 keV multiplied by *s*-process abundances (N) as a function of atomic mass for *s*-only nuclei and those produced predominately by the *s* process. It is not difficult to understand in first-order approximation that the product σN should be constant in the *s*-process synthesis. A nucleus with a small (large) neutron capture cross section must have a large (small) abundance to maintain continuity in the *s*-capture path. Figure 26 demonstrates this in the *plateaus* shown from $A = 90$–140 and from $A = 140$–206. The anomalous behavior below $A = 80$ is discussed in Almeida and Käppeler (1983), from which Fig. 26 is taken.

Nuclear shell structure introduces the *precipices* shown in Fig. 26 at $A \sim 84$, $A \sim 138$, and $A \sim 208$, which correspond to the *s*-process abundance peaks in Fig. 2. At these values for A the neutron numbers are "magic," $N = 50$, 82, and 126. The cross sections for neutron cap-

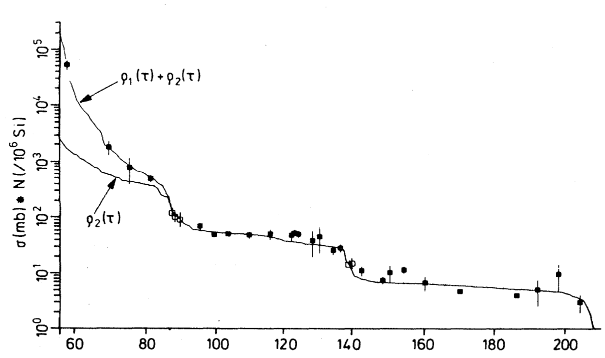

FIG. 26. Neutron capture cross section at 30 keV in mb multiplied by solar system abundances relative to Si $= 10^6$ vs atomic mass for nuclei produced in the *s* process, from Almeida and Käppeler (1983). Theoretical calculations are shown for a single exponential distribution, in neutron exposure, τ, $\rho_2(\tau)$, and for two such distributions, $\rho_1(\tau)+\rho_2(\tau)$.

ture into new neutron shells are very small at these magic numbers. With a finite supply of neutrons it follows that the σN product must drop to a new plateau just as observed. Quantitative explanations of this effect have been given by Ulrich (1973) and by Clayton and Ward (1974).

What is the site of the *s* process and what is the source of the neutrons? A very convincing answer has been given by Iben (1975), that the site is the He-burning shell of a pulsating red giant, with the $^{22}\mathrm{Ne}(\alpha,n)^{25}\mathrm{Mg}$ reaction as the neutron source. Critical discussions have been given by Almeida and Käppeler (1983) and by Truran (1983). The latter reference reserves the possibility that the $^{13}\mathrm{C}(\alpha,n)^{16}\mathrm{O}$ reaction is the neutron source.

We turn now to the *r* process. This process has been customarily treated by the *waiting point* method of B^{2}FH (1957). Under explosive conditions a large flux of neutrons drives nuclear seeds to the neutron-rich side of the valley of stability where, depending on the temperature, the (n,γ) reaction and the (γ,n) reaction reach equality. The nuclei wait at this point until electron beta decay transforms neutrons in the nuclei into protons, whence further neutron capture can occur. At the cessation of the *r* process the neutron-rich nuclei decay to their stable isobars. In first order this means that the abundance of an *r*-process nucleus multiplied by the electron beta-decay rate of its neutron-rich *r*-process isobar progenitor will be roughly constant. At magic neutron numbers in the neutron-rich progenitors, beta decay must perforce open the closed neutron shell in transforming a neutron into a proton, and there the rate will be relatively small. Accordingly the abundance of progenitors with $N=50$, 82, and 126 will be large. The associated number of protons will be less than in the corresponding *s*-process nuclei with a magic number of neutrons. It follows then that the stable daughter isobars will have smaller mass numbers, and this is indeed the case, the *r*-process abundance peaks occurring at $A\sim 80$, $A\sim 130$, and $A\sim 195$, all below the corresponding *s*-process peaks, as illustrated in Fig. 2.

A phenomenological correlation of *r*-process abundances with beta-decay rates made by Becker and Fowler (1959) and a detailed illustration of this correlation between solar system *r*-process abundances and theory is given in Fig. 13 of Fowler (1964). It is too phenomenological to satisfy critical nuclear astrophysicists. They wish to know the site of the high neutron fluxes demanded for *r*-process nucleosynthesis and the details of the *r*-process path through nuclei far from the line of beta stability.

There is also general belief at the present time that the waiting point approximation is a poor one and must be replaced by dynamical *r*-process flow calculations taking into account explicit (n,γ), (γ,n), and beta-decay rates with time-varying temperature and neutron flux. Schramm (1982) has discussed such calculations in some detail and has emphasized that nonequilibrium effects are particularly important during the freezeout at the end of the *r* process, when the temperature drops and the neutron flux falls to zero. Simple dynamical calculations have been made by Blake and Schramm (1976) for a process they designated as the *n* process and Sandler, Koonin, and Fowler (1982) for their $n\beta$ process discussed in Sec. IX. The most ambitious calculations have been made by Cameron, Cowan, and Truran (1984). This paper gives references to their previous herculean efforts in dynamical *r*-process calculations. An example of their results is shown in Fig. 27. They emphasize that they have not been able to find a plausible astrophysical scenario for the initial ambient conditions required for Fig. 27. In spite of this I am convinced that they are on the right track to an eventual understanding of the dynamics and site of the *r* process.

Many suggestions have been made for possible sites of the *r* process, almost all in supernova explosions where the basic requirement of a large neutron flux of short duration is met. These suggestions are reviewed in Schramm (1982) and Truran (1983). To my mind the helium core thermal runaway *r* process of Cameron, Cowan, and Truran (1984) is the most promising. These authors do not rule out the $^{22}\mathrm{Ne}(\alpha,n)^{25}\mathrm{Mg}$ reactions as the source of the neutrons, but their detailed results shown in Fig. 27 are based on the $^{13}\mathrm{C}(\alpha,n)^{16}\mathrm{O}$ reaction as the source. They start with a star formed from material with the same heavy-element abundance distribution as in the solar system, but with smaller total amount. They assume a significant amount of ^{13}C in the helium core of the star after hydrogen burning. This ^{13}C was produced previously by the introduction of hydrogen into the core, which had already burned half of its helium into ^{12}C. For Fig. 27 they assume a ^{13}C abundance of 14.3% by mass, density equal to $10^6\ \mathrm{g\,cm^{-3}}$, and an initial temperature of 1.6×10^8 K, which is raised by the initial slow ^{13}C burning to an eventual maximum of 3.6×10^8 K. The electrons in the core are initially degenerate, but the rise in temperature lifts the degeneracy, producing a thermal runaway with expansion and subsequent cooling of the core. This event is the second helium-flash episode in the history of the core and, if it occurs, only a small amount

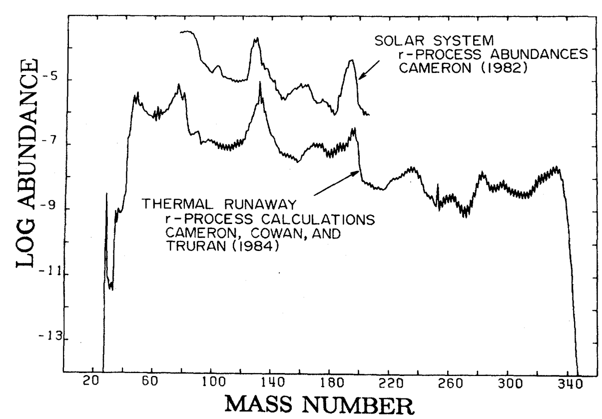

FIG. 27. Abundances produced in the *r* process vs atomic mass number in the thermal runaway model (lower curve) of Cameron, Cowan, and Truran (1984) compared with the solar system *r*-process abundances (upper curve) of Cameron (1982).

of the r-process material produced need escape into the interstellar medium to contribute the r-process abundance in solar system material. It is my belief that a realistic astrophysical site for the thermal runaway, perhaps with different initial conditions will be found. I rest the case.

XII. NUCLEOCOSMOCHRONOLOGY

Armed with their r-process calculations of the abundances of the long-lived parent of the natural radioactive series ^{232}Th, ^{235}U, and ^{238}U and with the then current solar system abundances of these nuclei, B^{2}FH (1957) were able to determine the duration of r-process nucleosynthesis from its beginning in the first stars in the Galaxy up to the last events before the formation of the solar system. The general idea was originally suggested by Rutherford (1929). B^{2}FH (1957) made a major advance in taking into account the contributions to the abundance of the long-lived *eon glasses* from the decay of their short-lived progenitors also produced in the r process. The parents of the natural radioactive series are indeed excellent *eon glasses* with their mean lifetimes: ^{232}Th, 20.0×10^9 years; ^{238}U, 6.51×10^9 years; ^{235}U, 1.03×10^9 years. The analogy with *hour glasses* is fairly good; the sand in the top of the *hour glass* is the radioactive parent, that in the bottom is the daughter. The analogy fails in that in the *eon glasses* "sand" (Th,U) is being added or removed, top and bottom by nucleosynthesis (production in stars) and astration (destruction in stars). Properly expressed differential equations can compensate for this failure.

The abundances used were those observed in meteorites assumed to be closed systems since their formation, taken to have occurred 4.55 billion years ago. It was necessary to correct for free decay during this period in order to obtain abundances for comparison with calculations based on r-process production plus decay over the duration of Galactic nucleosynthesis before the meteorites became closed systems. Fortunately ratios of abundances, ^{232}Th/^{238}U and ^{235}U/^{238}U, sufficed since absolute abundances could not, and still cannot, be calculated with the necessary precision. The calculations required only the elemental ratio Th/U in meteorites, since the isotopic ratio ^{235}U/^{238}U was assumed to be the same for meteoritic and terrestrial samples. The Apollo Program has added lunar data to the meteoritic and terrestrial in recent years.

B^{2}FH (1957) considered a number of possible models, one of which assumed r-process nucleosynthesis *uniform* in time and an arbitrary time interval between the last r-process contribution to the material of the solar nebula and the closure of the meteorite systems. A zero value for this time interval indicated that the production of uranium started 18 billion years ago. When this time interval was taken to be 0.5 billion years, the production started 11.5 billion years ago. These values are in remarkable, if coincidental, concordance with current values.

It is appropriate to point out at this point that nucleocosmochronology yields, with additional assumptions, an estimate for the age of the expanding universe completely independent of redshift-distance observations in astronomy on distant galaxies. The assumptions referred to in the previous sentence are that the r process started soon, less than a billion years, after the formation of the Galaxy and that the Galaxy formed soon, less than a billion years, after the "big bang" origin of the universe. Adding a billion years or so to the start of r-process nucleosynthesis yields an independent value, based on radioactivity, for the age or time back to the origin of the expanding universe.

Much has transpired over the recent years in the field of nucleocosmochronology. I have kept my hand in most recently in Fowler (1977; also see Fowler, 1972b). Exponentially decreasing nucleosynthesis, with the time constant in the negative exponent a free parameter to be determined by the observed abundance ratios along with the duration of nucleosynthesis, was introduced by Fowler and Hoyle (1960). For the time constant in the denominator of the exponent set equal to infinity, uniform synthesis results. When it is set equal to zero, a single spike of synthesis results. With two observed ratios, two free parameters in a model can be determined. As time went on the ratios ^{129}I/^{127}I and ^{244}Pu/^{238}U, with $\bar{\tau}(^{129}\mathrm{I})=0.023\times10^9$ years and $\bar{\tau}(^{244}\mathrm{Pu})=0.117\times10^9$ years, were added to nucleocosmochronology to permit the determination of two additional free parameters, the arbitrary time interval of B^{2}FH (1957) previously discussed and the fraction of r-process nucleosynthesis produced in a last gasp "spike" at the end of the exponential time dependence.

Sophisticated models of Galactic evolution were introduced by Tinsley (1975). A method for model-independent determinations of the mean age of nuclear chronometers at the time of solar system formation was developed by Schramm and Wasserburg (1970). In this method the mean age is one-half the duration for uniform synthesis and is equal to the actual time of single-spike nucleosynthesis. This indicates that one can expect no more than a range of a factor of 2 in the time back to the beginning of nucleosynthesis in widely different models for its time variation. These developments are reviewed in Schramm (1982).

The most recent calculations are those of Thielemann, Metzinger, and Klapdor (1983). Their results, revised by his own recent calculations, are shown in Fig. 28, prepared by F.-K. Thielemann. The presolar spike and its time of occurrence before the meteorites became closed systems depend primarily on the *minute glasses*, ^{129}I and ^{244}Pu. The *eon glasses*, ^{232}Th/^{238}U and ^{235}U/^{238}U, indicate that r-process nucleosynthesis in the Galaxy started 17.9 billion years ago with uncertainties of $+2$ billion years and -4 billion years according to Thielemann *et al.* (1983). This value is to be compared with my value of 10.5 ± 2.3 billion years ago given in Fowler (1977). Inputs of production and final abundance ratios have changed (Thielemann *et al.*, 1983)! Thielemann and I are now recomputing the new value for the duration using an initial spike in Galactic synthesis plus uniform synthesis thereafter.

Thielemann *et al.* indicate that the age of the expanding universe is 19 billion years, give or take several billion

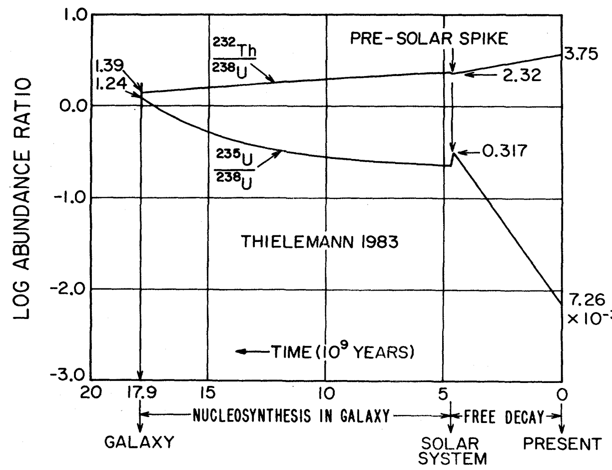

FIG. 28. The abundance ratios for $^{232}Th/^{238}U$ and for $^{235}U/^{238}U$ produced by theoretical r-process nucleosynthesis over the lifetime of the Galaxy prior to the formation of the solar system, from Thielemann, Metzinger, and Klapdor (1983). The free decays over the lifetime of the solar system to reach the present values for these ratios, $(^{232}Th/^{238}U)_0=3.75$ and $(^{235}U/^{238}U)_0=7.26\times10^{-3}$, are also shown. The production ratios in each r-process event was theoretically calculated to be 1.39 for $^{232}Th/^{238}U$ and 1.24 for $^{235}U/^{238}U$. Compare with Fig. 10 in Fowler (1977).

years. This to be compared to the Hubble time or reciprocal of Hubble's constant given by Sandage and Tammann (1982) as 19.5 ± 3 billion years. However, the Hubble time is equal to the age of the expanding universe only for a completely open universe with mean matter density much less than the critical density for closure, which can be calculated from the value for the Hubble time just given to be 5×10^{-30} g cm^{-3}. The observed visible matter in galaxies is estimated to be ten percent of this, which reduces the age of the universe to 16.5 billion years. Invisible matter, neutrinos, black holes, etc., may add to the gravitational forces which decrease the velocity of expansion and may thus decrease the age to that corresponding to critical density, which is 13.0 billion years. If the expansion velocity was greater in the past, the time to the present radius of the universe is correspondingly less. Moreover, there are those who obtain results for the Hubble time equal to about one-half that of Sandage and Tammann (1982), as reviewed in van den Bergh (1982). There is much to be done on all fronts!

A completely independent nuclear chronology involving the radiogenic ^{187}Os produced during Galactic nucleosynthesis by the decay of ^{187}Re ($\bar{\tau}=65\times10^9$ years) was suggested by Clayton (1964). Schramm (1982) discusses still other chronometric pairs. Clayton's suggestion involves the s process, even though the ^{187}Re is produced in the r process. It requires that the abundance of ^{187}Re, the parent, be compared to that of its daughter, ^{187}Os, when the s-only production of this daughter nucleus is subtracted from its total solar system abundance. This was to be done by comparing the neutron capture cross section of ^{187}Os with that of its neighboring s-only isotope ^{186}Os, which does not have a long-lived radioactive parent, and using the $N\sigma$=const rule for the s process.

Fowler (1972a) threw a monkey wrench into the works by pointing out that ^{187}Os has a low-lying excited state at 9.75 keV which is practically fully populated at $kt=30$ keV, corresponding to the temperature $T=3.5\times10^8$ K, at which the s process is customarily assumed to occur. Moreover, with spin $J=\frac{3}{2}$ this state has twice the statistical weight $(2J+1)$ of the ground state with spin $J=\frac{1}{2}$. Measurements of the ground-state neutron capture cross section yield only one-third of what one needs to know.

All of this has led to a series of beautiful and difficult measurements for neutron-induced reactions on the isotopes of osmium. Winters and Macklin (1982) found the Maxwell-Boltzmann average ground-state (laboratory) cross-section ratio for $^{186}Os(n,\gamma)$ relative to that for $^{187}Os(n,\gamma)$ to be 0.478 ± 0.022 at $kT=30$ keV, with a slow dependence on temperature. This ratio must be multiplied by a theoretical factor to correct the ^{187}Os cross section in the denominator of the cross-section ratio for that of its excited state. The larger the theoretical ^{187}Os excited-state capture, the smaller this factor. Woosley and Fowler (1979) used Hauser-Feshbach theory to give an estimate for this factor in the range 0.8—1.10, which is little comfort in view of the fact that it multiplies one number comparable to the number from which it must be subtracted. These factors translated into a time for the beginning of the r process in the Galaxy in the range 14—19 billion years. In desperation I suggested privately that inelastic neutron scattering off the ground state of ^{187}Os to its excited state at 9.75 keV might yield information on the properties of the excited state. Measurements by Macklin *et al.* (1983) and Hershberger *et al.* (1983) determined these inelastic neutron scattering cross sections, which yielded inherent support of the lower value of the Woosley and Fowler (1979) factor and thus a greater value for the time back to the beginning of r-process nucleosynthesis on the 18-to-20 billion-year range. It has to be admitted that this is concordant with the latest value from the Th/U nucleocosmochronology.

Once again in desperation I privately suggested that measurement of the neutron capture cross section on the ground state of ^{189}Os might be helpful. In ^{189}Os the ground state has the same spin and Nilsson numbers as the excited state of ^{187}Os and an excited state corresponding to the ground state of ^{187}Os. Measurements by Browne and Berman (1981) are available but are now being checked by an Oak Ridge National Laboratory, Denison University, and University of Kentucky consortium.

It will be clear that the lifetime of ^{187}Re comes directly into the calculations under discussion. There has been some discrepancy in the past between lifetimes measured geochemically and those measured directly by counting the electrons emitted in the 2.6-keV decay $^{187}Re(e^-\nu)^{187}Os$. Direct measurement yields only the lifetime for electron emission to the continuum, while geochemistry yields the lifetime for electron emission both to the continuum and to bound states in ^{187}Os. The entire matter is treated in considerable theoretical detail by Williams, Fowler, and Koonin (1984), who found that the bound-state decay is negligible and that the direct measurements by Payne (1965) and Drever (1983), which agree with the geochemical measurements of Hirt *et al.* (1963), are correct.

There is also the vexing problem of the possible decrease in the effective lifetime of ^{187}Re in the Galactic environment. The ^{187}Re included in the material of the interstellar medium which forms new stars is subject to destruction by the *s* process (astration) as well as being produced by the *r* process. This decreases the effective lifetime of the ^{187}Re and all chronometric time based on the Re/Os chronology. This problem is discussed in elaborate detail by Yokoi, Takahashi, and Arnould (1983). The time back to the beginning of *r*-process nucleosynthesis could be as low as 12 billion years. It is appropriate to end this last section before the concluding section with considerable uncertainty in nucleocosmochronology, indicating that, as in all Nuclear Astrophysics, there is much exciting experimental and theoretical work to be done for many years to come. Amen!

XIII. CONCLUSION

In spite of the past and current researches in experimental and theoretical Nuclear Astrophysics, illustrated in what I have just shown you, the ultimate goal of the field has not been attained. Hoyle's grand concept of element synthesis in the stars will not be truly established until we attain a deeper and more precise understanding of many nuclear processes operating in astrophysical environments. Hard work must continue on all aspects of the cycle: experiment, theory, observation. It is not just a matter of filling in the details. There are puzzles and problems in each part of the cycle which challenge the basic ideas underlying nucleosynthesis in stars. Not to worry—that is what makes the field active, exciting, and fun! It is a great source of satisfaction to me that the Kellogg Laboratory continues to play a leading role in experimental and theoretical Nuclear Astrophysics.

And now permit me to pass along one final thought in concluding my lecture. My major theme has been that all of the heavy elements, from carbon to uranium have been synthesized in stars. Let me remind you that your bodies consist for the most part of these heavy elements. Apart from hydrogen you are 65% oxygen and 18% carbon, with smaller percentages of nitrogen, sodium, magnesium, phosphorus, sulfur, chlorine, potassium, and traces of still heavier elements. Thus it is possible to say that you and your neighbor and I, each one of us and all of us, are truly and literally a little bit of stardust.

Charles Christian Lauritsen taught me a Swedish toast. I conclude with this toast to my Swedish friends: *"Din skål, min skål, alla vackra flickor skål. Skål!"*

ACKNOWLEDGMENTS

My work in Nuclear Astrophysics has involved collaborative teamwork with many people, and I am especially grateful to Fay Ajzenberg-Selove, Jean Audouze, C. A. Barnes, E. M. Burbidge, G. R. Burbidge, G. R. Caughlan, R. F. Christy, D. D. Clayton, G. M. Fuller, J. L. Greenstein, Fred Hoyle, Jean Humblet, R. W. Kavanagh, S. E. Koonin, C. C. Lauritsen, Thomas Lauritsen, D. N. Schramm, T. A. Tombrello, R. V. Wagoner, G. J. Wasserburg, Ward Whaling, S. E. Woosley, and B. A. Zimmerman.

For aid and helpful cooperation in all aspects of my scientific work, especially in the preparation of publications, I am grateful to Evaline Gibbs, Jan Rasmussen, Kim Stapp, Marty Watson, and Elizabeth Wood. I acknowledge support for my research over the years by the Office of Naval Research (1946 to 1970) and by the National Science Foundation (1968 to present).

REFERENCES

Ajzenberg-Selove, F., R. E. Brown, E. R. Flynn, and J. W. Sunier, 1984, Phys. Rev. Lett. (in press).

Allen, B. J., R. L. Macklin, and J. H. Gibbons, 1971, Adv. Nucl. Phys. **4**, 205.

Almeida, J., and F. Käppeler, 1983, Astrophys. J. **265**, 417.

Alpher, R. A., and R. C. Herman, 1950, Rev. Mod. Phys. **22**, 153.

Alvarez, L. W., 1949, University of California Radiation Laboratory Report UCRL-328.

Arnett, W. D., and F.-K. Thielemann, 1984, in *Stellar Nucleosynthesis,* edited by C. Chiosi and A. Renzini (Reidel, Dordrecht).

Arnould, M., H. Nørgaard, F.-K. Thielemann, and W. Hillebrandt, 1980, Astrophys. J. **237**, 931.

Audouze, J., and H. Reeves, 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 355.

Audouze, J., and S. Vauclair, 1980, *An Introduction to Nuclear Astrophysics* (Reidel, Dordrecht), p. 92.

Axelrod, T. S., 1980, "Late Time Optical Spectra from the ^{56}Ni Model for Type I Supernovae," Ph.D. thesis (University of California, Berkeley), UCRL-52994.

Bahcall, J. N., W. F. Huebner, S. H. Lubow, P. D. Parker, and R. K. Ulrich, 1982, Rev. Mod. Phys. **54**, 767.

Bahcall, N. A., and W. A. Fowler, 1969, Astrophys. J. **157**, 645.

Barnes, C. A., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 193.

Becker, R. A., and W. A. Fowler, 1959, Phys. Rev. **115**, 1410.

Becker, R. H., S. S. Holt, B. W. Smith, N. E. White, E. A. Boldt, R. F. Mushotsky, and P. J. Serlemitsos, 1979, Astrophys. J. Lett. **234**, L73.

Becker, S. A., 1983, private communication.

Bethe, H. A., 1939, Phys. Rev. **55**, 434.

Bethe, H. A., 1967, in *Les Prix Nobel 1967* (Almqvist and Wiksells, Stockholm).

Bethe, H. A., G. E. Brown, J. Cooperstein, and J. R. Wilson, 1983, Nucl. Phys. A **403**, 625.

Bethe, H. A., and C. L. Critchfield, 1938, Phys. Rev. **54**, 248.

Bethe, H. A., A. Yahil, and G. E. Brown, 1982, Astrophys. J. Lett. **262**, L7.

Blake, J. B., and D. N. Schramm, 1976, Astrophys. J. **209**, 846.

Bloom, S. D., and G. M. Fuller, 1984, Lawrence Livermore National Laboratory, in preparation.
Bodansky, D., D. D. Clayton, and W. A. Fowler, 1968, Astrophys. J. Suppl. **16**, 299.
Boyd, R. N., 1981, in *Proceedings of the Workshop on Radioactive Ion Beams and Small Cross Section Measurements* (Ohio State University, Columbus).
Branch, D., 1984, in Proceedings of Yerkes Observatory Conference on "Challenges and New Developments in Nucleosynthesis," edited by W. D. Arnett (University of Chicago, Chicago).
Branch, D., C. H. Lacy, M. L. McCall, P. G. Sutherland, A. Uomoto, J. C. Wheeler, and B. J. Wills, 1983, Astrophys. J. **270**, 123.
Brown, G. E., H. A. Bethe, and G. Baym, 1982, Nucl. Phys. A **375**, 481.
Browne, J. C., and B. L. Berman, 1981, Phys. Rev. C **23**, 1434.
Buchmann, L., M. Hilgemaier, A. Krauss, A. Redder, C. Rolfs, and H. P. Trautvetter, 1984, Z. Phys. (in press).
Burbidge, E. M., G. R. Burbidge, W. A. Fowler, and F. Hoyle, 1957, Rev. Mod. Phys. **29**, 547.
Burnett, D. S., M. I. Stapanian, and J. H. Jones, 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 144.
Cameron, A. G. W., 1957, Publ. Astron. Soc. Pac. **69**, 201.
Cameron, A. G. W., 1958, Annu. Rev. Nucl. Sci. **8**, 249.
Cameron, A. G. W., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 23.
Cameron, A. G. W., J. J. Cowen, and J. W. Truran, 1984, in Proceedings of Yerkes Observatory Conference on "Challenges and New Developments in Nucleosynthesis," edited by W. D. Arnett (University of Chicago, Chicago).
Canizares, C. R., and P. F. Winkler, 1981, Astrophys. J. **246**, L33.
Caughlan, G. R., W. A. Fowler, M. J. Harris, and B. A. Zimmerman, 1984, At. Data Nucl. Data Tables (in press).
Champagne, A. E., A. J. Howard, and P. D. Parker, 1983, Astrophys. J. **269**, 686.
Chandrasekhar, S., 1939, *Stellar Structure* (University of Chicago, Chicago).
Chen, J. H., and G. J. Wasserburg, 1981, Earth Planet. Sci. Lett. **52**, 1.
Chevalier, R. A., and R. P. Kirshner, 1979, Astrophys. J. 233, 154.
Chevalier, D. D., 1964, Astrophys. J. **139**, 637.
Clayton, D. D., 1975, Astrophys. J. **199**, 765.
Clayton, D. D., 1979, Space Sci. Rev. **24**, 147.
Clayton, D. D., 1983, Astrophys. J. **268**, 381.
Clayton, D. D., 1984, Astrophys. J. (in press).
Clayton, D. D., and F. Hoyle, 1974, Astrophys. J. Lett. **187**, L101.
Clayton, D. D., and F. Hoyle, 1976, Astrophys. J. **203**, 490.
Clayton, D. D., and R. A. Ward, 1974, Astrophys. J. **193**, 397.
Cook, C. W., W. A. Fowler, C. C. Lauritsen, and T. Lauritsen, 1957, Phys. Rev. **107**, 508.
Davis, R., Jr., B. T. Cleveland, and J. K. Rowley, in *Science Underground,* AIP Conference Proceedings No. 96, edited by M. M. Nieto, W. C. Haxton, C. M. Hoffman, E. W. Kolb, V. D. Sandberg, and J. W. Toevs (American Institute of Physics, New York), p. 2.
Drever, R. W. P., 1983, private communication.
Dunbar, D. N. F., R. E. Pixley, W. A. Wenzel, and W. Whaling, 1953, Phys. Rev. **92**, 64.
Dyer, P., and C. A. Barnes, 1974, Nucl. Phys. A **233**, 495.
Fowler, W. A., 1958, Astrophys. J. **127**, 551.
Fowler, W. A., 1964, Proc. Natl. Acad. Sci. USA **52**, 524.
Fowler, W. A., 1967, *Nuclear Astrophysics* (American Philosophical Society, Philadelphia).
Fowler, W. A., 1972a, Bull. Am. Astron. Soc. **4**, 412.
Fowler, W. A., 1972b, in *Cosmology, Fusion and Other Matters,* edited by F. Reines (Colorado Associated University Press, Boulder), p. 67.
Fowler, W. A., 1977, in *Proceedings of the Welch Foundation Conferences on Chemical Research, XXI, Cosmochemistry,* edited by W. D. Milligan (Robert A. Welch Foundation, Houston), p. 61.
Fowler, W. A., G. R. Caughlan, and B. A. Zimmerman, 1967, Annu. Rev. Astron. Astrophys. **5**, 525.
Fowler, W. A., G. R. Caughlan, and B. A. Zimmerman, 1975, Annu. Rev. Astron. Astrophys. **13**, 69.
Fowler, W. A., J. L. Greenstein, and F. Hoyle, 1962, Geophys. J. **6**, 148.
Fowler, W. A., and F. Hoyle, 1960, Ann. Phys. **10**, 280.
Fowler, W. A., and F. Hoyle, 1964, Astrophys. J. Suppl. **9**, 201.
Fuller, G. M., 1982, Astrophys. J. **252**, 741.
Fuller, G. M., W. A. Fowler, and M. J. Newman, 1980, Astrophys. J. Suppl. **42**, 447.
Fuller, G. M., W. A. Fowler, and M. J. Newman, 1982a, Astrophys. J. **252**, 715.
Fuller, G. M., W. A. Fowler, and M. J. Newman, 1982b, Astrophys. J. Suppl. **48**, 279.
Fuller, G. M., W. A. Fowler, and M. J. Newman, 1984, California Institute of Technology, in preparation.
Goodman, C. D., C. A. Goulding, M. B. Greenfield, J. Rapaport, D. E. Bainum, C. C. Foster, W. G. Love, and F. Petrovich, 1980, Phys. Rev. Lett. **44**, 1755.
Greenstein, J. L., 1954, in *Modern Physics for the Engineer,* edited by L. N. Ridenour (McGraw-Hill, New York), Chap. 10.
Greenstein, J. L., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 45.
Haight, R. C., G. J. Mathews, R. M. White, L. A. Avilés, and S. E. Woodward, 1983, Nucl. Instrum. Methods **212**, 245.
Harris, M. J., W. A. Fowler, G. R. Caughlan, and B. A. Zimmerman, 1983, Annu. Rev. Astron. Astrophys. **21**, 165.
Hauser, W., and H. Feshbach, 1952, Phys. Rev. **78**, 366.
Hemmindinger, A., 1948, Phys. Rev. **73**, 806.
Hemmindinger, A., 1949, Phys. Rev. **74**, 1267.
Hershberger, R. L., R. L. Macklin, M. Balakrishnan, N. W. Hill, and M. T. McEllistrem, 1983, Phys. Rev. C **28**, 2249.
Hirt, B., G. R. Tilton, W. Herr, and W. Hoffmeister, 1963, in *Earth Sciences and Meteorites,* edited by J. Geiss and E. D. Goldberg (North-Holland, Amsterdam).
Holmes, J. A., S. E. Woosley, W. A. Fowler, and B. A. Zimmerman, 1976, At. Data Nucl. Data Tables **18**, 305.
Hoyle, F., 1946, Mon. Not. R. Astron. Soc. **106**, 343.
Hoyle, F., 1954, Astrophys. J. Suppl. **1**, 121. See also Chandrasekhar, 1939, for a discussion of earlier ideas.
Hoyle, F., and W. A. Fowler, 1960, Astrophys. J. **132**, 565.
Hoyle, F., W. A. Fowler, E. M. Burbidge, and G. R. Burbidge, 1956, Science **124**, 611.
Hoyle, F., and M. Schwarzschild, 1955, Astrophys. J. Suppl. **2**, 1.
Huck, A., G. Klotz, A. Knipper, C. Miéhé, C. Richard-Serre, and G. Walter, 1981, CERN 81-09, 378.

Hulke, G., C. Rolfs, and H. P. Trautvetter, 1980, Z. Phys. A **297**, 161.

Iben, I., 1975, Astrophys. J. **196**, 525.

Jeffery, P. M., and J. H. Reynolds, 1961, J. Geophys. Res. **66**, 3582.

Kettner, K. U., H. W. Becker, L. Buchmann, J. Gorres, H. Kräwinkel, C. Rolfs, P. Schmalbrock, H. P. Trautvetter, and A. Vlieks, 1982, Z. Phys. A **308**, 73.

Kirshner, R. P., and J. B. Oke, 1975, Astrophys. J. **200**, 574.

Koonin, S. E., T. A. Tombrello, and G. Fox, 1974, Nucl. Phys. A **220**, 221.

Langanke, K., and S. E. Koonin, 1983, Nucl. Phys. A **410**, 334, and private communication.

Lederer, C. M., and V. S. Shirley, 1978, Eds., *Table of Isotopes: Seventh Edition* (Wiley, New York).

Lee, Typhoon, 1978, Astrophys. J. **224**, 217.

Lee, Typhoon, D. A. Papanastassiou, and G. J. Wasserburg, 1977, Astrophys. J. Lett. **211**, L107.

Macklin, R. L., and J. H. Gibbons, 1965, Rev. Mod. Phys. **37**, 166.

Macklin, R. L., R. R. Winters, N. W. Hill, and J. A. Harvey, 1983, Astrophys. J. **274**, 408.

Mahoney, W. A., J. C. Ling, A. S. Jacobson, and R. E. Lingenfelter, 1982, Astrophys. J. **262**, 742.

Mahoney, W. A., J. C. Ling, W. A. Wheaton, and A. S. Jacobson, 1984, Astrophys. J. (in press).

Mann, F. M., 1976, Hanford Engineering and Development Laboratory internal report HEDL-TME-7680.

Merrill, P. W., 1952, Science **115**, 484.

Michaud, G., and W. A. Fowler, 1970, Phys. Rev. C **2**, 2041.

Mitchell, L. W., and D. G. Sargood, 1983, Aust. J. Phys. **36**, 1.

Niederer, F. R., D. A. Papanastassiou, and G. J. Wasserburg, 1979, Astrophys. J. Lett. **228**, L93.

Nomoto, K., 1982a, Astrophys. J. **253**, 798.

Nomoto, K., 1982b, Astrophys. J. **257**, 780.

Nomoto, K., 1984, in *Stellar Nucleosynthesis,* edited by C. Chiosi and A. Renzini (Reidel, Dordrecht).

Nomoto, K., F.-K. Thielemann, and J. C. Wheeler, 1984, Astrophys. J. (in press).

Osborne, J. L., C. A. Barnes, R. W. Kavanagh, R. M. Kremer, G. J. Mathews, and J. L. Zyskind, 1982, Phys. Rev. Lett. **48**, 1664.

Payne, J. A., 1965, "An Investigation of the Beta Decay of Rhenium to Osmium Using High Temperature Proportional Counters," Ph.D. thesis (University of Glasgow).

Pontecorvo, B., 1946, Chalk River Laboratory Report PD-205.

Reynolds, J. H., 1960, Phys. Rev. Lett. **4**, 8.

Robertson, R. G. H., P. Dyer, T. J. Bowles, Ronald E. Brown, Nelson Jarmie, C. J. Maggiore, and S. M. Austin, 1983, Phys. Rev. C **27**, 11.

Rutherford, E., 1929, Nature **123**, 313.

Salpeter, E. E., 1952a, Astrophys. J. **115**, 326.

Salpeter, E. E., 1952b, Phys. Rev. **88**, 547.

Salpeter, E. E., 1955, Phys. Rev. **97**, 1237.

Sandage, A. R., and M. Schwarzschild, 1952, Astrophys. J. **116**, 463.

Sandage, A., and G. A. Tammann, 1982, Astrophys. J. **256**, 339.

Sandler, D G., S. E. Koonin, and W. A. Fowler, 1982, Astrophys. J. **259**, 908.

Sargood, D. G., 1982, Phys. Rep. **93**, 61.

Sargood, D. G., 1983, Aust. J. Phys. **36**, 583.

Schramm, D. N., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 325.

Schramm, D. N., and G. J. Wasserburg, 1970, Astrophys. J. **163**, 75.

Seeger, P. A., W. A. Fowler, and D. D. Clayton, 1965, Astrophys. J. Suppl. **11**, 121.

Shull, J. M., 1982, Astrophys. J. **262**, 308.

Shull, J. M., 1983, private communication.

Skelton, R. T., and R. W. Kavanagh, 1984, Nucl. Phys. A (in press).

Skelton, R. T., R. W. Kavanagh, and D. G. Sargood, 1983, Astrophys. J. **271**, 404.

Spinka, H., and H. Winkler, 1972, Astrophys. J. **174**, 455.

Starrfield, S. G., A. N. Cox, S. W. Hodson, and W. D. Pesnell, 1983, Astrophys. J. **268**, L27.

Staub, H., and W. E. Stephens, 1939, Phys. Rev. **55**, 131.

Suess, H. E., and H. C. Urey, 1956, Rev. Mod. Phys. **28**, 53.

Thielemann, F.-K., J. Metzinger, and H. V. Klapdor, 1983, Z. Phys. A **309**, 301, and private communication.

Thiemens, M. H., and J. E. Heidenreich, 1983, Science **219**, 1073.

Tinsely, B. M., 1975, Astrophys. J. **198**, 145.

Tollestrup, A. V., W. A. Fowler, and C. C. Lauritsen, 1949, Phys. Rev. **76**, 428.

Truran, J. W., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 467.

Truran, J. W., 1983, in *The Neutron and its Applications, 1982,* Institute of Physics Conference Series No. 64, edited by P. Schofield (Institute of Physics, Bristol/London), p. 95.

Ulrich, R. K. 1973, in *Explosive Nucleosynthesis,* edited by D. N. Schramm and W. D. Arnett (University of Texas, Austin), p. 139.

van den Bergh, S., 1982, Nature **229**, 297.

Vogt, E. W., 1969, Adv. Nucl. Phys. **1**, 261.

Wagoner, R. V., W. A. Fowler, and F. Hoyle, 1967, Astrophys. J. **148**, 3.

Ward, R. A., and W. A. Fowler, 1980, Astrophys. J. **238**, 266.

Wasserburg, G. J., W. A. Fowler, and F. Hoyle, 1960, Phys. Rev. Lett. **4**, 112.

Wasserburg, G. J., and D. A. Papanastassiou, 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 77.

Weaver, T. A., S. E. Woosley, and G. M. Fuller, 1983, in *Proceedings of the Conference on Numerical Astrophysics,* edited by R. Bowers, J. Centrella, J. LeBlanc, and M. LeBlanc (Science Books International, Boston).

Whaling, W., 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 65.

Wheeler, J. C., 1981, Rep. Prog. Phys. **44**, 85.

Williams, J. H., W. G. Shepherd, and R. O. Haxby, 1937, Phys. Rev. **52**, 390.

Williams, R. D., W. A. Fowler, and S. E. Koonin, 1984, Astrophys. J. (in press).

Wilson, H. S., R. W. Kavanagh, and F. M. Mann, 1980, Phys. Rev. C **22**, 1696.

Winters, R. R., and R. L. Macklin, 1982, Phys. Rev. C **25**, 208.

Woosley, S. E., T. S. Axelrod, and T. A. Weaver, 1984, in *Stellar Nucleosynthesis,* edited by C. Chiosi and A. Renzini (Reidel, Dordrecht).

Woosley, S. E., and W. A. Fowler, 1979, Astrophys. J. **233**, 411.

Woosley, S. E., W. A. Fowler, J. A. Holmes, and B. A. Zimmerman, 1978, At. Data Nucl. Data Tables **22**, 371.

Woosley, S. E., and T. A. Weaver, 1982, in *Essays in Nuclear Astrophysics,* edited by C. A. Barnes, D. D. Clayton, and D. N. Schramm (Cambridge University, Cambridge), p. 381.

Wu, S.-C., 1978, "Fusion Cross Section Measurements for the Reactions $^{14}N + {}^{10}B$ and $^{16}O + {}^{16}O$," Ph.D. thesis (California Institute of Technology).

Yokoi, K., K. Takahashi, and M. Arnould, 1983, Astron. Astrophys. **117**, 65.

Zyskind, J. L., C. A. Barnes, J. M. Davidson, W. A. Fowler, R. E. Marrs, and M. H. Shapiro, 1980, Nucl. Phys. A **343**, 295.

Zyskind, J. L., J. M. Davidson, M. T. Esat, M. H. Shapiro, and R. H. Spear, 1978, Nucl. Phys. A **301**, 179.

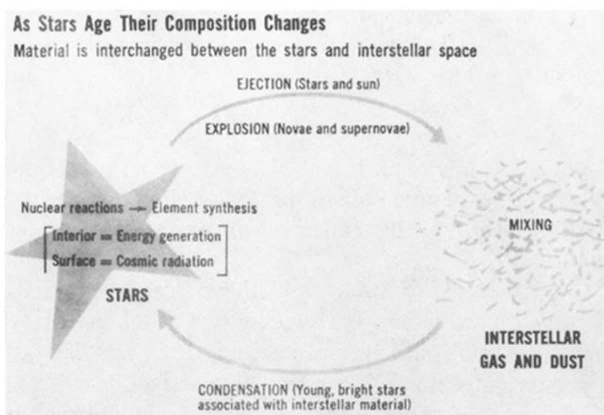

FIG. 1. Synthesis of the elements in stars.